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The origin and abundances of the chemical elements*†

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A NATO Advanced Studies Institute on the Origins and Abundances of the Chemical Elements was held at the Institute of Astronomy, Cambridge, England from 23 July to 9 August 1974. The problems discussed there are the basis for this review. Three major areas can be identified. First, the normal distribution of chemical and isotopic abundances and deviations from it must be determined observationally. Second, the relevant nuclear processes and their sites and products must be identified. Third, a model of the evolution of galaxies which includes the relevant processes and produces the observed abundances must be formulated. There has been rapid progress recently in all three of these areas.

CONTENTS

* The literature search for this review was terminated on 1 April 1975.
† Dedicated to my father, Lyne Starling Trimble, who taught me that, when you have a long story to tell, you must start at the beginning, keep going until you get to the end, and then stop, and to my late mother, Virginia Farmer Trimble, who listened with great love and patience to my early efforts at story-telling.
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I. INTRODUCTION AND HISTORY

The problems of determining the abundances of the chemical elements and their isotopes and accounting for them in terms of a theory of nucleosynthesis and the evolution of our Galaxy involves virtually every branch of astronomy, and many related sciences, from objective prism spectroscopy and neutron activation techniques to shell models of the nucleus and computerized histories of the Galaxy. It is not perhaps surprising then that no completely unique, clear picture of nucleosynthesis and evolution has yet emerged.

In order to be able to account for the abundances of the chemical elements, we must first know what they are. The first attempt at compiling a table of "cosmic abundances" (we will discuss shortly what this is supposed to mean) was made by Goldschmidt (1937), who relied primarily on data taken from meteorites and stellar spectra. He distinguished elements as siderophilic, chalcophilic, or lithophilic, depending on whether they were found primarily in the metal, sulphide, or silicate phases of meteorites, and adopted a very high abundance of Li, Be, and B on the basis of igneous rocks.

Other important compilations were those of Brown (1949) and Suess and Urey (1956), who were able to make considerable improvements in combining the data from various sources because of better understanding of the ratio of metal to silicate phases in meteorites and the Earth, the ratio of volatile to refractory elements in the Sun and stars (which allows you to link up stellar and meteorite data), and the ratios of isotopes which were by then also regarded as containing important information.

The current standard abundance distributions to which theoretical predictions are usually compared are those by Cameron (1968, 1973). The latter of these does not differ very much from the former and is not regarded by everyone as being an improvement. In fact the standard abundances have been remarkably stable. Only 8 elements out of the 81 tabulated had abundances which changed by factors of 3 or more between 1956 and 1968: C, Ga, and Pb were revised upward by factors of 3, 4, and 8.5, respectively, and Be, La, Sm, Ta, and W downward by factors of 25, 4, 3, 3, and 3, respectively.

The most surprising thing, in a way, is that such a compilation is possible at all. To what extent can we expect to represent with a single set of numbers objects as diverse as the Earth (more than half of whose atoms, at least in the crust, are oxygen), the Sun (mostly hydrogen), and the iron meteorites (whose name is self-explanatory)? If we aim at "cosmic" rather than Solar System abundances, the situation might appear even worse. There are stars like R CrB (whose atmosphere is evidently about half helium and half carbon) and α^2 CVn (whose spectrum shows 1000 times as much Eu as our best guess at the solar value), and large numbers of Population II stars whose entire heavy element abundance, (i.e. everything but hydrogen and helium) is a percent or less of the solar.

But, in fact, once allowance is made for some secondary processes, abundances are remarkably similar nearly everywhere we look. The meteorites, for instance, have obviously lost much of their volatile material (H, He, C, N, O, and the rare gases), but the relative proportions of the heavier elements in the Type I carbonaceous chondrites (those which show the least evidence of remelting after formation) are very similar to the solar ratios. This is even true to a lesser extent for terrestrial rocks, given that the siderophilic and chalcophilic elements are preferentially hidden in the core, and that volatile materials have again been lost. Moon rocks are similarly made up of only the more refractory elements, but again in standard proportions. Meteoritic, terrestrial, and lunar samples are particularly valuable because isotopic ratios (which are generally not much affected by chemical fractionation) can be accurately determined in them, which is not generally possible from solar or stellar spectroscopy. Thus, it is possible to arrive at a representative set of Solar System abundances, which (with a few exceptions) represents the surface of the Sun now, and which should represent the original solar nebula out of which the Sun and planets were formed, in the conventional view of things.

In addition, abundances very similar to those in the Solar System are found in the interstellar gas in H II regions around newly formed stars (called H II regions because the hydrogen there has been ionized by the ultraviolet radiation from the hot young stars) and in planetary nebulae (the ejected outer layers of older low mass stars). Among the stars as well, abundances are quite similar, provided that we confine our attention to stars which (a) are not evolved off the main sequence so that their surfaces have not been polluted by the products of their own nuclear reactions, (b) have convective atmospheres or are rapidly rotating, so that their surface layers are kept well mixed, and (c) belong to Population I or the disk population of stars which, on kinematic grounds, must have formed after the first billion years or so of the life of our galaxy. Even the very old Population II stars, although their total metal[1] abundance is very low, have those heavy elements that can be detected present in nearly solar relative proportions. The integrated spectra of other galaxies and quasars do not exhibit gross differences from "normal" abundances either.

Thus, it is possible to define a standard abundance distribution which applies to the Solar System and many other objects, and, then, to discuss how other objects deviate from this norm. This will be done in Sec. II, special attention being given to the problem of differentiating abundance anomalies that reflect nuclear processes from those that result from fractionation, diffusion, and so forth. Such a distribution will be referred to hereinafter as "cosmic" or "normal" abundances.

The next step is to account for the observed abundances in terms of nuclear processes and their products. Modern discussions of this problem build on two fundamental approaches. The first is synthesis in an early, dense phase of the universe, as formulated in the work of Gamow and his colleagues (Alpher *et al.*, 1948, often called Alpha, Beta, and Gamma; Alpher and Herman, 1950, 1953). This is now believed to be responsible for most of the synthesis of the very light nuclides (^{2}H, ^{3}He, ^{4}He, and very possibly ^{7}Li). The second is nucleosynthesis in stars, as formulated by Burbidge, Burbidge, Fowler, and Hoyle (1957,

[1] Astronomers have customarily called all elements but hydrogen and helium "metals." There is probably nothing that can be done about this.

usually known as B^2FH) and Cameron (1957). They considered seven processes which are (with some modifications) now believed to occur in stars, the earlier ones (hydrogen and helium burning) being associated with well-defined evolutionary stages and definite positions of the stars on a color-magnitude (Hertzsprung–Russell) diagram. The others appear to occur in later, more rapid stages of evolution, generally of massive stars. Finally, they postulated an x process, which was necessary to account for the observed amount of Li, Be, and B (which tend to be destroyed rather than produced in stars because they burn at low temperatures). This is now attributed to spallation of C, N, and O in the interstellar medium by passing cosmic ray protons and alpha particles (and conversely). These processes and their probable sites and products are discussed in Sec. III. The section concludes with a discussion of how to sum over the various zones of a star to determine its total yield of heavy elements, and how to sum over stars of various masses to get the yield per generation.

Thus, very broadly, we must think in terms of material, with an initial composition determined by the Big Bang, condensing into galaxies. Then the first generations of stars more massive than the Sun undergo nuclear reactions and return the products to the interstellar medium. Successive generations of stars, including the Sun and its planetary system, will then be formed from the material thus enriched in heavy elements. The pioneering work of Eggen, Lynden-Bell, and Sandage (1962) made clear that this process was accompanied, in the case of our Galaxy (and presumably other spirals), by the collapse of an initially spherical gas distribution to a flat disk of gas and stars. Models of the chemical and dynamical evolution of galaxies of various types will be discussed in Sec. IV.

Section V summarizes the present state of these various problems and concludes with exhortations to observers, experimenters, and theorists.

II. OBSERVED ABUNDANCES

A. The problem of the retrieval of meaningful data

The chief difficulty in determining the observed standard abundance distributions is deciding just which bits of the wealth of available raw data really belong to the problem and which are telling us about some process other than nucleosynthesis. Unfortunately, not everyone agrees on which bits of data belong in which pile. We discuss here briefly data which the majority of speakers at the Institute (NATO, 1974) believe do not belong to the problem.

1. D/H in molecules

For many years, the only available abundance for deuterium came from terrestrial (Craig, 1961) and meteoritic (Boato, 1954) water, which give D/H = $(1.5\text{–}1.6) \times 10^{-4}$, an uncomfortably high value for some theorists. This inspired a discussion by Geiss and Reeves (1972) of the probable preferential formation of HDO over H_2O by molecular exchange interactions. They estimated a prefractionation ratio of 2.5×10^{-5} for the solar nebula. This is more or less confirmed by a measurement of D/H = 2 $\times 10^{-5}$ in the hydrogen molecules in Jupiter (Trauger *et al.*, 1973). We should not, therefore, be surprised by the high D/H ratios (up to 6×10^{-3}) found for HCN in the Orion Nebula by Jefferts *et al.* (1973) and interpreted, again in terms of exchange reactions, by Solomon and Woolf (1973) as being consistent with the value of D/H $= 1.4 \times 10^{-5}$ in the atomic phase of the present interstellar medium (Rogerson and York, 1973).

2. The solar corona/photosphere iron discrepancy

For many years, the coronal iron abundance (as well as the meteoritic) seemed to be about 7.8 (Pottasch, 1964), on the logarithmic scale where H = 12.00, while the photospheric abundance was more like 6.5, prompting various theories of separation in the solar nebula as well as in the Sun itself. Improved values of the oscillator strengths (f values) for the permitted lines of Fe (e.g., Bridges and Wiese, 1970; Wolnick *et al.*, 1971) have increased the photospheric abundance to 7.3–7.6 and the discrepancy to a tolerable level. Additional recent determinations of transition probabilities for Fe are given by Smith *et al.* (1970), Whaling *et al.* (1970), and Martinez-Garcia *et al.* (1971), and the solar abundance derived from them by Smith and Whaling (1973). Ni has had a similar, but less spectacular history. The f values are given by Lennard *et al.* (1973) and the solar abundance derived by Lennard *et al.* (1975).

It might seem that most of the determinations of relative metal richness or poorness in stars, which are usually expressed and frequently measured as log(Fe/H) relative to the same ratio in the Sun, would now be seen to be nonsense. But luckily the same (bad) f values were used for both solar and stellar studies, so that relative abundances are nearly unaffected.

3. Weak He lines in QSOs and at the galactic center

The helium emission lines are sometimes much weaker relative to the hydrogen lines in quasistellar objects than in typical emission nebulae (e.g., Wampler, 1968). Bahcall and Oke (1971), among others, interpreted the observed line strengths as implying a ratio by number He/H as low as 0.003 (for 3C273) and 0.01 for some other objects. If this represented the original material of the QSO's, it would be a serious objection to synthesis of helium in the early universe. Williams (1971) points out, however, that a blob of gas which is (a) ionized by a synchrotron (power-law) spectrum of ultraviolet photons and (b) optically thick in the Lyman continuum of both H^0 and He^+ will show weak He lines even for a "normal" ratio He/H = 0.1. The assumptions are likely ones in light of the radio and (probably) optical synchrotron radiation of QSOs and the large amount of gas required to produce their emission lines. Shields (1974) has pointed out that the apparent variation in He/H in Seyfert galaxies (He/H = 0.09–0.23 from conventional analysis) is probably also an ionization effect.

Early reports (Mezger *et al.*, 1970) of a low He/H ratio at the galactic center, based on radio recombination lines, also caused some theoretical distress to believers in "cosmic" helium. But again, the distress seems to have been unnecessary. The catch in these low resolution observations is that the He^+ and H^+ lines really come from two Strömgren spheres of ionized helium and hydrogen, respectively. (Lines from He^{++} are never seen in normal H II regions.) There is direct observational evidence that this is a problem. The measured He/H ratio in the Orion Nebula is normal near the center and gradually declines to essentially zero about one parsec away from the center (Churchwell, in NATO, 1974), implying that the He line emitting region is smaller than the H line emitting region. With poorer resolution, this would look like low He/H. There are at least two circumstances under which the helium Strömgren sphere is expected to be smaller than the hydrogen one. These have been discussed by Mezger *et al.* (1974) and the observations reexamined in light of the discussion by Churchwell *et al.* (1974). First, a star with an equivalent blackbody temperature less than about 37 000°K (later than 09) will emit enough fewer helium-ionizing photons (228–504 Å) than hydrogen-ionizing photons (504–912 Å) to give a smaller He sphere. This may be relevant for some spiral arm H II regions and perhaps 30 Doradus in the Large Magellanic Cloud (C. Cesarsky, in NATO, 1974) but probably not for the galactic center. The problem is further discussed by Rodriguez *et al.* (1975), who include the effects of helium abundance in the ionizing star.

The second thing that can mislead us in this way is dust which preferentially absorbs the shorter-wavelength helium-ionizing photons. Mezger *et al.* (1974) point out that, where this is important, there should also be considerable infrared emission. Again, there is evidence for such a correlation. We can, for instance, compare M17, which shows essentially normal helium (He/H = 0.092) and little or no infrared emission by dust, with W3, which is a strong ir source and shows variable He/H across the nebula with the He disappearing completely in places (Churchwell, in NATO, 1974). This mechanism can account for the apparent absence of He at the galactic center if the ratio of dust to gas there is about 10 times what it is in spiral arm regions. Alternatively, or in addition, the dust could have a different dependence of absorption coefficient on wavelength there, so that it was about 10 times more effective at absorbing He-ionizing photons. There is observational evidence for both effects. The ir emission from the galactic center suggests that the dust to gas ratio there is 2–3 times higher than in spiral arms (Cesarsky, in NATO, 1974), while the anomalously high strength of the 9.7 μ silicate absorption feature in the direction of the galactic center compared to the visual extinction there suggests unusual wavelength dependence of the absorption cross section. [The visible and infrared extinction by small silicate particles has recently been remeasured by Day *et al.* (1974), who find the infrared efficiency at least 3 times higher than had previously been supposed.] Thus the low He/H ratio observed for some H II regions near the galactic center probably reflects dust distribution rather than low abundance. A genuinely low He abundance cannot be conclusively excluded by these considerations, and more work is clearly required. He/H appears normal in one galactic center H II region, Sgr A West (Pauls *et al.*, 1974).

4. Peculiar A stars (including helium-weak and helium-strong B stars, Mn–Hg stars, Ap stars of the Si λ 4200 and Eu–Cr–Sr types, Am stars, and their Population II analogs)

This heterogeneous and improbable group comprises main sequence (or almost so) Population I stars from about B0 to F5 and some Population II stars of similar temperature, which display a wide variety of surface abundance anomalies not attributable to unusual ionization or excitation conditions. They include a few percent of the early and mid B stars, about 10% of the late B and early A, and up to 30% of the main sequence near A5–6, declining to a low percentage of all stars again by F2. They have in common low rotation velocities (less than about 100 km/sec compared to about 200 km/sec for normal hot main sequence stars) often indistinguishable from zero rotation, and ratios of radiation pressure to surface gravity exceeding that of a 1.5 $\mathfrak{M}_\odot$ unevolved star. The incidence of spectroscopic binaries appears to be 100% among the Am's (Abt, 1965), normal among the Mn–Hg stars, and unusually low (20% or less) among the Si and Eu–Cr–Sr types (Abt and Snowden, 1974). The visual binary frequency is normal for the latter two groups. Most or all Ap's show evidence of magnetic fields in excess of 1000 G when the lines are sharp enough to measure, and they are frequently variable in luminosity, field strength, and spectrum with the same period for all variations. The other types have magnetic fields less than 100 G and no strong evidence for variability. Observations of Ap stars have been reviewed by Preston (1974) and by Jaschek and Jaschek (1974). Among the types of anomalies found are:

(a) Enhanced helium, accompanied by abnormally high O and N, but normal C, Si, Mg, and Al in a few very hot B stars, for instance σ Ori E, studied by Osmer and Peterson (1974).

(b) Helium apparently deficient by a factor of 5 or more among cooler B stars (Norris, 1971), accompanied by other anomalies in P, Ti, and CNO. An interesting example is 3 Cen A, in which $^3He/^4He$ is greater than unity (Sargent and Jugaku, 1961) and P is greatly enriched. A more typical example is α Sco in which weak He (^4He!) is accompanied by enhancements of Sr, Cr, and Ti (Jugaku and Sargent, 1961). No stars are found to have weak He and be otherwise normal.

(c) An overabundance of Mn and Hg (which defines the class as Mn or Mn–Hg stars) among late B stars with a general enhancement of the iron peak and heavier elements (except Cr), especially gallium and yttrium. Typical examples are 53 Tau (Aller and Bidelman, 1964) and K Cnc (Sargent, 1964). Helium is somewhat underabundant (Auer *et al.*, 1966). Mn stars are discussed by Wolff and Wolff (1974), and Cowley and Aikman (1975) have pointed out that some of the anomalies cannot be accounted for by any reasonable nuclear process, because odd Z's are more abundant than even Z's.

(d) An overabundance of Si in late B to middle A stars (defining the class as Si stars or λ 4200 stars), accompanied by weak He, C, O, and Ne and an enhanced iron peak except for Mn and Ni. These are probably a higher temperature variant of the next group.

(e) Large enhancements of the elements beyond Fe, especially the rare earths (hence the name europium stars), accompanied by normal iron peak abundances, among A stars. Adelman (1973) has studied a number of these classic Ap stars, of which the best known example is α^2 CVn, and finds that many of the anomalies correlate with effective temperature. Even more surprising is the great relative excess of the heavy isotopes of Hg (Preston, 1971) and Pt (Dworetsky and Vaughan, 1973), amounting in some cases to almost 100% cf ^{204}Hg or ^{198}Pt. The excess of neutron rich isotopes is greatest in the cooler stars.

(f) Deficiencies of Ca and Sc, sometimes accompanied by a general enhancement of Fe and heavier elements. Conti (1970) has discussed both the subtle Am stars like Sirius and the more extreme ones like τ U Ma and picked out weak Sc as the defining characteristic of the class.

(g) Apparent weakness of everything but hydrogen among A and early F stars that are, dynamically, Population I and rapid rotators. The prototype star, λ Boo, was discovered by Morgan *et al.* (1943), and the class is discussed by Sargent (1968).

(h) Population II analogies of most of the above types. For instance, Feige 86 (Sargent and Searle, 1968) is a He weak star; 38 Draconis a Mn star; and BD + 5°4268 a Si λ 4200 star.

There are several reasons to conclude that only the surface layers of these stars are involved in the abundance anomalies. First, they all have temperatures such that there will be no large surface convection zone, and nearly all are slow rotators or have strong magnetic fields (or both), which also tend to reduce mixing. Second, even a few supernova explosions of stars whose whole volume was as much enriched in rare earths as the surface of α^2 CVn would raise the average abundance of these things in the galaxy well above what is observed in the solar system. The concentration of certain of the enhanced abundances in spots on the stellar surface in the spectrum variable Ap stars, the close correlation of some anomalies with surface temperature, and the similarity of the effects in Population I and Population II stars also seem inconsistent with the whole mass of the star being involved, as in the model of Fowler *et al.* (1965), in which surface spallation and interior nuclear reactions are followed by mixing. Explanations involving nuclear reactions have, therefore, concentrated either on surface nuclear reactions, as discussed by Brancazio and Cameron (1967, who considered reactions of seed nuclei with alpha particles accelerated in a surface magnetic field of 10^{6-7} G) and Novikov and Sunyaev (1967, for ^{3}He) and more recently by Kuchowicz (1971) and Tjin a Djie *et al.* (1973), with special reference to promethium, or on contamination due to the supernova of a binary companion, the accretion being influenced by the magnetic field of the Ap star, as discussed in two series of papers by van den Heuvel (1967) and by Guthrie (1970, 1971a,b, 1972). Drobyshevskii (1974) has presented yet another model of this type.

It now seems probable that all of these theories should be rejected. The arguments discussed by Michaud (1970), by Pikel'ner and Khokhlova (1971), and by Vauclair and Reeves (1972) with special reference to ^{3}He production by spallation, include the large amounts of energy required for surface nuclear reactions, the absence of stars with normal spectra and strong magnetic fields, and the details of the anomalies and their correlations with the temperature and surface gravity of the stars rather than with evolutionary status.

A promising alternative mechanism involving diffusion in the stars' atmospheres was originally proposed by Greenstein *et al.* (1967) to account for weak He in old halo B stars. Michaud (1970) considered diffusion in the atmospheres of Ap stars and was able to account for the deficiencies of He and O and the overabundance of Si, Mn, and many other heavy elements in the temperature regions where they are observed, provided that the atmospheric turbulence is sufficiently low. The excess of ^{3}He in 3 Cen A and other stars can also be accounted for in this way (Vauclair *et al.*, 1974), since being lighter, it is less prone to sink than ^{4}He. A modification of the mechanism by Smith (1971) and Watson (1971) also considers the lifting of certain elements by radiation pressure and the sinking of others under the influence of gravity, but in a stable radiative zone above the HeII ionization region and below the hydrogen convection zone in A stars. The abundances seen in Am stars as a function of temperature are well reproduced, even at the low temperature end (Smith, 1973).

Further work (Michaud, 1973, 1974, and in NATO, 1974) seems to indicate that the anomalies can be built up on a reasonable time scale (e.g., 10^5 years) and that large excesses of either the heaviest or the lightest isotopes of some elements will result, depending on whether the dominant effect is lifting an element into the line formation zone or blowing it off the star completely. Atoms can be trapped in the line formation region either by magnetic fields or by recombining to an ionization state that has few lines in the part of the spectrum where the star is brightest. Additional work, including non-LTE effects and detailed models for particular stars, is in progress.

5. Low energy solar flare particles

It has recently become possible to determine the composition of particles from solar flares down to energies less than 1 MeV/amu. Typical trends include a ratio Fe/He gradually increasing from close to the photospheric value at 20–30 MeV/amu to 10 or 20 times it at a few MeV/amu and the appearance of fluctuations in He/CNO below 1 MeV/amu which do not occur at higher energies (Price, 1973). Since we know that these particles come from the Sun, there is no ambiguity in attributing abundance variations to the acceleration mechanism. This must, of course, also be true for more energetic solar flare particles, but since they provide our primary standard for the Solar System abundance of He and Ne and may also have something to tell us about how acceleration affects cosmic ray abundances, they are discussed in Sec. II.B.4. below.

6. Galactic cosmic rays (as observed)

The gross distribution of the chemical elements in the cosmic rays coming from outside the Solar System in those energy ranges where we have detailed information (100 MeV—a few GeV/amu) is very much like that of everything else: a great predominance of H and He (with He/H about 0.1 by number), about 1% as much C, N, and O, and a gradual decrease of abundance with atomic number, interrupted by the Fe peak, toward $Z \sim 45$, followed by a leveling off at $N(Z)/N(\mathrm{Fe}) = 10^{-6}$ (to within a factor of 10) beyond that. But there are also some significant differences. If we normalize at, e.g., carbon, then H and He are depleted by about a factor of 3–10, while the iron peak and heavier elements are enhanced by similar factors. The cause of these effects is not clear. It is discussed below in Sec. II.E on cosmic ray source composition.

In addition, the large "holes" in the cosmic abundances at Li, Be, and B, F, and Sc to Mn are partially filled in, and the odd–even effects partly smoothed over. This filling in of gaps is reasonably well understood on the basis of spallation, heavy cosmic ray nuclei being fragmented into lighter ones by collisions with protons and helium nuclei in the interstellar medium. Thus, ^{4}He gives rise to ^{3}He and ^{2}H; C, N, and O to Li, Be, and B; Ne, Si, and Mg to F; Fe to Sc, Ti, and V; nuclei in the Pb and Pt peaks to nuclei in the range $Z = 60$–75, and so forth. We should, therefore, be able, at least in principle, to take the observed abundance of each isotope and extract the abundances that must have left the source. In practice, this is not a trivial operation, since there are, for instance, about 1000 paths that lead from Fe to Li, most of them involving spallation cross sections that have never been measured. A general scheme for extracting the source abundances is discussed by Shapiro and Silberberg (1970, 1971), and a semiempirical approach to getting many of the necessary cross sections by Silberberg and Tsao (1973a,b). Shirk *et al.* (1973) have adopted the alternative approach of assuming an initial composition (e.g., Solar System or pure *r* process) and propagating it forward through interstellar space and comparing the results with the observed abundances. The derived source abundances are discussed in Sec. II.E.

The amount of interstellar matter that must be traversed to produce the observed amount of spallation varies with both the atomic number and the energy range considered. For instance, to produce the Li, Be, and B seen from C, N, and O requires 5–7 g/cm^2, assuming He/H = 0.1 in the interstellar medium (Shapiro and Silberberg, 1971), while Comstock *et al.* (1972) find that the observed ^{2}H and ^{3}He can be made from ^{4}He by 4 g/cm^2. The persistence of nuclei as heavy as U and Th requires that some material have come through much less than 1 g/cm^2, since they are very easily destroyed. All of the data for energies below a few GeV/amu can be fitted by an exponential distribution of path lengths, $N(x) = (1/x) \times \exp(-x/x_0)$, with x_0 in the range 4–6 g/cm^2 (Shirk *et al.*, 1973; Price, 1973). This corresponds to the cosmic rays being confined in the galactic disk (with an average density of 1 atom/cm^3) for 3 million years or in the galactic halo (with an average density of 0.01 atom/cm^3) for 300 million years. A decision between the two kinds of confinement must be made on the basis of an independent measurement of the age of the cosmic rays. A suitable clock exists in the form of ^{10}Be, which is produced at a well-known ratio to the other spallation products, and then decays with a half-life of 1.6×10^6 yr. Other secondary nuclei with half-lives between 10^5 and 2×10^6 years which might also be useful in this regard are ^{26}Al, ^{36}Cl, and ^{53}Mn. Unfortunately, individual isotope abundances cannot be measured for any of these at the present time, but the total abundance ratio, Be/(Li + B) is consistent with part, but not all, of the ^{10}Be having decayed, and a confinement time of order 10^7 yr (Meneguzzi *et al.*, 1971).

Above a few GeV/amu, the average ratio of secondary nuclides to primaries gradually decreases with increasing energy. For instance, ($Z = 17$–25)/(Fe + Ni) drops from about 1.2 at 1 GeV/amu to 0.2 or less above 10 GeV/amu. The ratios (Li + Be + B)/(C + N + O) and N/(C + O) also seem to decline with increasing energy. In addition, (C + O)/(Fe + Ni) decreases with increasing energy, although all of these are primaries. Recent data are given by Ormes and Balasubrahmanyan (1973), Webber *et al.* (1973), Juliusson *et al.* (1972), and O'Sullivan *et al.* (1973), and discussed by Audouze and Cesarsky (1973), Meneguzzi (1973), and Price (1973). A consistent explanation of all the data is obtained if the mean path length traversed by the cosmic rays decreases from 6–7 g/cm^2 at 1–2 GeV/amu to less than 1 g/cm^2 at 100 GeV/amu, with x_0 varying as $E^{-\frac{1}{2}}$ to E^{-1}. The effect of this on secondary to primary ratios is obvious. (C + O)/(Fe + Ni) is also affected because the iron peak nuclei are more readily broken up by spallation than carbon and oxygen, thus at low energies, where much of the Fe and Ni have been eroded away, the (C + O)/(Fe + Ni) ratio looks artificially high. That this is the correct explanation is indicated by the fact that the ratio (C + O)/(Fe + Ni + $Z = 17$–25) is about the same at all energies observed in the range 1–50 GeV/amu. This means that Fe and Ni are gradually transformed to lighter nuclei with increasing spallation, while the destruction of C + O is less important [as it is indicated by (Li + Be + B)/(C + O), which is less than 0.3 at all energies].

The ratios at the highest measured energies tend to resemble those in the calculated source abundances. It therefore seems hopeful that detailed composition measurements at sufficiently high energies may eventually be able to tell us the source abundances directly. The highest energies presently available preserve the main features of H and He depleted by about a factor of 10, and Fe and heavier elements enhanced by a factor of 3–10 if we normalize at carbon.

B. Solar system abundances

1. The Earth–Moon system

By far the most convenient sample of Solar System material available for analysis is the Earth. Unfortunately, 4.6 billion years of geologic processing have produced so much chemical fractionation that the only statements we can make naively are of the form that silicon seems to be a good deal commoner than europium. But on the same basis we might have said that oxygen seems to be a good deal commoner than helium, which will turn out to be entirely wrong for the Solar System as a whole, so that although a great deal is known about the composition of various components of the Earth's crust (see, for instance, the papers in Sec. V of Ahrens, 1968), the Earth is not a primary standard for any chemical abundances. Rather, effort has gone into trying to account for the composition of the Earth in terms of condensation from a solar nebula with a composition determined in other ways, and subsequent geological processing.

Terrestrial materials are, however, our primary standard for the ratios of isotopic abundances within individual chemical elements. All isotopic abundances in the compilation by Cameron (1973) discussed in Sec. II.B.2 are terrestrial, except those of H, He, Ne, Xe, and ^{40}Ar (the last being an interpolation). In addition, those isotopic abundances which involve the parent or daughter nuclide of radioactive decay processes have been corrected back to the birth of the Solar System. It is believed that isotopic ratios (except for the very light elements; see the discussion of D/H in Sec. II.A.1) are not greatly affected by chemical processes. A typical result is that of Reynolds (1953), who found that the ratio ^{30}Si/^{28}Si did not vary by more than 0.2% from his standard value for olivine (obtained from dunite from North Carolina) in any of a wide variety of minerals. The standard compilation of isotopic abundances, which was used by Cameron (1968, 1973), is that of Fuller and Nier (1965). The same data is also available in, e.g., any recent edition of *The Handbook of Chemistry and Physics* or *The American Institute of Physics Handbook*.

The Moon appears to be even more severely fractionated than the Earth, being relatively more depleted in volatile elements and enriched in those with low condensation temperatures, by factors of about 10 and 3, respectively (Ganapathy and Anders, 1974). Again the primary effort has gone into accounting for the observed surface composition in terms of condensation processes and a theory of subsequent fractionation. A recent review of the problems is given by Anderson (1973); it is not even absolutely clear whether the interior of the Moon is richer or poorer in iron and associated elements than the surface. Of course, the same thing could be said of the Earth, if Lyttleton's (1973 and references therein) phase change theory is allowed as a viable alternative to the conventional iron–nickel core. There are no significant (for nucleosynthesis) Earth–Moon differences in isotope ratios, apart from those involved in radioactive decays, cosmic ray production and the like.

2. Meteorites

Meteoritic data is the primary source for the standard solar system elemental abundances. It also contains information on possible interstellar components in the Solar System and on extinct radioactivities.

a. The standard elemental abundance distribution. The standard Solar System abundance distribution (Cameron, 1973) is obtained primarily from meteoritic data, especially the Type I carbonaceous chondrites (those which show the least evidence of remelting after condensation). The first column of Table I shows essentially Cameron's adopted elemental abundances, and the same data is shown graphically in Fig. 1. Slightly modified values for K, Ca, Fe, Zr, Hf, and Th, recommended by Anders (in NATO, 1974) are included. Some elements (especially those which are gases under normal circumstances) are greatly depleted in all types of meteorites, and their abundances must be obtained in other ways. We discuss these first, before returning to the general question of the reliability and meaning of the meteoritic data.

H, C, N, and O are observed spectroscopically in the Sun. Thus, if we believe that the solar surface is a good sample of its total composition [neglecting, for instance, Joss's (1974) suggestion that the surface is much enhanced in heavy elements due to the accretion of comet-like material, and Hoyle's classic remarks about the surface composition of chimney sweeps], the only problem remaining is one of normalization. This arises because solar abundances are determined fairly directly with respect to hydrogen (which supplies most of the continuous opacity) and are usually expressed on a scale in which $\log N(\mathrm{H}) = 12.00$, while meteoritic abundances are customarily determined and given with respect to silicon, with $N(\mathrm{Si}) = 10^6$. Cameron's (1973) solution to the problem was to take the average of the ratios of the meteoritic to solar (as given by Withbroe, 1971) abundances of the eight elements Mg, Al, Si, P, S, Ca, Fe, and Ni as his normalization constant. This results in N(meteoritic scale) $= 0.03175 \times N$(solar scale) for any element. Other authors have normalized at Si alone, which appears less satisfactory.

He and Ne do not appear in either meteorites or the solar photosphere (because of their very high first excitation potentials), but do occur in energetic particles from solar flares. The observations are discussed in Sec. II.B.4. Cameron has adopted the results of Bertsch *et al.* (1972), who give Ne/O $= 0.16 \pm 0.03$ and He/O $= 103 \pm 10$. In view of the variable composition of solar flare particles, these are clearly rather uncertain, and cannot be thought of as inconsistent with any other values (e.g., solar corona). He/H is well determined in other stars and the interstellar medium (see Secs. II.C and II.D) and appears to be constant at 0.1 where not affected by processes in the objects themselves. Ne is much more difficult, and "well determined" modern values range at least from 2.5×10^6 (for nearby H II regions, Peimbert and Costero, 1969) to 1.3×10^7 (for τ Sco, analyzed by Hardorp and Scholz, 1970), with the solar coronal value (10^6) on the low end of the range. It is hard to take the variations very seriously.

Four elements, Ar, Kr, Nb, and Xe, are not observed well anough anywhere in the solar system for their abundances to be measured. Cameron's adopted values for these are interpolations between nearby nuclei on the basis of the systematics of nuclear reactions. Their abundances cannot, therefore, be used as tests of theories of nucleosynthesis, although their isotope ratios can be. Kr, Nb, and Xe are not well enough observed anywhere else in the universe for conflicts to arise, but there is some evidence (discussed at the end of Sec. II.B.3) that the adopted value for argon may be too high by a factor of about 4. The adopted values for S, Cl, and K have also been influenced by considerations of nuclear systematics, so that the adopted values for Cl and K are about a factor

TABLE I. Relative abundances by number of the chemical elements in the meteorites, solar photosphere, solar corona, and galactic cosmic ray sources. The abundances of Th and U have been corrected to the time of formation of the Solar System. The solar abundances are normalized at hydrogen and the cosmic ray source at carbon. The meteoritic data are those of Cameron (1973) and include solar values for boron and the volatile elements; the photospheric data are those of Engvold and Hauge (1974); the coronal data are those of Withbroe (1971) modified by subsequent measurements discussed by Engvold and Hauge (1974) and by Walker *et al.* (1974a,b). The cosmic ray source data are those of Shapiro and Silberberg (1974) and Price (1973). Q is the quality of the photospheric determinations as estimated by Engvold and Hauge (1974). a represents errors of less than 30%, b less than a factor of 2.5, and c less than a factor of 10. Unstable elements are omitted.

Z, Element		Meteorites	Solar photosphere	Q	Solar corona	Cosmic ray sources
1	H	3.18×10^{10}	2.5×10^{10}	a	2.5×10^{10}	4.1×10^{9}
2	He	2.21×10^{9}	2×10^{9}	b	2.0×10^{9}	3.1×10^{8}
3	Li	49.5	0.2	b	...	...
4	Be	0.81	0.2	b	...	...
5	B	3.2	<4.0	b	...	...
6	C	1.18×10^{7}	10^{7}	a	1.4×10^{7}	1.18×10^{7}
7	N	3.74×10^{6}	3×10^{6}	a	2.8×10^{6}	1.3×10^{6}
8	O	2.15×10^{7}	1.6×10^{7}	a	2.0×10^{7}	1.3×10^{7}
9	F	2450	1000	c	...	...
10	Ne	3.44×10^{6}	10^{6}	b	1.7×10^{6}	1.8×10^{6}
11	Na	6.0×10^{4}	5×10^{4}	a	5.3×10^{4}	9.4×10^{4}
12	Mg	1.061×10^{6}	8×10^{5}	a	9.4×10^{5}	2.7×10^{6}
13	Al	8.5×10^{4}	8×10^{4}	a	7.9×10^{4}	2.4×10^{5}
14	Si	1×10^{6}	10^{6}	a	1.1×10^{6}	2.4×10^{6}
15	P	9600	10^{4}	b	7100	2.4×10^{4}
16	S	5.0×10^{5}	4×10^{5}	a	3.5×10^{5}	3.5×10^{5}
17	Cl	5700	8×10^{3}	b	...	...
18	Ar	1.172×10^{5}	2.4×10^{4}	c	8×10^{4}	8.3×10^{4}
19	K	3790	8×10^{3}	c	1.4×10^{4}	...
20	Ca	6.25×10^{4}	6×10^{4}	a	6.3×10^{4}	2.6×10^{5}
21	Sc	35	30	b	315	...
22	Ti	2775	1600	b	5000	...
23	V	262	250	b	1.6×10^{4}	...
24	Cr	1.27×10^{4}	1.6×10^{4}	b	1.8×10^{4}	3.5×10^{4}
25	Mn	9300	6000	b	8900	2.4×10^{4}
26	Fe	8.9×10^{5}	6×10^{5}	b	8.2×10^{5}	2.6×10^{6}
27	Co	2210	800	b	5600	...
28	Ni	4.80×10^{4}	8×10^{4}	a	8.5×10^{4}	9.4×10^{4}
29	Cu	540	400	b	1000	7.4×10^{3}
30	Zn	1244	630	b	...	9750 ($Z = 30 - 31$)
31	Ga	48	16	b	...	...
32	Ge	115	80	b	...	1.2×10^{3} ($Z = 32 - 34$)
33	As	6.6	No lines		...	...
34	Se	67.2	No lines		...	...
35	Br	13.5	No lines		...	113 ($Z = 35 - 39$)
36	Kr	46.8	No lines		...	...
37	Rb	5.88	10	b	...	...
38	Sr	26.9	20	a	25	...
39	Y	4.8	1.6	b	...	...
40	Zr	15.1	16	b	...	61 ($Z = 40 - 44$)
41	Nb	1.4	5	b	...	...
42	Mo	4.0	8	b	...	...
44	Ru	1.9	2.5	b	...	...
45	Rh	0.4	0.8	b	...	0.9 ($Z = 45 - 49$)
46	Pd	1.3	0.6	b	...	...
47	Ag	0.45	0.2	b	...	...
48	Cd	1.48	2.5	b	...	...
49	In	0.189	1.3	b	...	...
50	Sn	3.6	0.8	b	...	14.4 ($Z = 50 - 54$)
51	Sb	0.316	0.25	b	...	...
52	Te	6.42	No lines		...	...
53	I	1.09	No lines		...	...
54	Xe	5.38	No lines		...	...
55	Cs	0.387	<2.0	b	...	10.6 ($Z = 55 - 59$)
56	Ba	4.8	2.5	b	10	...
57	La	0.445	1.6	b	...	...
58	Ce	1.18	2.0	b	...	...
59	Pr	0.149	1.0	b	...	...
60	Nd	0.78	1.6	b	...	0.5 ($Z = 60 - 64$)
62	Sm	0.226	1.3	b	...	...
63	Eu	0.085	0.13	b	...	...
64	Gd	0.297	0.3	b	...	...
65	Tb	0.055	No f values		...	0.2 ($Z = 65 - 69$)
66	Dy	0.36	0.3	b	...	...
67	Ho	0.079	No f values		...	...
68	Er	0.225	0.16	c	...	...

TABLE I. *(Continued)*

Z,	Element	Meteorites	Solar photosphere	Q	Solar corona	Cosmic ray sources
69	Tm	0.034	0.06	b	···	···
70	Yb	0.126	0.16	c	···	1.2 (Z = 70 − 44)
71	Lu	0.036	0.16	b	···	···
72	Hf	0.15	0.2	c	···	···
73	Ta	0.021	No lines		···	···
74	W	0.16	10	c	···	···
75	Re	0.053	<0.01	b	···	13.4 (Z = 75 − 79)
76	Os	0.75	0.16	c	···	···
77	Ir	0.717	4.0	c	···	···
78	Pt	1.4	3.2	c	···	···
79	Au	0.202	0.13	b	···	···
80	Hg	0.4	<3	b	···	5.6 (Z = 80 − 84)
81	Tl	0.192	0.2	b	···	···
82	Pb	4	2.0	a	···	···
83	Bi	0.143	<2.0	b	···	···
90	Th	0.045	0.20	c	···	6.1 (Z = 90 − 94)
92	U	0.0262	<0.27	c	···	···

of 2 higher than the average of all measured values, while S is about the same as the measurements in C1 meteorites.

The boron value in Table I is that of Boesgaard *et al.* (1974) for Vega, while that in Fig. 1 is a modified meteoritic abundance ($B = 350$) given by Cameron (1973). Cameron now believes this meteoritic abundance to be too high (private communication to H. Reeves, 1974). We will return to this point at the end of Sec. II.B.4.

All other adopted values are based entirely on meteoritic data, whose sources Cameron (1973) references. For the majority of the elements, only Type I carbonaceous chondrites are considered. Sc, Os, and Pt come from the mean of all types of chondrites. The relative abundances of the rare earths are taken from ordinary chondrites, but their sum normalized to the value for C1's. Rh is taken from enstatites and high-iron ordinary chondrites (in the absence of C1 data); and Hg from enstatites, in view of the great variability of the C1 data.

The use of meteorites, especially Type I carbonaceous chondrites, as the primary Solar System abundance standard has been vigorously and convincingly defended by

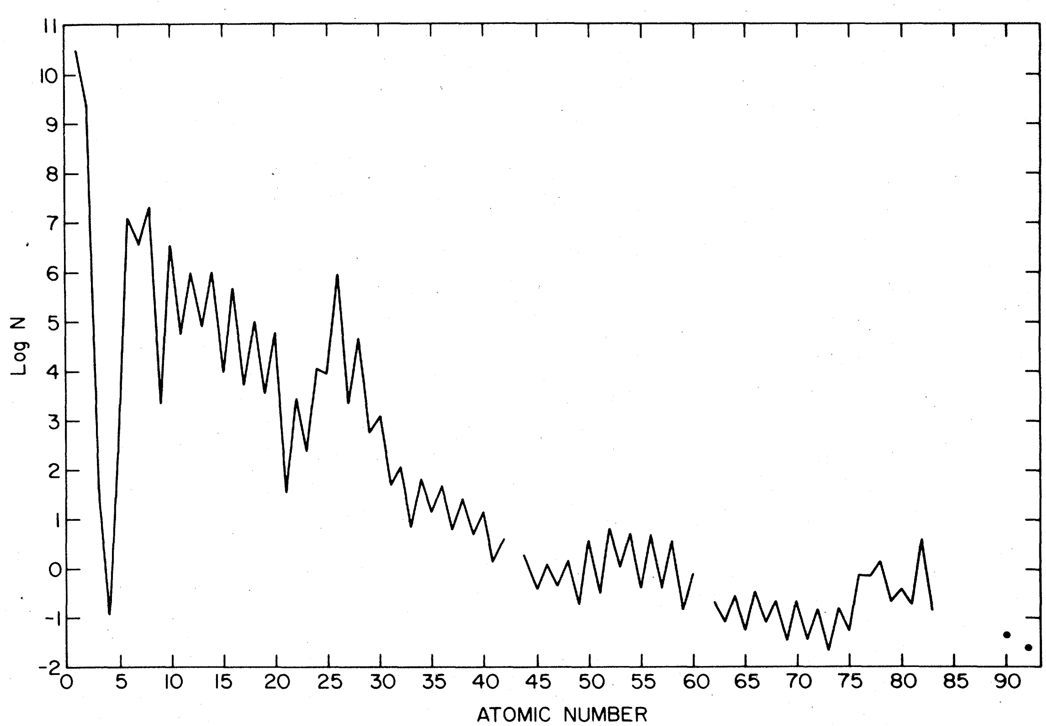

FIG. 1. The abundances by number of the chemical elements in Solar System material. The abundances are largely derived from Type I carbonaceous chondritic meteorites and are those given by Cameron (1973), except for slight revisions in the values for K, Ca, Fe, Zr, Hf, and Th which were suggested by Anders (in NATO, 1974). Boron is probably too high by a factor of about 100. Breaks in the line represent elements all of whose isotopes have half-lives short compared to the age of the Solar System. Sources of the abundances are further discussed in Sec. II.B.

Anders (1971a,b, 1972, and in NATO, 1974). Two of the lines of evidence, the agreement with solar and cosmic ray values, are discussed later. The other three, continuity of isotopic and elemental abundance trends among groups of elements with very different chemical properties, fractionation patterns among chondrites, and absence of chondrules, will require us to understand something about how meteorites are classified and how they probably formed.

All the chemical elements can be divided into five groups on the basis of the temperature at which they will condense from a gaseous to a solid phase (normally in the form of molecules of 4–12 atoms) at a pressure characteristic of the solar nebula (about 10^{-4} atm). In order from high to low temperatures of condensation, the groups are the refractories or early condensates (T_c above 1400°K), the metals and silicates (both with T_c between 1300° and 1400°K but a bit higher for the metals), and two groups of volatiles with condensation temperatures between 600° and 1300°K and below 600°K respectively. This should also be the order in which the materials condense in a gradually cooling nebula, and we will expect the composition variations between different types of meteorites to reflect these groupings.

Of the four principal types of meteorites (chondrites, achondrites, stony irons, and irons), the chondrites, which alone contain all the nonvolatile elements in roughly solar proportion, are surely the most primitive. The chondrites, in turn, can be classified in a two-dimensional scheme, based on (a) the relative proportion of the iron which is in metallic form, and several other closely correlated chemical criteria, and (b) the relative proportions and properties of chondrules and matrix. The ratio of metallic iron to total iron ranges from 80% in the enstatites (E), through 60%, 30%, and 10% in the H, L, and LL types (standing for high, low, and low–low total iron) of ordinary chondrites (O), to essentially zero in the carbonaceous chondrites (C). At least 50 elements show well-defined abundance differences among these five classes. Independently (almost) six petrologic types can be defined, called 1–6 in the direction of increasing chondrule-matrix intergrowth. There are composition variations correlated with petrologic type in the C's and E's but generally not in the O's. Meteorites with sharply delineated chondrules (typical sizes are in the millimeter range) and fine-grained (micron-sized) matrix are assigned to Types 2 and 3, while the chondrules become less distinct and the matrix coarser grained among the higher types, 4–6. These distinctions appear to reflect increasing amounts of remelting after accretion (and, therefore, presumably increasing depth from the surface of the parent body whose fragments are the meteorites). Types 3–6 occur among the E's and O's, Types 1–4 among C's. Type I comprises the class of carbonaceous chondrites which have no detectable chondrules at all (C1's). Both Roman and Arabic numerals are used for the petrologic types, the distinction between the two being comprehensible only to meteorite experts.

The two sequences E,O,C and 1–6 and their chemical and textural properties can be understood in terms of a small number of processes occurring in a definite order. They are (1) condensation of dust made of the most refractory elements (Li, Be, Al, Ca, Sc, Ti, V, Sr, Y, Zr, Nb, Mo, Ru, Rh, Ba, the rare earths, Hf, W, Os, Ir, and Pt); (2) fractionation of refractories; early condensate is gained (in the case of the Earth and Moon) or lost (in the case of E and O chondrites) in particular areas of the protoplanetary nebula; (3) condensation of metallic (P, Cr, Fe, Co, Ni, Pd, Au) and silicate (Mg, Si; this uses up only a small fraction of the available O) dust; (4) accretion of intermediate temperature volatiles (F, Na, S, K, Mn, Cu, Zn, Ga, Ge, Rb, Ag, Sn, Sb) onto the existing dust with particle sizes 1–10 μ; (5) fractionation of metal; nickel–iron is gained (Earth, Venus?) or lost (E's, O's, Moon?) from particular regions of the nebula, perhaps based on its magnetic properties, since the apparent temperature for this (800–1050°K) is close to the Curie temperature of typical Fe–Ni alloys; (6) chemical reactions with H_2S and H_2O from the accreted volatiles; (7) accretion of low-temperature volatiles (H, He, C, N, O, Ne, Cl, Ar, Cd, In, Tl, Hg, Pb) onto the existing grains; (8) remelting and chondrule formation, during or just before accretion; some of the dust is remelted (probably due to collisions), forming chondrules; volatiles are lost from these grains and their FeS converted back to metallic form; the C1's contain no detectable amount of this remelted material, but all other chondrites as well as the Earth and Moon seem to have been affected; (9) accretion to form the parent body; the material now found in each type of meteorite accretes at a well-defined temperature, highest for the enstatites and lowest for the carbonaceous chondrites, indicating the order in which they accreted; (10) shortly after formation, the interior of the parent body was reheated to about 1200°K and then cooled rather rapidly, affecting the texture of matrix and chondrules; the maximum temperature and time scale are indicated by the types of crystallization. The parent body of the presently observed chondrites fragmented about 520 million years ago, as deduced from cosmic ray exposure ages.

With this as a background, we can now understand that absence of chondrules in the C1's is evidence that the minimum of metamorphosis and chemical fractionation has occurred in them, as are the relative proportions of the various condensation temperature components in the different kinds of meteorites (the alternative of enrichment of the C1's in early condensates can probably be excluded). Finally, in light of the different times of condensation of the various components, we might expect gross abundance discrepancies between them if significant fractionation had occurred. These are not observed. The smoothness of observed abundances as we go from one group to another is shown clearly by Anders (1971b), and the lack of enhancement or depletion of the various groups relative to solar abundances is shown in Fig. 3.

Several of the other arguments that have been advanced against carbonaceous chondrite abundances can also be more or less eliminated. Early results that seemed to indicate large variations in trace element abundances from one sample to another have largely been replaced by more recent results with smaller errors bars and much more homogeneity (Krähenbühl *et al.*, 1973). The components of O and Ne with anomalous isotope ratios discussed below are in all cases a very small percentage of the element. The very large variations in Hg abundance among

C1's are confined to that component of the Hg which boils off at low temperatures and is probably terrestrial contamination from the laboratories and museums in which the samples are stored. And, finally, although the types and distributions of certain crystals indicate that the meteorites have at some time been exposed to water, we can conclude that soluble materials have not had their abundances much affected from the fact that the total ratio of Mg to Si is not correlated with the fraction of Mg in soluble form.

It must be said that not everyone agrees with the above picture of the evolution of the solar nebula. It assumes, firstly, that the nebula was homogeneous and hot enough to vaporize all solid material in the region where the meteorites were formed. Cameron (1972) has questioned this, and, in light of the discussion of oxygen isotope ratios below, it cannot be entirely true. The second assumption is that the condensation occurred in chemical equilibrium, down to temperatures of 300–400°K. Blander and Abdel-Gawad (1969) and Arrhenius and Alfvén (1969; see also Alfvén and Arrhenius, 1974, and references therein) have proposed alternative models in which the condensation occurs in a supercooled neutral gas and a highly ionized plasma, respectively.

b. Abundances of the nuclides. We will now proceed to abundances of individual nuclides, armed with whatever faith in chondritic elemental abundances the above discussion has produced. Since different isotopes of an element are sometimes produced in quite different nuclear processes, these nuclide abundances are really the fundamental data for comparison with theories of nucleosynthesis. Cameron (1973) has derived nuclide abundances directly from the chemical ones by use of observed terrestrial isotopic ratios for all elements except H, He, Ne, and Xe (which are discussed below, because several different components with different isotopic ratios can be identified) and the elements involved in radioactive decays, whose ratios have been extrapolated back to the time of formation of the Solar System. Figure 2 shows his data (apart from modifications of the H, He, and Ne isotope ratios indicated below, the changes in chemical abundances mentioned in Fig. 1, and adoptions of solar rather than meteoritic or interpolated abundances of B, Ar, S, Th, and U). The individual nuclides are plotted with symbols that indicate the nuclear processes in which they are probably produced. The deuterium abundance is the value for Jupiter, $D/H = 2 \times 10^{-5}$, found by Trauger *et al.* (1973); and $^3He/^4He = 1.5 \times 10^{-4}$. There is no evidence from either meteoritic or solar work that the terrestrial isotope ratios are not generally applicable to the Solar System, apart from the special cases in Secs. II.B.2.c and II.B.2.d below.

Many of the chapters in Mason (1971) contain discussions of the meteoritic isotope ratios of the elements in question. Apart from the rare gases, carbon and oxygen (discussed below), and elements with radioactive decay parent or daughter isotopes, there is little evidence for nonterrestrial ratios. Pb shows quite variable isotopic composition; but the ratios can always be represented as some mixture of a single primitive composition (found in regions where there is little U) and radiogenic lead. Similarly, the variable $^{186}Os/^{187}Os$ ratio is closely correlated with variations of Re/Os in the samples in the expected way. On the other hand, $^6Li/^7Li$ may be somewhat variable around the terrestrial ratio; and there is some evidence that boron and silicon show slight deficiencies of the heaviest isotope (3% and 1%, respectively). Variations in Mg isotope ratios occur between fragments of the Allende meteorite (Grey and Compston, 1974), but they are not of a type to suggest that large amounts of ^{26}Al were ever available to heat the parent bodies (Lee and Papanastassiou, 1974), nor are they correlated with the oxygen anomalies. The isotope ratios have been measured to be equal to the terrestrial to within 1% or so for S, V, Ag, Sn, Ba, and Tl, and the small variations found for S, for instance, can be attributed to chemical fractionation (Hulston and Thobe, 1965). No isotopic studies had been made for many elements up to 1971, even where techniques would, in principle, permit it, e.g., Ti, In, and Ir. The neutron activation techniques, which are used to determine the chemical abundances of the rare earths and many other trace elements, do not permit the measurement of isotope ratios as a rule. Instead, the abundance is found for a single isotope (or occasionally, two), and the total chemical abundance inferred, assuming the terrestrial isotopic composition.

c. Helium, neon, carbon, and oxygen. Isotope ratios of the rare gases in meteorites display considerable variability in an exceedingly complicated way. Ar and Kr are relatively simple; their variations are small and linear in nuclide mass, in the fashion expected from chemical fractionation. Xe is so complicated that we give it a separate section (II.B.2.d) below. He and Ne will be discussed together. The He isotope ratio is particularly important, because it is presently our only hope of getting hold of the primitive solar nebula value. The Sun has long since converted all its deuterium to ^3He, perturbing the original value of $^3He/^4He$ beyond recognition. The present solar wind value, $^3He/^4He = 4 \times 10^{-4}$ derived from Apollo and Surveyor data, (Geiss and Reeves, 1972) is, therefore, not immediately relevant to the problem at hand. (Although, if nothing else could be done, we migh *assume* He/H $= 0.1$ and $D/H = 2 \times 10^{-5}$, subtract off the amount of ^3He made from D, and arrive at $^3He/^4He = 2 \times 10^{-4}$ as the primitive value.)

The He and Ne found in meteorites (as well as the other noble gases) can have three sources, (a) production by cosmic ray bombardment, (b) radioactive decay, and (c) "other," representing gas not produced *in situ.* This third component is sometimes called "primeval" or "primitive" or "primordial," but the nonjudgmental "trapped" is perhaps a better term. The first two processes produce definite isotope ratios, and the amount of radiogenic gas will also typically bear some relation to the amount of U, Th, and K present. Their contributions can therefore be estimated and subtracted off, leaving the trapped component to be analyzed.

The trapped component in carbonaceous chondrites, in turn, must be broken down into further subdivisions. Initially, two kinds of trapped rare gases were recognized, and called "planetary" and "solar" (Signer and Suess, 1963). The solar component is systematically found on the surfaces of crystals and is believed to be due to im-

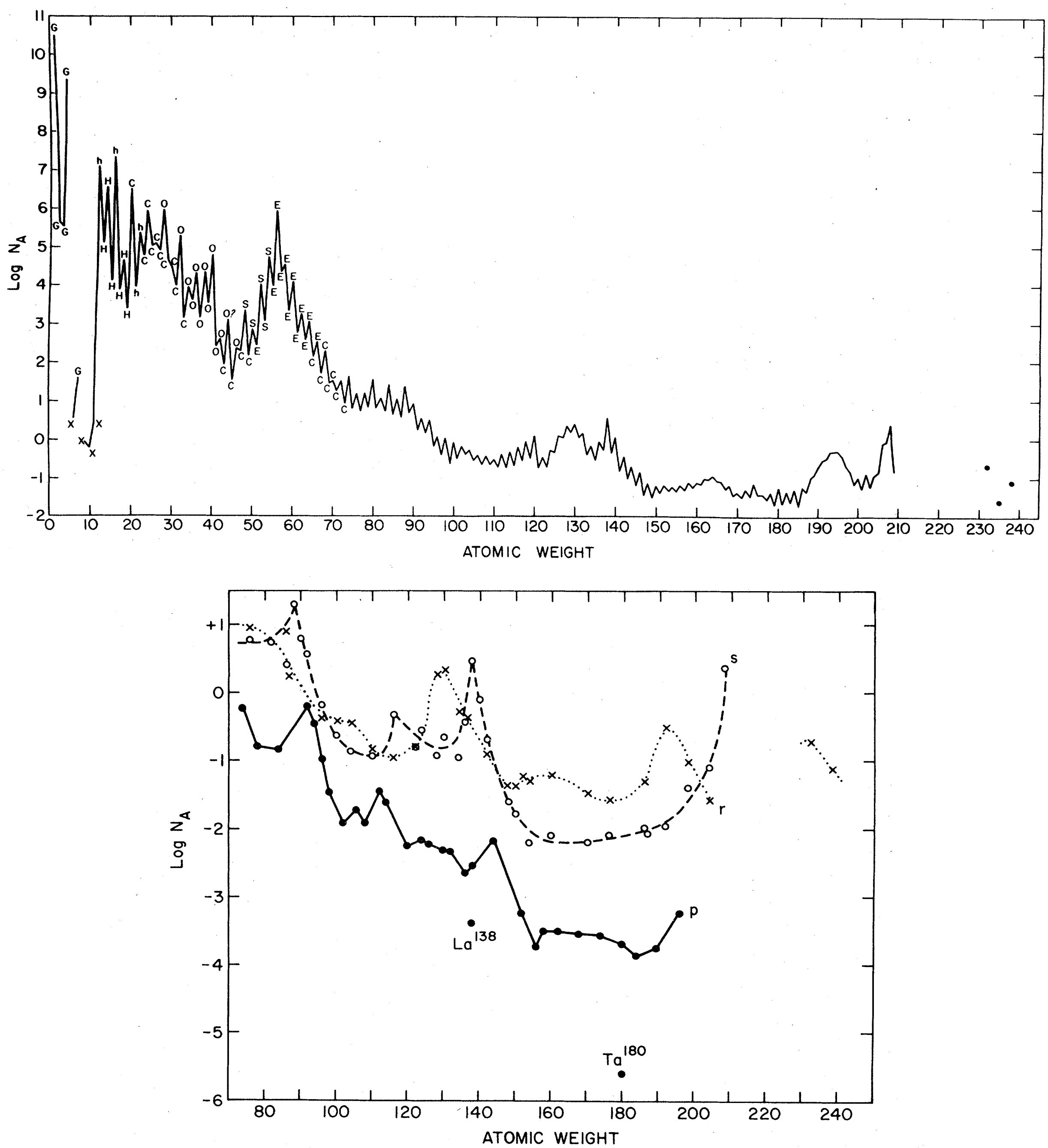

FIG. 2. (a) The abundances by number of the nuclides in Solar System material. The abundances for each value of atomic weight, A, is the sum of the abundances of all nuclides with that A. The data are those of Cameron (1973), except for (1) slight modifications in the elemental abundances of K, Ca, Fe, Zr, Hf, and Th suggested by Anders (in NATO, 1974), (2) a reduction in the ratios of ^{2}H/^{1}H and ^{3}He/^{4}He to agree with observed deuterium in Jupiter and the ^{3}He in a particular meteoritic component, (3) adoption of the solar, rather than meteoritic, elemental abundances of B, Ar, S, Th, and U, and (4) the adoption of a different neon isotope ratio, based on the work of Black (1972a,b,c). Sources of the data are further discussed in Sec. II.B. The letters indicate the nucleosynthetic processes (discussed in Sec. III) which are thought to be primarily responsible for each nuclide. G = cosmological; X = cosmic ray spallation; H = hydrogen burning; h = Helium burning; C = carbon burning; O = oxygen burning; S = silicon burning; E = equilibrium. ^{44}Ca is not adequately produced by any process. The attributions are those discussed in Sec. III. Nuclides beyond $A = 70$ are largely produced by the s, r, and p processes. Breaks in the lines represent atomic numbers with no stable isobar. (b) The abundances by number of the nuclides which can be attributed primarily to one of the processes s, r, or p (discussed in Sec. III.E). The peaks are related to the closed neutron shells at neutron numbers $N = 50$, 82, and 126. The open circles and dashed line represent nuclei produced by the slow (s) addition of neutrons; the X's and dotted line, nuclei produced by the rapid (r) addition of neutrons to iron peak seed nuclei; and the solid points and lines, nuclei produced primarily by (p,γ) and (γ,n) processes.

plantation by the solar wind. This component has a ratio of ^{20}Ne/^{22}Ne = 12.5 and ^{21}Ne/^{22}Ne = 0.0335 (Black, 1972a) and a high ^{3}He/^{4}He ratio (up to 5×10^{-4}), representing the enrichment of ^{3}He by deuterium burning. The "planetary" component, as originally described by Anders *et al.* (1970), had ^{20}Ne/^{22}Ne = 8 and ^{3}He/^{4}He = 1.25×10^{-4}, a plausible initial Solar System value. The component is called "planetary" because its isotopic ratios and proportions of the various rare gases are not unlike those in the Earth's atmosphere. It does not seem possible that this component in fact represents pure, pre-main sequence, solar nebula material in view of its very non-solar neon. We do not, therefore, have a clear measurement of the $^3He/^4He$ initial ratio. Keep in mind, for the moment, that in these carbonaceous chondrites, a low value of $^3He/^4He$ is associated with a low value of $^{20}Ne/^{22}Ne$.

Black (1972a,b,c) has examined both carbonaceous chondrites and gas-rich ordinary meteorites, and, by means of stepwise heating and careful measurements of isotopic ratios, succeeded in distinguishing five separate components in the Ne, and associated He components for most of them. His components A and B have the properties just described for the "planetary" and "solar" (wind) components. Component C may be due to implantation by solar flare particles and need not concern us. Component D, which has $^3He/^4He = 1.5 \times 10^{-4}$ and $^{20}Ne/^{22}Ne = 14.5$, is attributed by him to implantation by solar-wind-type particles before the Sun reached the main sequence and should, therefore, represent the primitive solar nebula composition. Finally, Ne–E is distinguished by boiling out of the samples only at high temperatures (about 1000°K) and by having very low ratios $^{20}Ne/^{22}Ne \leq 3.4$ and $^{21}Ne/^{22}Ne \leq 0.02$. The amount of E component present is well correlated with the amount of trapped ^{36}Ar. Components A, B, and E are found in carbonaceous chondrites and B, C, and D in gas-rich ordinary meteorites. Separating them out is helped by the fact that in the gas-rich O's, the ratios of $^3He/^4He$ and $^{20}Ne/^{22}Ne$ are anticorrelated, contrary to the carbonaceous chondrite case. Black (1972a,c) has suggested that the Ne–E component may represent grains which were never vaporized in the solar nebula. This is suggested by the high temperature at which it boils out and by the excess of ^{22}Ne, which is the sort of thing that can be expected to be produced in excess by an explosion that also makes ^{36}Ar, with which the component is correlated. The Ne–E represents only a few percent of the meteoritic neon, so that the unvaporized material is not likely to be an important part of the objects as a whole.

Cameron (1973) adopted essentially the component A isotope ratios of He and Ne. Black (1972a) regards this component as a mixture of D and E. The isotope ratios from components B and D seem to be a more suitable choice to represent the solar nebula, and they were used in Fig. 2.

Carbon also displays some variation in isotope ratio, in the sense that $^{13}C/^{12}C$ is 7%–8% higher in carbon found in carbonates than in the carbon in insoluble, aromatic polymers ("organic" material). It is possible that this variation represents derivation from two different supplies of carbon (Urey, 1967), but it now seems more probable that it is a natural result of the chemical processes that produced the complex compounds (Anders, 1971a, 1972). These bear some resemblance to the Fischer–Tropsch reaction, used industrially to produce hydrocarbons from CO and H_2 in the presence of Fe or Co catalysts, which results in similar isotope selection. The reaction has been studied in the laboratory under conditions that may be relevant to the early solar system (Anders *et al.*, 1974). The terrestrial carbon isotope ratio can, therefore, be assumed to represent the solar nebula.

The isotope effects in oxygen are a bit more subtle. Whole-rock analyses trace out nearly the same chemical fractionation curve in the $^{18}O/^{16}O$–$^{17}O/^{16}O$ plane as terrestrial and lunar materials. But the anhydrous, high-temperature phases of C2 and C3 meteorites (which occur as inclusions, fragments, chondrules, or isolated crystals) show small amounts of an additional component of nearly or completely pure ^{16}O (Clayton *et al.*, 1973). This component cannot be attributed to chemical fractionation or other processes occurring *in situ* in the meteorites. The ^{16}O component amounts, at most, to about 5% of the oxygen in the inclusions studied, and its site within them has not been identified, but it clearly reflects material which was not completely homogenized in the solar nebula. Clayton *et al.* (1973) have suggested that the ^{16}O represents interstellar grains which have had a different nucleosynthetic history from the solar nebula as a whole. The component may be associated with the slight excess in the $^{28}Si/^{30}Si$ ratio mentioned at the end of Sec. II.B.a, since ^{16}O and ^{28}Si are likely to be produced in the same nuclear processes. One would also, therefore, expect comparable excesses of ^{12}C, ^{20}Ne, and ^{24}Mg. The small variations in the Mg isotope ratios mentioned above could be interpreted as slight excesses of ^{24}Mg (Clayton, in NATO, 1974). Note that this anomalous ^{16}O component is not associated with the Ne–E just discussed, first, on theoretical grounds, since Ne–E has a deficiency of ^{20}Ne relative to ^{22}Ne, rather than the expected excess. Second, and more certain, they are found in different phases of the meteorites, the ^{16}O material occurring only in the high condensation temperature phases of C2 and C3 meteorites, while Ne–E is found in whole-rock analyses and occurs in C1's which have no chondrules.

More recent data (Clayton *et al.*, 1974) indicate a slight difference in isotope ratios between the ordinary chondrites and terrestrial samples, in the direction of the chondrites being slightly more ^{16}O-poor. This implies that at least 0.2% of the oxygen in terrestrial rocks was derived from the ^{16}O-rich component.

d. Xenonology and nucleocosmochronology. The most complicated case of all is xenon. In addition to the normal trapped, "planetary" gas, which occurs abundantly in C1's and is the source of Cameron's (1973) Xe isotope abundances, there are no less than three other, well-identified xenon components, two of which enter into age determination techniques for the meteorites. The isotopic composition of "planetary" xenon is rather different from that of xenon in the Earth's atmosphere, but quite similar to that of the solar wind, as determined from the surfaces of lunar rocks (Reynolds, in NATO, 1974). It is, therefore, presumably a suitable representative of the solar nebula.

The first anomaly to be discovered is a large excess of ^{129}Xe, which is found in whole-rock samples of numerous types of meteorites where the amount of trapped gas is not so large as to conceal it (C2–4, silicate inclusions from iron meteorites, enstatite chondrites and achondrites, H, L, and LL ordinary chondrites, and several rarer types), as well as in the magnetite from C1's (Herzog *et al.*, 1973). ^{129}Xe is the daughter of ^{129}I, which beta decays with a half-life of 17×10^6 yr. The excess ^{129}Xe therefore represents ^{129}I which was incorporated into the meteorites and decayed *in situ*.

Because ^{127}I and ^{129}I are produced in (r process) nuclear reactions in a known ratio, the observed ratio of ^{129}Xe to ^{127}I can be used as a clock which indicates the length of time between the synthesis of the iodine and the time the meteoritic material started retaining xenon. The analysis can be done rather elegantly by irradiating the sample with slow neutrons. This converts a (well-known) portion of the ^{127}I to ^{128}Xe in the same sites where the ^{129}Xe is. (The sites of xenon retention are not well understood.) Then, as the meteorite sample is gradually heated, a temperature is reached above which excess ^{128}Xe and ^{129}Xe are released in a fixed proportion for the given meteorite. This proportion is very durable; meteorite samples preheated to 1200°K and then exposed to neutrons and reheated show the same ratio as samples of the same object not subjected to this indignity. A number of meteorites have been analyzed in this way (Podosek, 1970a; Lewis, 1973). Relative dates are much easier to obtain than absolute ones, following directly from the ^{128}Xe/^{129}Xe ratio if the amount of neutron exposure is the same. Absolute ages require knowing the production ratio and are of order 10^8 yr. To first order, all meteorites that can be dated at all are isochronous, which says that, whatever event it was that caused the beginning of Xe retention, it was all over within about 14 million years. A few meteorites (e.g., Manych) cannot be dated at all; the amounts of ^{128}Xe and ^{129}Xe boiled off from different samples are not correlated (Podosek and Hohenberg, 1970). These objects appear to be unmetamorphosed assemblages of fresh chondrules, which genuinely have different "ages," in the Xe retention sense (Reynolds, 1974 and in NATO, 1974).

These ^{129}I ages clearly have some significance for our understanding of the formation of the Solar System. Just what that significance is depends on what event is being dated. Clayton (1974, 1975) has suggested that the radioactive progenitors were already trapped in interstellar grains surrounding the star in which the synthesis occurred, so that there is no significance for the Solar System. Although there is evidence for the incorporation in meteorites of unvaporized grains (Clayton *et al.*, 1973; the reader should at this point refer to the list of references and convince himself that two different Claytons are involved), this picture does not easily account for the fact that samples from a single meteorite generally define the same age to much higher precision than samples from different meteorites, or for the systematic relationships between the ages of meteorites of different types. We will, therefore, confine ourselves to the conventional view that it is Solar System events which are being dated.

Very crudely, it appears that C1–2 carbonaceous chondrites are the oldest, C3–4's next (by 1–2 million years), and most examples of other types still younger. This is in good accord with our picture of the condensation and accretion of the meteorite parent body. All types will have their clocks set by the condensation process (the time when the 400°K isotherm swept across the relevant region of the solar nebula, since magnetite is only stable below that temperature). All chondrules (in C2–6's) will have their clocks reset by the event of chondrule remelting and formation that accompanied accretion of the parent body. Finally, C3–6's will begin to retain Xe permanently only when they again cool below about 900°K after the reheating that occurred within the parent body. Since Karoonda (C4) is only 1.8×10^6 years younger than Murchison (C2) and Orgueil (C1) all the excitement must have been over within that time. This is in accord with the rapid cooling implied by crystal structures, and also requires rather rapid heating. The short cooling time scale implies small parent bodies (10–20 km radii). The only obvious cause of the rapid heating is short-lived radioactivity, but the only candidate, ^{26}Al (Urey, 1955) is not very promising (Schramm, 1971). It is not clear what we should make of all this, except to conclude that the relative sameness of meteorite ^{129}I ages probably implies that the stage of evolution of the Solar System that formed them went fairly fast.

The actual mean age between nucleosynthesis and Xe retention clearly has some implications for where and when the nucleosynthesis occurred; we discuss that later, since the age is jointly determined from ^{244}Pu.

The second anomalous Xe component was first postulated by Kuroda (1960) on the grounds that if ^{129}I survived long enough to get into the meteorites, then so should ^{244}Pu, which ought to be formed in the same processes and has a longer half-life (82×10^6 yr). The relevant component resulting from fission of ^{244}Pu was subsequently found by Rowe and Kuroda (1965). It is most easily detected in achondrites, which are considerably outgassed (so that the trapped xenon component is small) and enriched in U, Th, and the rare earths. The identification of the ^{244}Pu fissionogenic xenon has been confirmed first by comparison of the isotope ratios with those from the fission of a laboratory sample of ^{244}Pu (Alexander *et al.*, 1971) and second by the discovery by Wasserburg *et al.* (1969) that the fission xenon component in meteorites is strongly correlated with the excess fission tracks first observed by Cantelaube *et al.* (1967) in the St. Severin amphoterite. This xenon component can also be used for dating, although the situation is a bit more complicated, since no reference isotope of Pu remains in the meteorites. What is done, instead, is to use neutron irradiation to induce fission of the uranium at the same sites in the meteorite, and then boil off the U and Pu fissionogenic xenon together. Since the isotope ratios produced by the two elements are known from laboratory experiments, the two components can be separated. We thus obtain the ratio ^{244}Pu/^{238}U at the time the meteorites began to retain xenon. This event has the same ambiguity with respect to events in the formation of the Solar System as the beginning of ^{129}Xe retention discussed above.

The first measurement of Pu/U at the beginning of Xe retention, by Wasserburg *et al.* (1969), gave the value 0.033 for the whitlockite in St. Severin. This value was large enough to require a significant spike in the nucleosynthesis giving rise to Solar System heavy elements just before the formation of the Sun (Hohenberg, 1969). The problem is that, although Pu is produced by the r process in larger amounts than ^{238}U, it also decays much faster, so that the larger the amount you find, the more recent the synthesis must have been; hence the late spike. It seems, however, to have been a false alarm. An analysis of a bulk sample of St. Severin (Podosek, 1970b, 1972) gives $^{244}Pu/^{238}U = 0.013$–0.015 at the beginning of Xe retention. An excess of Th in the whitlockite confirms that the Pu enhancement there is the result of chemical fractionation (Crozaz, 1974). The evidence for the late spike in nucleosynthesis has, therefore, more or less gone away. Only a few meteorites have been dated by $^{244}Pu/^{238}U$. There are no inconsistencies with the iodine dates.

The relatively low abundances of ^{244}Pu and ^{129}I at the time of the onset of gas retention, compared to their production ratios, tell us several interesting things. $^{129}I/^{127}I$ was 1.46×10^{-4} at the time the meteorites solidified (Anders, in NATO, 1974), while the production ratio is about 1.5 (Fowler, 1972) if nucleosynthesis occurred at a constant rate over a few half-lives of the unstable isotope. Thus, only 0.01% of the ^{129}I made was still around at solidification. This implies the passage of about 1–2×10^8 yr between the last bit of nucleosynthesis and solidification (Schramm, 1973). It is interesting that this is about the same length of time as the interval between Solar System passages through successive spiral density waves. This suggests (Reeves, 1972) that the last of the nucleosynthesis occurred in supernovae of massive stars that were formed in the spiral arm passage just before the one that caused the Sun to condense out of the interstellar medium. Similarly, we find $^{244}Pu/^{322}Th = 0.006$ (Podosek, 1970b, 1972), while the ratio manufactured is 0.47 (Seeger and Schramm, 1970). This is consistent with nucleosynthesis having gone on at a roughly constant rate for several half-lives before the Solar System material separated out from the rest of the Galaxy (Schramm, 1973).

It is clearly possible in principle to add data from other radioactive nuclei with varying half-lives and trace out the entire history of galactic nucleosynthesis. This subject is generally called nucleocosmochronology. Although numerous such chronometers have been thought of (^{146}Sm, ^{176}Lu, ^{205}Pb, etc.), the only other ones so far to have yielded data interesting for the history of nucleosynthesis are the ratios $^{238}U/^{235}U$, $^{238}U/^{232}Th$, and $^{187}Re/^{187}Os$. The present vs r-process-production values of these ratios imply that the present mean age of the r-process elements is about 6.8×10^9 yr (Schramm, 1973). Since 4.6×10^9 yr of this is the age of the Solar System, during which no additional synthesis took place, the mean time scale for synthesis was 2.2×10^9 yr. This is surely not less than half the total time during which synthesis took place, implying an age for the universe of at least $(4.6 + 2 \times 2.2)$ or 9×10^9 yr. This is a sort of constraint on cosmological models (see Gott *et al.*, 1974), but the uncertainties are sufficiently large that one can really only say that the total time since the beginning of nucleosynthesis in the galaxy is between 6 and 16 billion years (Schramm, in NATO, 1974) or 9 and 15 billion years (Fowler, 1972). The average age of the r-process elements is obviously also a constraint on models of chemical evolution of the galaxy, and will be further discussed in that connection below (see also Tinsley, 1975). The data on the chronometers is summarized in Table IX.

Finally, there is evidence for one more component in meteoritic xenon. It is found in carbonaceous chondrites (where there is little U, Th, or, presumably, Pu to confuse things) with low condensation temperatures. It boils out at intermediate temperatures (600–1200°K) and is characterized by large excesses of the isotopes ^{134}Xe and ^{136}Xe. This bears all the hallmarks of xenon produced by the fission of a volatile element in the range $Z = 112$–119 (Anders and Heymann, 1969; Anders and Larimer, 1972) whose abundance was $(SH/^{238}U) = 6 \times 10^{-4}$ when the meteorites formed. Unfortunately, most nuclear physicists are no longer optimistic about the r process being able to form superheavy elements before the chain is terminated by fission (Schramm and Howard, in NATO, 1974). Even more difficult is the problem of getting the superheavies to live the 10^8 years between synthesis and incorporation into meteorites implied by the ^{129}I data. This latter problem is overcome if the anomalous xenon is contributed by the same, unvaporized grains that contribute the anomalous oxygen referred to in Sec. II.B.2.c (Clayton, 1974), but the synthesis problem is not. Black (1975) attributes this component to direct synthesis of a xenon peak at $Z = 54$, $N = 82$ in a modified r process (Sec. III.E.2).

The meteorites are not alone in displaying complicated xenon isotope ratios. The xenon from a CO_2 gas well in New Mexico has been shown to have a component from the decay of ^{129}I, and perhaps some fissionogenic xenon as well (Boulos and Manuel, 1971). As a result of continual outgassing, the Earth's atmosphere will inevitably also contain its share of these radiogenic components. It is, therefore, not surprising that Sabu *et al.* (1974) have been able to interpret the compositions of trapped xenon from various meteorites as mixtures of components which have properties and histories similar to those of atmospheric and solar-wind xenon.

Finally, it should be said that, within the conventional picture, we do not understand the origins of Ne–E, the excess ^{16}O, or the anomolous, fission-like Xe, or the sites at which the ^{16}O, iodine, and xenon are retained.

3. Solar photosphere and corona

Quantitative solar abundance work was pioneered by Russell (1929). He made use of the line intensities given in the revised Rowland Atlas and a reversing-layer model of the solar atmosphere, in which all the atoms producing the absorption lines are assumed to be at a single temperature and pressure in a layer above the photosphere which gives rise to the continuum. He correctly deduced the great predominance of hydrogen over all other elements ($Fe/H = 5 \times 10^{-5}$. The modern value is 2.5×10^{-5}.) on the basis of its rather strong lines, despite its large first excitation potential. His values relative to H for C, N, O, Si, and various other abundant elements were also within factors of 2 of modern determinations. A plot of his data clearly shows the odd–even effect, the iron peak, and the first s-process peak. Russell attributed the absence of He lines to a rather low abundance (comparable with that of oxygen) rather than to its very large first excitation potential. His values for 18 common elements were in good accord with those found for a series of bright stars (mostly giants) by Payne (1926). A couple of elements (Hf, Pt) have never been redetermined from first principles, and Russell's abundances for them are still quoted in modern compilations (e.g., Engvold and Hauge, 1974).

The next major step was the adoption of model atmosphere techniques, in which the distribution of temperature, pressure, and electron pressure as a function of optical depth is obtained from continuum observations, and the atoms producing the continuum and the lines are assumed to be intermingled. This approach was first discussed by Strömgren (1940). Since the strength of a line in this model depends on the ratio of line absorption to continuum absorption coefficient, and the latter is due mostly to H^- (Wildt, 1939), abundances relative to hydrogen result directly.

In order to obtain elemental abundances relative to hydrogen from a model atmosphere approach we need: 1. equivalent widths for lots of lines of the element, preferably weak ones (to avoid problems with microturbulence and collisional damping); 2. a good model atmosphere, giving temperature as function of optical depth in the continuum, and the velocity field as well if blended lines are to be utilized; 3. a good theory of radiative transfer and line formation and a good theory of ionization and excitation; and 4. reliable f values or transition probabilities and, in some cases, collision broadening parameters. The first requirement is met by about 20 of the most abundant elements which have permitted transitions arising from low-lying levels. At least another 20 elements have abundances which are determined from less than four lines (Withbroe, 1971), and this includes some of the more interesting ones, like Li, Be, Th, Ar, and F. High-resolution spectrometry is gradually improving this situation.

The second requirement, for an adequate model atmosphere, is well fulfilled, since continuum measurements of limb-darkening at many wavelengths overdetermine the problem (Gingerich *et al.*, 1971). Thus, a variety of models will give the same abundances to within 10% or so for weak lines observed at the center of the photospheric disk. The model atmosphere situation is not quite so good for elements that must be determined from lines in sunspots (F, Cs, Te, and the isotopes of Mg).

The standard theory of line formation (curve of growth under the assumption of local thermodynamic equilibrium) is more controversial. Curves of growth make use of a parameter called microturbulence (representing motions with characteristic size scales smaller than the mean free path of a typical photon in the line) in analyzing the equivalent widths of partially saturated lines. Viewed as a fitted parameter, microturbulence in the Sun has a value of 1–2 km/sec, implying that the lines are slightly broader than can be attributed to thermal motions alone. Worrall and Wilson (1972) have questioned the reality of microturbulence and, therefore, the accuracy of abundances derived from other than very weak lines. But there is recent observational evidence (Canfield and Mehltretter, 1973; Lites, 1973) for small-scale velocities with about this amplitude. Concerning local thermodynamic equilibrium, there is no question that non-LTE effects occur in the Sun (Mihalas and Athay, 1973). All lines formed at depths above $\tau = 1$ in the continuum will be affected. This includes both the cores of strong lines and weak lines of elements of low abundance. Even forbidden lines are not exempt: although they are likely to have LTE source functions, their opacity may be very non-LTE. The real question for us is, how much, if at all, do the non-LTE effects affect abundance analyses? The evidence is largely negative. For instance, Garwood and Evans (1974) find that the derived abundance of Cr does not change across the face of the Sun, as might be expected if non-LTE were important in this regard. Pagel (1973, 1971) points out that permitted and forbidden lines (which, while both may have non-LTE effects, will surely have different effects) give consistent results for C, O, S, Ca, Fe, and Ni. This absence of evidence should probably not be taken as proof that no important effects occur. Freire and Praderie (1974) have estimated the importance of microturbulence and various non-LTE effects under a range of conditions appropriate to stellar atmospheres. They find that abundances can be significantly over- or underestimated from either strong or weak lines.

The fourth requirement, good f values, is by far the most serious problem. The sad case of Fe was discussed in Sec. II.A.2 above. It is still the case for some of the iron peak elements and nearly all those with $Z > 40$ that the only available f values come from free-burning arcs (usually the work of Corliss and Bozman, 1962) whose properties are complicated, poorly understood, and poorly calibrated (Bell and Upson, 1971, and references therein). Many of the elements for which there are large discrepancies in the photospheric vs meteoritic abundances do not have good modern f values. Numerous groups using a variety of techniques (shock tubes, stabilized arcs, beam foil spectroscopy, etc.) are in the process of improving this situation. Recent results are nearly all in the direction of reducing the f values and therefore increasing the solar abundances. Apart from Fe (reviewed by Unsöld, 1971), typical examples include a downward revision from the Corliss and Bozman (1962) values of Ti I f values by a factor of about 1.5 for lines arising from low-lying levels and a factor of about 5 for higher levels (Roberts *et al.*, 1973a) and their similar result for V II (Roberts *et al.*, 1973b), requiring a drop in the oscillator strengths by a factor of 3 for low levels and 10 for higher levels. These are both based on beam foil lifetimes and branching ratios from gas-flow stabilized arcs. Roberts *et al.* (1975) present additional measurements for Ti I, II, and III. Analogously, Garz *et al.* (1970) find (using a high current arc in an argon atmosphere) and Lennard *et al.* (1973) confirm (from beam foil techniques) that Corliss and Bozman's f values for Ni I are, on the average, too high by a factor of 3.7, the error again being largest for lines arising from highly excited levels, while Cocke *et al.* (1973) find that Cr I oscillator strengths should be reduced by a factor of 5, and Curtis *et al.* (1973) that the Mn I oscillator strengths should be reduced by a factor of 3, with the usual dependence on excitation potential. The scandium situation is less clear. Buchta *et al.* (1971) reduce the Sc I f values by about 50%, but leave the Sc II ones unchanged, thus not resolving the photospheric–coronal discrepancy discussed later in this section. And the thulium oscillator strengths should evidently be increased (and thus the abundance decreased) by about a factor 2.4 (Curtis *et al.*, 1973). There are several reviews of the whole situation given in Bashkin (1973).

The classic compilation of abundances determined from photospheric absorption lines is the work of Goldberg, Müller, and Aller (1960). No major changes in technique have occurred since then, except the use of spectrum synthesis to improve the measurements of equivalent widths in crowded regions of the spectrum, and occasional applications of non-LTE. Extensive redeterminations have been made by Grevesse (1970, and references cited therein) and his colleagues, and by Lambert and Warner (Lambert *et al.*, 1969a,b, and references therein) and their colleagues. Many authors have considered single elements or small groups.

The second column of Table I gives the adopted solar atmospheric abundances of Engvold and Hauge (1974), who referenced the sources of all recent determinations of each element. Two changes have been made on the advice of Hauge (private communication): the Cr abundance is that of Garwood and Evans (1974), and the Ge abundance that of Ross and Aller (1974).

The adopted abundances are based for the most part on photospheric atomic absorption lines. The exceptions are He (from coronal lines and solar flare particles), N (partly from coronal atomic lines and photospheric molecular lines), Ne, Ar, and K (from coronal lines), Cs and Te (from atomic lines in sunspot spectra), and F (from molecular lines in spot spectra). The lower case letters indicate Hauge's (private communication) estimate of the reliability of the solar values; *a*-errors $\leq 30\%$, *b*-errors ≤ 3, *c*-errors ≤ 10.

A word needs to be said about the normalization of the abundances. Solar values are derived with respect to hydrogen, which supplies most of the continuous opacity, and are normally quoted on the scale $\log N(\mathrm{H}) = 12.00$, while meteoritic values are normally quoted on the scale $N(\mathrm{Si}) = 10^6$. Engvold and Hauge (1974) have converted their adopted solar values to the meteoritic scale by considering only their adopted value for Si, $\log N(\mathrm{Si}) = 7.6$, and thus multiplying solar values by 0.025 to convert them to the meteoritic scale. This is placing a great deal of faith in silicon. Cameron (1973), as discussed in Sec. II.B.2, put solar values of H, C, N, and O onto the meteoritic scale by using the average of the solar (Withbroe, 1971) to meteoritic values of eight elements. This resulted in his multiplying solar values by 0.03175 to get on the meteoritic scale. The analogous operation on the Engvold and Hauge (1974) adopted solar values results in a conversion factor of 0.0273 (or 0.029 excluding Ni), 10%–15% higher than they used. Thus, we might expect the solar values to look a bit low in the table, as a result of the normalization at silicon. This appears to be responsible for a portion of the meteoritic–solar discrepancy discussed below.

Solar abundances can also be determined from coronal and chromospheric emission lines. Both forbidden lines in the optical and permitted (resonance) lines in the far ultraviolet can be used. In the latter case, only the line intensity and the local electron density are necessary to infer the number of atoms in the upper level involved. In the former case, the line excitation mechanism (which can sometimes involve complicated resonances and wavelength coincidences) must also be known. The electron density can be determined either from the local brightness of the K (electron scattering) corona, or from high resolution radio observations. Once the population of the upper state is known, it is necessary to correct both for excitation effects and for unobserved ionization stages. Difficulties at this stage are responsible for most of the uncertainties in coronal abundance determinations. The techniques are further discussed by Withbroe (1971).

Some representative coronal and chromospheric abundances are given in the third column of Table I. Most of them are the values adopted by Withbroe (1971). A few more recent determinations from Walker *et al.* (1974a,b) and Engvold and Hauge (1974) are also included. The values are normalized to Si in the same way as the photospheric values. For the most part, the coronal and photospheric abundances are in good accord, and the best assumption may well be that they are identical and the apparent discrepancies will eventually get sorted out (Withbroe, 1971). Apart from the disagreements for argon and neon, which are meaningless since we do not really know the photospheric abundance, the differences which seem large enough to be significant are all in the sense of higher coronal or chromospheric abundances, especially for Sc, Ti, V, and Co. This is reminiscent of the situation for Fe and Ni a few years ago, and some of the coronal observers expect that the photospheric abundances will rise to match the coronal ones as additional f values are obtained (de Boer *et al.*, 1972; Nikolsky *et al.*, 1971). This has already happened to a certain extent for Mn and Cr, whose abundances increased by a factor of 2 between Goldberg *et al.* (1960) and Grevesse (1970), so that the discrepancy no longer looks so serious. Evidence that the coronal abundances may not be entirely blameless either was given by Nakayama (1972), who found that the abundances he determined at various heights above the photosphere increased with height. He was not convinced that the phenomenon had any physical significance.

Where the coronal abundances are larger than the photospheric, the meteoritic ones tend to be larger too. This could be taken as evidence either that the photospheric values are too low or that some complicated kind of fractionation has occurred.

For the rest of this discussion, "solar" abundances will be taken to mean solar photosphere, with the reservation that it may still be necessary to revise some more of the iron group abundances upward to nearer the coronal values.

We are now in a position to compare the solar and meteoritic abundances. It is at least possible, and perhaps rather probable (Pagel, 1973; Withbroe, 1971) that none of the differences are significant. But some of them are, nevertheless, rather large, and it will be useful for purposes of comparison with theory if we can decide which of the two (or more) values is likely to be more representative of the primitive solar nebula. Considering the data by groups of elements, we find that the early condensates

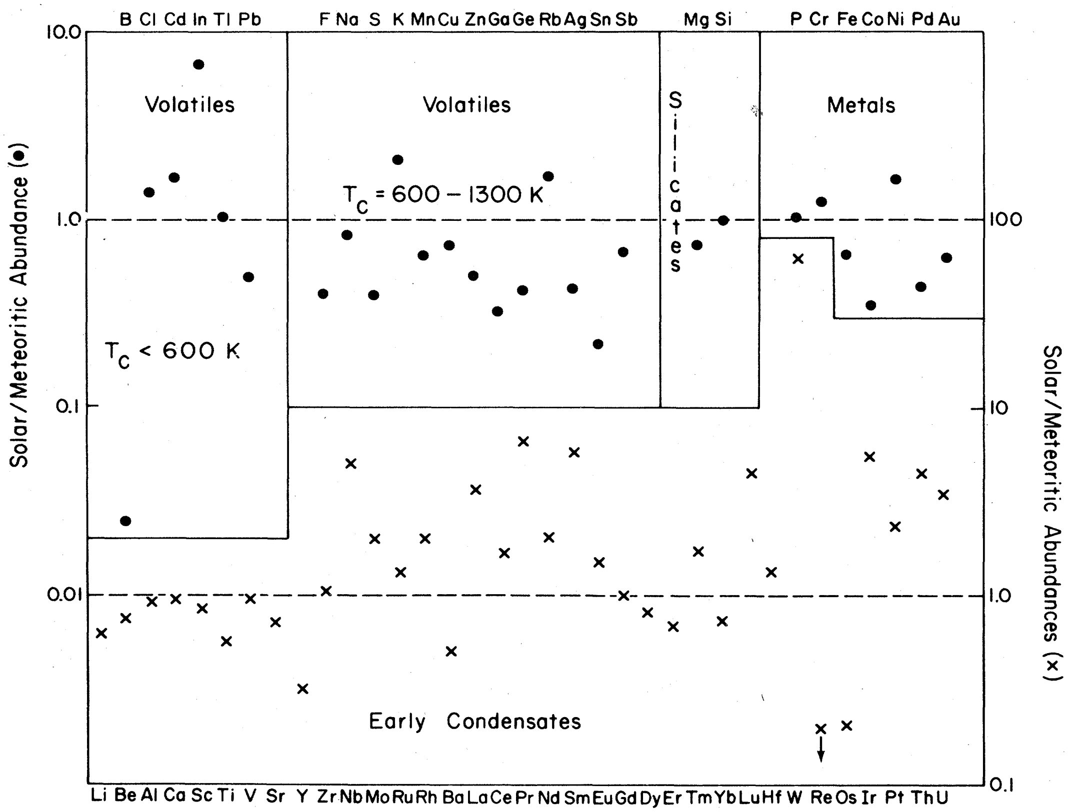

FIG. 3. A comparison of the abundances of the elements, grouped by condensation temperature, as determined from the Sun and from meteorites. The meteoritic abundances are those of Cameron (1973) and the solar those of Engvold and Hauge (1974). The tendency for the points to cluster around the line solar/meteoritic = 1 suggests that the Type I carbonaceous chondrites have not been subjected to severe chemical fractionation. The early condensates form solids above 1400°K and the silicates and metals between 1300 and 1400°K. No comparison can be made for the very volatile H, He, C, N, O, Ar, Kr, and Xe, which have been largely lost from the meteorites, or for the rare heavy elements (Tb, Ho, Ta, etc.) which have not been identified in the Sun. The only interesting upper limit is that for Re, which is shown with an arrow. The left-hand scale is to be used for the elements shown as points and the right-hand scale for those shown as crosses.

are, on average, a bit high in the Sun, and the remaining groups (metals, silicates, volatiles) a bit low in the Sun compared to the meteorites. The former effect is attributable largely to the lanthanides and heavier elements, most of whose abundances are rather poorly determined. The latter effect seems to be an artifact of the normalization at silicon and disappears if we normalize at Mg or some average of several elements. The data is shown graphically in Fig. 3, which is derived from Table I, in a fashion suggested by Anders (in NATO, 1974). The large discrepancies for Os and W and the moderate ones for F, Cl, K, and In all involve poorly determined values, and, luckily, are not critical for comparison with theory.

Argon is more worrying. Cameron's abundance is not a measurement, but an interpolation based on nuclear theory. That abundance is, therefore, easily accounted for. The solar value is based entirely on coronal lines, but it is indirectly confirmed by rather low values for the argon abundance in (a) the interstellar medium [$N(\mathrm{Ar}) = 3 \times 10^{-4}$], where it is not easily accounted for by depletion on grains (Field, 1974) since Ar is exceedingly volatile and, unlike C, N, and O, which are comparably depleted, cannot be hidden in complex molecules either (Greenberg, 1974), (b) the cosmic ray "source" composition, where $N(\mathrm{Ar}) = 2 \times 10^4$ (Price, 1973), and (c) at least some of the data on energetic solar flare particles (Price, in NATO, 1974) with $N(\mathrm{Ar}) \sim 3 \times 10^4$. By the usual standards of the abundance game, this would count as overwhelming evidence that the argon abundance should be reduced by a factor of 4–5 from the Cameron (1973) value, were it not for the utter inability of explosive nucleosynthesis theories to cope with the lower abundance (Woosley in NATO, 1974). Neon is similarly a bit lower in the corona and cosmic rays "sources" than Cameron's value (which is derived from a subset of the flare particle observations), but the consequences are not similarly traumatic for theory.

Thorium, on the other hand, seems to be somewhat more abundant in the Sun. In particular, the useful (for comparison with r-process theory) ratio Th/Pb is about 0.01 in the meteorites now (note that Cameron's values for radioactive substances are extrapolated back to the birth of the Solar System) and 0.08 in the Sun (now). Once again, there is some indirect confirmation for the solar value. Preliminary analysis of Skylab data on large Z cosmic rays gives an observed ratio of nuclei $(\mathrm{Th} + \mathrm{U})/(Z = 70\text{–}83) = 0.07$ (Price, in NATO, 1974). This is too large to be accounted for by the propagation of Solar-System-type material through the interstellar medium unless Th (and presumably U) have the higher, solar, abundance (Shirk *et al.*, 1973). The alternative is a considerable enrichment of the cosmic rays in young r-process material. In this case, the solar abundance is somewhat easier to account for theoretically (Seeger *et al.*, 1965).

We have left the most difficult cases, Li, Be, and B for last. According to the data of Cameron (1973) and Engvold and Hauge (1974), all are somewhat depleted in the Sun, Li by a factor of 250, Be by about 4, and B by at least 90. Since the light elements are destroyed by nuclear reactions at relatively low temperatures, and are, therefore, likely to have been at least partly consumed in the Sun, this, at first glance, seems all right. The catch is that Li is the easiest to destroy and B the hardest, with Be intermediate. Thus the Sun has no business losing most of its boron but keeping much of its beryllium. The simplest alternatives are that either the Be lines have been misindentified, and the solar abundance is really much lower, or that the meteoritic B abundance is somehow much too high. There are difficulties with both of these, which we will consider in turn. Lithium appears to be well behaved. The meteoritic value is consistent with what is found in very young stars ($Li/H \sim 10^{-9}$, Zappala, 1972) and the interstellar medium ($Li/H = 6 \times 10^{-10}$, Van den Bout and Grupsmith, 1973) if we allow for a little depletion, while field stars show a gradual reduction of Li/H with stellar age (Reeves, 1974, and references therein) down to the solar value for older objects.

There is some indirect evidence on Be. A variety of other stars show abundances in the range $Be/H = 4 \times 10^{-11}$ to less than 10^{-12} (Hauge and Engvold, 1968). There is no strong age correlation, but F stars on average have a bit less than G stars (Boesgaard, 1970), which may reflect a gradually declining abundance in the interstellar medium (since G stars are the older, on average). The present interstellar medium limit is less than 7×10^{-11} (Boesgaard, 1973), which can be expected to reflect some depletion on grains as well as any decrease with time that may occur. A recent detailed study of Vega (Boesgaard *et al.*, 1974) gives $Be/H = 1\text{–}2 \times 10^{-11}$, suggesting that the decrease with time is not very steep, Vega being a comparatively young star, while the presence of Be in many field G stars indicates that the nuclear destruction cannot be very great either. These stellar results are, of course, based on the same weak lines which Cameron *et al.* (1973) think may have been misidentified.

Shipman (1974) has recently redetermined the solar Be abundance from rocket uv observations and finds $Be/H = 2 \times 10^{-10}$. This is an uncomfortably high value to account for theoretically (Sec. III.B below), and E. Müller (private communication to Reeves, 1975) suggests that lines have again been misidentified.

The peripheral evidence on boron is even more ambiguous. On the "low-B" side we have an interstellar upper limit of $B/H < 7 \times 10^{-11}$ (Morton, private communication to Reeves, 1974) and the analysis of Vega by Boesgaard (1974), which gave $B/H = 10^{-10}$, as well as the solar upper limit, which Hall (private communication to Reeves, 1974) has recently lowered to $B/H < 10^{-10}$. (See also Wöhl, 1974, for another low solar upper limit.) The "high-B" side, on the other hand, is represented by the carbonaceous chondrites, for which an average observed value is $B/H = 4 \times 10^{-9}$ (Cameron *et al.*, 1973; Baedecker, 1971) for meteorites of types C1–C4. Cameron *et al.* (1973) adopted the measured value for the C2–C4's and corrected it for the relative depletion of volatiles (factors of 2–3) in these types to get a Solar System value $B/H = 10^{-8}$. Typical lunar rocks also show relatively high boron abundances, $N(B) \sim 50$ on the usual scale (Anderson, 1973). Since boron is highly volatile, and the Moon on average is even more enhanced in refractories and depleted in volatiles than the Earth and meteorites, this can be interpreted as implying a solar nebula abundance at least as high as that derived from the meteorites (Ganapathy and Anders, 1974; Anders, in NATO, 1974). Enstatites give an intermediate value, $B/H = 2\text{–}5 \times 10^{-10}$ (Baedecker, 1971). Cameron (private communication to Reeves, 1974) now believes that boron in most meteorites has been enhanced well above the solar nebula value by the same sort of process which has produced the high and variable Hg abundance in some meteorites.

Most of the theoretical discussion below will be based on the "low" values for both boron and beryllium: $B/H \sim 10^{-10}$, $Be/H \sim 10^{-11}$, but reference will be made to the difficulties encountered in accounting for the other possible values.

Solar isotope ratios, or limits thereto, have been obtained for 16 elements. None of the ratios deviates from the terrestrial or meteoritic values by more than the uncertainties, except for D/H, which has been reduced below 2.5×10^{-7} (Beckers, 1975) by nuclear reactions in the Sun and enhanced by chemical fractionation in the Earth and meteorites. The data, with references, are tabulated by Engvold and Hauge (1974).

A very general discussion of the relationships between solar, meteoritic, terrestrial, and lunar abundances and their meanings is given by Alfvén and Arrhenius (1974).

4. The solar wind and solar flares

Since the solar wind consists of a real sample of solar material brought into our vicinity, one might think that it would be the best possible source of solar abundances. Unfortunately, its atoms are only partially ionized, and differential acceleration effects are, therefore, very important. In addition, the samples available are so small that only the most abundant nuclei can be detected at all. Approximate abundances have been determined for He, Ne, O, and Fe (Hundhausen, 1972) as follows: He/H has an average value of 0.05, but is variable and largest when the solar wind velocity is largest (Hischberg *et al.*, 1972). This approximate agreement with the photospheric value at 1 A.U. from the Sun is more or less fortuitous, as the ratio probably goes up to 0.15–0.2 in the outer corona. $^3He/^4He = 4 \times 10^{-4}\text{–}10^{-3}$, which is somewhat higher than Cameron's (1973) meteoritic value. The third most abundant nucleus after H and He appears to be ^{16}O, with the ratio $^{16}O/^4He$ variable over at least the range 0.015–0.04. (The coronal value is 0.01.) ^{20}Ne has also been detected with ratios $^{20}Ne/^4He$ of 0.0016 to 0.0023 (Cameron's value is 0.0016). ^{28}Si has probably been detected, and argon, with $Ne/Ar \sim 30$. Finally, Fe has probably been found with a ratio $Fe/^{16}O = 0.17$, from one measurement (Cameron's value is 0.04). The data has been reviewed by Geiss (1972, 1973).

Particles accelerated in solar flares also bring samples directly to us. The abundances at low particle energies are quite erratic, for instance O/He can be enhanced over the photospheric value by factors from 2–5 and Fe/He by factors of 10–100. This seems to reflect acceleration effects (Sec. II.A.5 above). Above about 15 MeV/amu, some of the variations stop, and there are no obvious correlations of abundances with ionization cross section or ionization potential. There are also no differences from standard photospheric abundances of factors greater than about 2, except for argon, which is down by a factor of about 4 from Cameron's value (Price, in NATO, 1974). This relative constancy of composition has allowed Bertsch *et al.* (1972, 1973) to build up a consistent picture of solar flare abundances in the range 20–50 MeV/amu, roughly consistent with photospheric abundances where they can be compared. These solar flare measurements act as the primary standard for the Solar System abundance of He and Ne, which cannot be measured either in the photosphere (since the Sun is not hot enough to show their lines) or in meteorites (since they are volatile). The Ne/O ratio found in this way, 0.16, is rather higher than the coronal value, 0.063, found from far ultraviolet lines (Withbroe, 1971). The He/O ratio, 103 ± 10, on the other hand, is about $\frac{2}{3}$ of that in, for instance, local HII regions (Peimbert and Costero, 1969).

Additional recent data from techniques other than nuclear emulsions have considerably confused the question of solar flare particle abundances. It now seems probable that, in the energy range from 5 to 50 MeV/amu, the CNO/He ratio varies by at least a factor of 2 from flare to flare (Teegarden *et al.*, 1972), and the Fe/O ratio by a factor of 10 or more, bracketing the photospheric value (Mogro-Campero and Simpson, 1972b). On the average, Si and the iron group elements may be enhanced, relative to oxygen, by factors of 5 and 10 respectively, with the distribution in the range $Z = 10$–28 more closely resembling that of cosmic ray sources than that of the photosphere (Crawford *et al.*, 1972; Mogro-Campero and Simpson, 1972a). We should probably conclude with Teegarden *et al.* (1972) that solar flare abundances even at high energies are telling us about structure in the solar atmosphere and acceleration processes (a two-stage acceleration mechanism appears to be required; Cartwright and Mogro-Campero, 1972), rather than about the average solar composition. If so, then the general resemblances between solar flare and cosmic ray abundances may eventually help us to understand some of the puzzling features of the latter. The very great low energy enhancement (by about a factor of 120) of nuclei beyond $Z = 44$ with respect to Fe (Shirk, 1974) is particularly interesting in this respect, in light of the possible *r*-process enhancement of the cosmic rays.

The whole situation has been rediscussed on the basis of a large number of new observations by Crawford *et al.* (1975). They confirm flare-to-flare variability, even at high energies, especially for H/He and ^{3}He/^{4}He, and to a lesser extent for ratios of He:O:Fe. Their data are consistent with the conventional solar He and Ne abundances, but require the abundances of Ar and S to be lowered by factors of 3–4 from the values adopted by Cameron (1973) on the basis of nuclear systematics. Their argon abundance is consistent with the rather low value found in the solar wind by Geiss (1973).

C. Other stars

Abundances in stars other than the Sun can provide some indication of how "cosmic" our cosmic abundance distribution is, while those deviations from Solar System composition which can be attributed to nuclear processes will yield additional constraints on the sites and products of nuclear reactions and on models of the evolution of the Galaxy. The stars will be considered in two main groupings: stars which are sufficiently unevolved that their surface composition has probably not been affected by nuclear reactions within the stars themselves, and evolved stars whose surface composition probably has been affected. Data will frequently be given in the notation $[A/B]$ which means $\log(N(A)/N(B))_* - \log(N(A)/N(B))_\odot$.

We will also take this opportunity to remind the non-astronomical reader that X, Y, and Z are hydrogen, helium, and metal abundances by weight, normalized so that the sum is unity. The solar composition given in Table I corresponds to $X = 0.746$, $Y = 0.238$, $Z = 0.016$.

1. Unevolved stellar compositions

The available information is of two kinds—detailed analyses of a few stars, and statistical information on average abundances (of He, metals in general, etc.) for large groups of stars.

a. Atmospheric compositions of individual stars. Detailed abundance analyses have been done for some dozens of stars, both Population I and Population II, by either curve of growth or model atmosphere methods. These determinations suffer, in principle, from all the difficulties with f values, non-LTE, microturbulence, and so forth, discussed for the solar photospheric abundance measurements. Many of the problems, luckily, go away (especially f values) when the determination is done differentially with respect to the Sun or some other star. Microturbulence tends to be more serious, however, since very weak lines are generally not visible at the resolutions available for any but the brightest stars.

Table II gives six more or less typical examples of spectroscopic abundances of stars covering a wide range of total metal abundance, from α Cen with $[\mathrm{Fe/H}] = +0.17$ to HD 122563, with $[\mathrm{Fe/H}] = -2.6$. The stars, their spectral types, and the sources of the data are given in the table. Where the data given are averages of more than one determination, uncertainties are shown which are large enough to span the various determinations. These must be taken as optimistic estimates of the errors involved, since all authors use roughly similar methods, curves of growth, f values, and so forth, so that there may be systematic errors over and above the differences. Most authors, on the other hand, quote probable errors which are rather smaller than the differences.

Allowing for realistic uncertainties in the measured abundances, one's first impression of the table is that of monotonous regularity. Procyon and i Herc are typical of vast numbers of stars whose composition does not seem to be distinguishable from that of the Sun at all, while

TABLE II. Composition of stars whose surface abundances are thought to reflect those of the interstellar medium out of which they formed. The stars are arranged in order of decreasing total metal abundance, and where several determinations have been averaged, uncertainties are indicated which are sufficient to bracket the several values averaged. The numbers tabulated are logarithmic abundances by number relative to the Solar System, normalized at hydrogen, i.e., log $(N_M/N_H)/(N_M/N_H)_\odot$ = [M/H].

	i Herc[a] B3 V	α Cen[b] G0V + K0V	Procyon[c] F 5 IV–V	α Boo[d] K2 III	HD161817[e] A(HorBr)	HD122563[f] F III
He	−0.1	...	...	...	...	...
C	−0.15 ± 0.35	+0.32	−0.2	...	−1.1	−3.3 ± 0.3
N	+0.1 ± 0.3	...	0.0	...	...	...
O	−0.1 ± 0.3	...	−0.1	...	<−1.0	...
Ne	+0.7	...	...	...	...	...
Na	...	+0.65	0.0	−0.33	−1.3	−2.3 ± 0.2
Mg	−0.2	−0.13	−0.1	−0.19	−0.8	−2.5 ± 0.1
Al	−0.05 ± 0.15	+0.22	−0.1	−0.22	−1.4	−2.6 ± 0.2
Si	−0.3 ± 0.1	+0.05	−0.05	−0.61	−1.2	−2.0 ± 0.4
S	−0.05 ± 0.1	...	−0.1	...	...	...
Ca	...	+0.31	−0.05	−0.43	−1.3	−2.5 ± 0.1
Sc	...	+0.23	0.0	−0.72	−1.3	−2.8 ± 0.3
Ti	...	+0.29	−0.05	−0.42	−0.8	−2.6 ± 0.2
V	...	+0.37	−0.15	−0.64	−1.9	−2.1 ± 0.3
Cr	...	+0.25	−0.05	−0.52	−1.1	−2.5 ± 0.1
Mn	...	+0.25	−0.2	−0.82	−1.7	−2.9 ± 0.2
Fe	+0.15 ± 0.2	+0.17	0.0	−0.70	−1.6	−2.6 ± 0.1
Co	...	+0.29	−0.1	−0.58	−1.2	−2.7 ± 0.2
Ni	...	+0.24	0.0	−0.55	−1.7	−2.4 ± 0.2
Cu	...	...	0.0	...	...	...
Zn	...	+0.21	0.0	−0.56	...	−2.2 ± 0.3
Sr	...	+0.30	...	−0.66	−1.2	−3.3 ± 0.5
Y	...	+0.05	−0.1	−0.82	−0.6	−2.8 ± 0.3
Zr	...	+0.17	0.0	−0.75	−1.7	−3.1 ± 0.6
Mo	...	...	...	−0.69	...	...
Ba	...	−0.05	−0.1	−0.87	−0.7	−3.7 ± 0.6
La	...	+0.37	−0.3	−0.62	...	...
Ce	...	+0.18	−0.2	−0.66	...	−4.0 ± 0.3
Nd	...	+0.18	−0.3	−0.55	...	...
Sm	...	...	−0.4	−0.51	...	...
Eu	...	...	−0.3	...	...	−3.5 ± 0.7

[a] Average of Peters and Aller (1970) and Kodaira and Scholz (1970).
[b] Average of α Cen A and B. French and Powell (1971).
[c] Griffin (1971).
[d] van Paradijs and Meurs (1974).
[e] Kodaira (1973).
[f] Average of Wallerstein *et al.* (1963), Ball and Pagel (1967), Wolffram (1972), and Pagel (1965).

stars which have their heavy elements enhanced (α Cen) or depleted (α Boo, HD 161817 and 122563) seem to have them all enhanced or depleted by about the same amount.

The critical question is, which (if any) of the differences which various observers have reported should be regarded as real, apart from general excesses or deficiencies of heavy elements. Modern opinion ranges from Unsöld's (1969) view that it is obvious that none of the relative abundance variations are real to Arnett's (1971) conclusion that ratios like Na/Fe, Mg/Fe, and Al/Fe vary significantly with Fe/H, displaying the dependence of amount of odd–even effect (or alpha-rich effect) on total metal abundance that would be expected theoretically from the dependence of neutron excess on Z in hydrostatic carbon burning.

Among the abundance variations for which rather strong cases have been made are an excess of Ne among some, but not all, young B stars (Dufton, 1972; but Auer and Mihalas, 1973, attribute this to non-LTE effects), and an extra depletion of Ba among stars with very low metal abundances (Pagel, 1968). Since barium is a secondary element (made from iron-peak seed nuclei in the s process, probably in stars of relatively low mass), its relative depletion is not unexpected. The effect shows up for α Boo and HD 122563 in Table II, but the relative depletion of Ba in metal-rich α Cen and relative excess in metal-poor HD 161817 do not fit the picture. Huggins and Williams (1974) and Williams (1975) find a statistical correlation of the relative enhancement or depletion of Ba and other heavy s-process products with total metal abundance. There is considerable scatter in the correlation for Ba-peak elements, and almost no correlation for lighter s products (Sr, Y, Zr).

Other trends which have been suggested are extra depletion of Mn in stars where the total metal abundance is low (Wallerstein, 1962) and relative enhancement of [O/Fe] in low Z stars (e.g., Lambert *et al.*, 1974). The latter trend, as well as the one in nitrogen discussed below, is supported by the nature of gradients in external galaxies of N/H and O/H (see Sec. II.F).

Some evidence of more substantial relative abundance variations is found for nitrogen. Unfortunately, atomic N is generally not visible in old main sequence stars, but the strengths of CN bands can, with considerable caution, be interpreted to yield the nitrogen abundance. Pagel (1973, and in NATO, 1974) regards the observations of weak-lined halo stars as supporting a trend [N/Fe] $\approx$ [Fe/H]—that is, wherever metals in general are depleted by a given factor, the relative nitrogen abundance is low by as much again. This is what we would expect if N is a secondary nucleus, produced from ^{12}C and ^{16}O in regions where hydrogen is burned by the CNO cycle. Its high abundance in the CNO cycle results from its low proton-capture cross section. A few stars go counter to the trend, for instance HD 25329, with [Fe/H] $= -1.2$ but [N/Fe] $\gtrsim 0.5$, and no systematic trend is found among disk stars, even when their metal abundances are quite low. Pagel also suggests [O/H] $\approx \frac{3}{4}$ [Fe/H], so that oxygen is somewhat enhanced, relatively, in metal-poor stars. These results must be treated with some caution, since CN intensity is also influenced by the abundance of carbon and by how much of the carbon is tied up in other molecules. The analysis of NH lines in metal-poor stars by Sneden (1974, and references therein) is, therefore, exceedingly important. It is also quite contradictory. Sneden finds no evidence for relative depletion of nitrogen in metal-poor dwarfs, and regards all the observations as being consistent with [N/Fe] ~ 0.0 and [C/Fe] ~ 0.0 in all stars whose surface abundances are still indicative of the material from which they formed. The relative enhancement of N and depletion of C in, e.g., HD 122563, is then to be interpreted as mixing to the surface of material from a hydrogen burning shell.

We are forced, reluctantly, to the Scots verdict of "not proven" with respect to systematic relative metal abundance variations in unevolved stars. This will not be the case when we come to consider evolved stars. There is also considerable evidence for large-scale abundance variations from place to place in external galaxies, which is discussed in Sec. II.F.

b. Stellar helium abundances. A very wide variety of arguments from both stellar atmospheres and stellar interiors indicates that He/H ~ 0.1 (this corresponds to $Y = 0.28$) wherever it has not been influenced by nuclear reactions in the stars themselves. Norris (1971a) has analyzed the spectra of a number of main sequence B stars and finds He/H $= 0.09$–0.15 for all stars which do not show anomalies of the peculiar A star type discussed in Sec. II.A. Peterson and Shipman (1973), taking account of non-LTE effects, also find He/H $= 0.09 \pm 0.01$ for the B stars in three young clusters. The interior arguments are very numerous; each is given a separate paragraph below. Many of them are discussed by Iben (1974a).

It was once hoped that the flux of solar neutrinos, which is very sensitive to the central temperature of the Sun, and therefore to the present and initial central helium abundance (since, the less hydrogen there is left, the hotter it has to be to keep the luminosity going), would provide accurate measure of He/H. Unfortunately, the upper limit to the observed flux is so much lower than the predicted flux that even lowering the initial helium abundance to zero does not quite solve the problem (Ulrich, 1974). Standard models of the Sun, which match the present luminosity and radius, given the known mass, age, and Z, all have He/H about 0.1 (Sackmann, 1974).

On the negative side, we can say that normal stars ($M = 1$–$30\ M_\odot$; $Z \geq 0.01$) do not produce very much helium in the course of their evolution anyway, so that He/H must be about constant. To produce He/H $= 0.1$ in stars at a roughly uniform rate over the history of the universe would also make galaxies considerably brighter than they are seen to be in the recent past, (but see Kaufman, 1975, for indications that the energy may not come out in an observable region of the spectrum).

More positive evidence comes from the comparison of observed and calculated mass–luminosity, mass–radius, and luminosity–temperature relations for main sequence, Population I stars. The general effect of raising the helium abundance of a main sequence star is to make it more compact (since there are fewer particles to supply pressure). Thus, as can be seen from the standard homology relations, raising Y raises the luminosity some, but the effective temperature of a star even more, so that the main sequence moves to the left in the HR diagram. Raising Y therefore mimics lowering Z. Several authors (Percy and Demarque, 1967; Morton, 1968) have considered the observed mass–radium relation for unevolved stars in well-studied spectroscopic, eclipsing binary systems. The most recent such determination (Popper *et al.*, 1970) yielded $Y = 0.09$–0.14 for these stars.

The mass–luminosity relation for binaries in the Hyades has, historically, been considered to imply a high helium abundance ($Y \sim 0.4$). This is, however, dependent on using the distance that is derived from the moving cluster method (Eggen, 1969, and references therein). A variety of lines of evidence now suggest that the distance should be increased by about 15% (van Altena, 1974; Upgren, 1974). Since the masses for Hyades stars come from visual binaries, increasing the distance by 15% increases the masses by roughly 45%. Now, on the main sequence, luminosity varies as about $M^{3.5}$. Hence, increasing the distance (by increasing the calculated masses) increases the luminosity which the stars can be expected to produce by a large factor. The observed luminosities then become consistent with a lower helium abundance ($Y \sim 0.3$), the higher mass taking the place of high helium in making the stars bright enough to be consistent with observation.

The HR diagrams of both young and old galactic clusters are also consistent with He/H ~ 0.1, but the method is not very sensitive (Iben, 1967; Aizenman *et al.*, 1969; Demarque and Schlesinger, 1969; and reviews by Danziger, 1970, and in NATO, 1974). Some of these lead to rather high values of either Y or Z, but Iben (in NATO, 1974) now believes the observations all to be consistent with $Y \sim 0.3$.

Finally, among young stars, the period–luminosity–color relation observed for Cepheids can be compared with pulsation theory calculations to derive the helium abundance. A typical result, found for the Small Magellanic Cloud Cepheids, (Iben, in NATO, 1974) is $Y = 0.29 + 1.7(19.25 - DM) - 15(T_b - 3.8)$, where DM is the distance modulus of the Galaxy, and T_b is the log of the effective temperature of the blue edge of the instability strip. For the most probable values of the distance modulus and the edge of the instability strip, $Y = 0.29$, and $Y = 0.28$

$\pm$ 0.1 for Cepheids in the Galaxy, Andromeda, LMC, and SMC (Iben and Tuggle, 1975).

Comparison of observation and theory for Population II stars also leads to various determinations of the helium abundance. The mere existence of F and G subdwarfs says that the shifting to the left of the main sequence that is caused by the lower Z value is not compensated by a lower Y. From careful photometry of cool, high velocity stars, Cayrel (1968) has concluded that there are indeed, late type main sequence stars with low metal abundance and normal helium.

Gross (1973) has compared spectroscopically determined values of surface gravity and effective temperature of horizontal branch field stars with zero age horizontal branch models (that is, models of stars which have just started core helium burning). He finds $Y = 0.3$ for these stars (and 0.23 for 7 open clusters).

Globular clusters, since they present us with groups of stars all with the same age and composition (we trust), are amenable to a variety of statistical arguments, in which the time scales, colors, and luminosities along evolutionary tracks are compared with counts of stars at various positions in the cluster HR diagram. For instance, Iben (1968) and Iben and Rood (1969a) found that the relative numbers of blue and red horizontal branch stars in M15 implied $Y = 0.29$. This determination is dependent upon the assumption that all the horizontal branch stars have the same mass and the same age to rather higher precision than is, perhaps, justified.

The ratio of luminosities on the horizontal branch and at the main sequence turnoff point is sensitive to helium abundance, but also to age and Z. Iben and Faulkner (1968) concluded that $Y = 0.3$ was consistent with all the data from several clusters.

The absolute luminosity of the horizontal branch is relatively insensitive to age and semiconvection, since it depends only on the mass of the helium core (not the convective core). Unfortunately, an accurate distance for the cluster is then required. Iben (in NATO, 1974) finds $Y = 0.27 + 0.012\ (\log Z + 3)$ from this method.

The comparison of theoretical and observed luminosity functions also determines Y only if you know the cluster age. Relevant evolutionary tracks are given by Simoda and Iben (1970) and Simoda (1972). Simoda's analysis of the data implies $Y \geq 0.2 \pm 0.1$.

The ratio of the number of red giants to the number of horizontal branch stars in a globular cluster is a function of the respective lifetimes of the two stages, and, therefore, of the initial helium abundance. Unfortunately, it also depends on the treatment of semiconvection, so that the final result, $Y = 0.3 + 0.4 \log_{10}(R/f)$ (Iben, in NATO, 1974), not only involves $R = N(\text{hor. br.})/N(\text{red gi.})$, but also f, a fudge factor, which is unity if both semiconvection and convective overshoot are neglected, but 2 if both are included. The resulting value of Y, given that the observed R is about 0.9, is between 0.16 and 0.30. If we knew Y independently, we would learn something about convection this way, which might be even more useful. Other authors (e.g., Demarque *et al.*, 1972; Tarbell and Rood, 1975) have gotten low values of the helium abundance this way ($Y \sim 0.13$), but it does not seem to be necessary.

Finally, the theory of pulsational variables can be compared with observed properties of RR Lyrae stars to determine Y for clusters in which they occur. There is a continuing problem in all such efforts that the masses of these stars calculated from their pulsation properties are a bit smaller than those obtained from evolutionary tracks (Iben and Rood, 1970). Fortunately, the derived helium abundance is not too sensitive to this. A typical analysis is that of Tuggle and Iben (1972), who find $Y = 0.22 \pm 0.09$, from the position of the blue edge of the instability strip. But the real range of uncertainty is, perhaps, better illustrated by comparison with the result of Sandage (1969), who finds $Y = 0.32 \pm 0.09$ from similar data and similar pulsation models. Recent results also by Iben (in NATO, 1974) for four clusters give $Y = 0.23$ from $Y = 0.22 + 5(\log T_{\text{ed}} - 3.863) + (\log P_{\min} + 0.55)/(1.5 + Y)$, where T_{ed} is the temperature at the left-hand (blue) edge of the instability strip and $P_{\min}$ is the period, in days, of the RR Lyrae stars at the edge. The main uncertainty is in the masses of the stars involved. Consistent results come from the ratio of the periods of RR Lyrae stars pulsating in the fundamental and first overtone modes. The value implied is $Y = 0.27 + 0.012(\log Z + 3)$.

On the basis of the data available, it is very hard to exclude the null hypothesis that He/H ~ 0.1 for all unevolved objects which have been studied. It is also impossible to exclude the views of Mezger and Iben (in NATO, 1974) that $Y \sim 0.28$ for young objects and $Y \sim 0.23$ for objects formed early in the history of the galaxy. A third view, that helium abundance varies quite widely from place to place is also defensible, and firmly defended by Fowler, Hoyle, and others (private communications, 1975), who emphasize the apparently low He/H ratio at the galactic center (Sec. II.A.3) and in II Zw 40 (Sec. II.F.1) and the ambiguous nature of the stellar and solar evidence discussed above. Since the solar helium abundance is uncertain by at least a factor of 2 (Hirshberg, 1973), it does not provide a useful discriminant, but Fowler believes that the solar Y is nearer 0.1 than 0.2 on the basis of coronal and solar flare data and the difficulty in accounting for the low solar neutrino production rate in a high helium model.

c. Stellar metal abundance as a function of age. Although there are no unequivocal variations of the relative abundances of the heavy elements in unevolved stars, there is no question that the observed total metal abundance varies over almost a factor of 1000. The fact that the stars with very low metal abundances ([Fe/H] ≤ -1.5) all belong, by either location or dynamics, to the halo population, which presumably formed before the Galaxy had collapsed from a spheroid to a disk, suggests that we should look for a detailed correlation of Z with the ages of stars. This correlation (which may or may not also depend on position in the galaxy) should, then, tell us how Z in the interstellar gas has varied over the history of the galaxy, and will be a constraint on models of chemi-

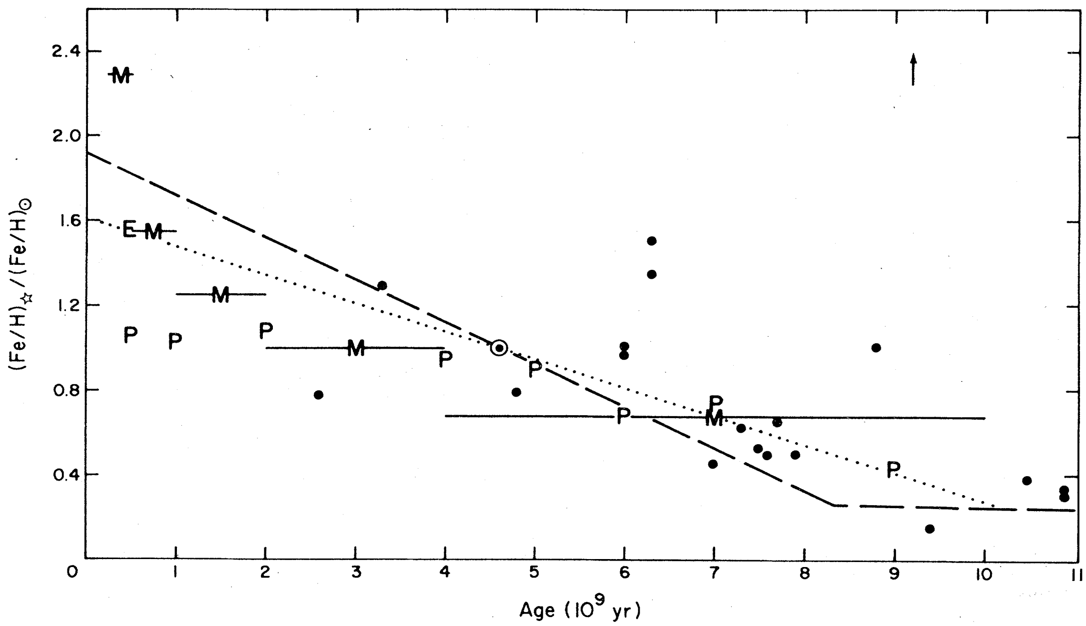

FIG. 4. The distribution of metal abundances of F, G, and K dwarfs in the solar neighborhood as a function of their age. The vertical scale is the ratio Fe/H by number in the stars divided by that in the Sun, which is represented by the circled dot. The M's represent the mean points for groups of stars observed by Mayor (1974). The bars indicate the spread in ages of the stars contributing to the means. The E is Eggen's (1970) data for the Hyades, and the P's are mean points derived by Powell (1972), each point representing the average metal abundance of a number of stars with ages varying by $(0.5\text{–}2.0) \times 10^9$ yr. Each mean point should be regarded as having 0.05 or so of statistical uncertainty in Fe/H, as well as whatever systematic errors there may be. The dots represent individual stars studied by Hearnshaw (1972). The arrow in the upper right corner is 30 Aql, whose metal abundance is 3.3 times that of the Sun. The dashed line is the mean relation of age and metal abundance suggested by Clegg and Bell (1973), and the dotted line represents the less steep increase in metallicity with age suggested by Pagel (in NATO, 1974). The slope of this line corresponds to $dZ/dt = 0.002$ per 10^9 yr, starting from an initial $Z = 0.004$ for the oldest disk stars.

cal evolution. There are, clearly, two difficult observational problems involved. First, we must find a spectroscopic measure of metal abundance which can be used at low enough resolution to be applied to a large number of stars. Second, we need some way of determining the ages of the stars whose metal abundances have thus been measured. The most usual approach is via wide or intermediate band photometry. For instance, uvby colors and their combinations can be interpreted to yield a metallicity index, effective temperature, and absolute visual magnitude. Then the height of an individual star above a theoretical zero age main sequence gives its age from calculated evolutionary tracks or isochrones. This is essentially the approach followed by Mayor (1974), Powell (1972), and Eggen (1970), whose data is shown in Fig. 4 and discussed below. Hearnshaw (1972), whose data is also shown in Fig. 4, has done differential curve of growth analyses for 20 individual elements for a number of field G stars to get the metal abundance.

Mayor's data is shown as M's, with bars indicating the spread in ages of the stars contributing to the means. Powell's data is shown as P's, each point representing the mean of the metal abundances of a number of stars with ages varying by $(0.5\text{–}2.0) \times 10^9$ yr. All mean points should be regarded as having 0.05 or so of purely statistical uncertainty in [Fe/H], in addition to whatever systematic errors there may be. Hearnshaw's points for individual stars are shown as small dots; the arrow in the upper right-hand corner represents 31 Aql, whose metal abundance is 3.3 times that of the Sun. These points all represent A, F, and G disk-population dwarfs within a few dozen parsecs of the Sun, which is shown in the usual fashion. E is the Hyades, from Eggen's work. Clegg and Bell (1973), who have also studied a large number of nearby dwarf stars, regard all their data as being reasonably consistent with a straight line, $[\mathrm{Fe/H}] = \log(T_9/5)$, where T_9 is the time elapsed (in units of 10^9 yr) since collapse of the Galaxy to disk was completed, with $[\mathrm{Fe/H}] = -0.6$ at that time. This relation, normalized to the Sun, is shown as a dashed line. M67 and NGC 188 are both too far away to be included in the samples, but both would lie above the mean points.

The various mean points present a general impression of a monotonic increase, with no peak in the past and no flattening off in the present. [Mayor's high point for very young stars is, according to Pagel (in NATO, 1974), probably contaminated by Am stars.] There is clearly considerable scatter around any mean relation. Most of this is probably real, and some of it is undoubtedly related to the phenomena of supermetallicity (discussed in Sec. II.c.E) and gradients in abundance across the Galaxy (discussed in Sec. II.F). Clegg and Bell's mean relation is not grossly inconsistent with any of the data, but Pagel believes a slightly less steep increase in metallicity, like that shown by the dotted line, to be more representative. The slope of this line corresponds to $dZ/dt = 0.0022$ per 10^9 yr, starting from an initial $Z = 0.004$ for the oldest disk stars. It is obviously impossible to exclude much more complicated curves than straight lines, but the data do not seem to warrant more complicated treatment.

The main properties of $Z(t)$ appear to be a monotonic increase with time, at least in the immediate solar neighborhood, and a nonzero initial value, accompanied by considerable scatter. We will see in Sec. IV that all "reasonable" models of galactic evolution are able to reproduce this sort of $Z(t)$.

d. The number of stars as a function of metal abundance. A model of the evolution of the Galaxy must also be able to account for the total numbers of stars with various metal abundances seen at the present time, as the sum of all surviving stars produced at all stages in the past. This will turn out to be a serious difficulty for many models, because of the very small number of stars with very low values of Z. The problem was first realized by van den Bergh (1962) and Schmidt (1963), who proposed as a solution a model of the Galaxy in which early generations of stars consisted largely of massive stars which have not survived to the present time. Other solutions are discussed in Sec. IV. Comparison with theory will be easiest if we limit our data to main sequence stars of late enough types that all the ones that have ever formed are still there (later than G2).

Relevant data are available from a variety of sources, including Bond's (1970) objective prism survey of several hundred thousand stars (which includes all spectral types, but has no proper motion bias); the Yale Bright Star Catalogue (unfortunately complete only to $m_v = 6.5$, providing a sample of only 30 dwarfs with UBV photometry later than G2); Eggen's (1964) survey of a limited part of the sky down to $m_v = +9$ (which is likely to contain some unrecognized subgiants); Clegg and Bell's (1973) survey of F and G dwarfs within 100 parsecs of the Sun (which does not go quite late enough); and the catalogs of stars within 20 and 25 parsecs of the Sun by Gliese (1957) and Woolley *et al.* (1970).

Figure 5 includes points from the data of Bond (B's); Clegg and Bell (C's; F9–G2 stars only); Gliese (G's; as compiled by Pagel, 1973, stars later than G2 only); Eggen (E's; stars later than G0); and the data available to Schmidt in 1963 (S's). The figure gives as its ordinate the log of the fraction of all stars in each sample which have [Fe/H] $\leq$ the value in the abscissa. Highest weight should probably be given to the Gliese data since they represent a complete sample, should be uncontaminated by giants and stars of earlier spectral types than G2, and incorporate a modern calibration of [Fe/H] in terms of the measured photometric quantity $\delta(U-B) = (U-B)_\star$

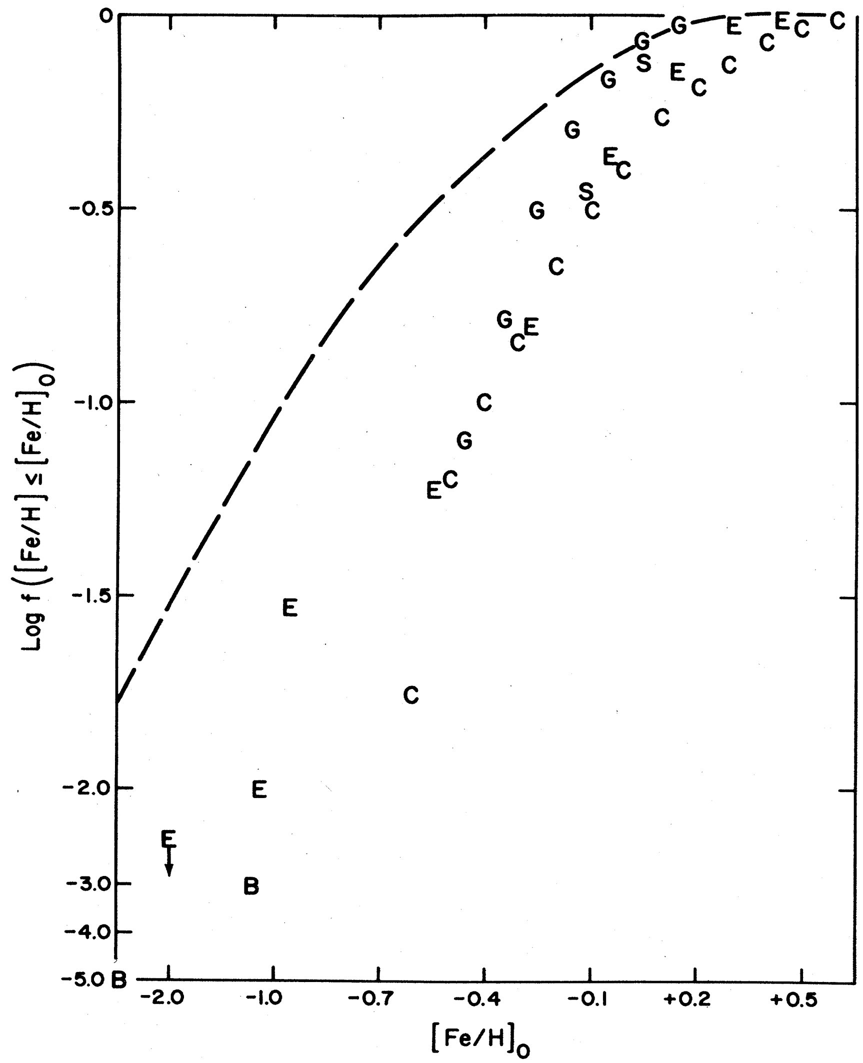

FIG. 5. Fraction of F, G, and K dwarfs in the solar neighborhood having metal abundances [Fe/H] less than particular values $[Fe/H]_0$. The S's are the data available to Schmidt (1963); the B's are the points from Bond's (1970) objective prism search for stars of very low metal abundance; the C's are the F9–G2 stars studied by Clegg and Bell (1973); the E's are Eggen's (1964) stars later than G0; and the G's represent the catalog of stars within 20 parsecs of the Sun, compiled by Gliese (1957), as analyzed by Pagel (1973). The dashed line represents the prediction of a typical naive model of galactic evolution (e.g., Pagel, 1973b) and requires the presence of far more stars of low-metal abundance than are, in fact, seen in the data.

$-(U-B)_{\mathrm{Hyades}\odot}$. Considering the differences in the various samples and calibrations, the various sets of data agree quite well, and the differences are in the direction that would be expected in light of the different spectral types included and the behavior of $Z(t)$. Thus, Gliese's data, for stars later than G2 only, have the lowest fraction of metal-rich stars; and Clegg and Bell's data, for F9–G2 stars, the highest fraction of metal-rich stars. The Eggen points, for stars later than G0, fall in between. Since Bond has considered all spectral types, the "B" points should be regarded as a continuation, most nearly, of the Clegg and Bell data. That is, a sufficiently large sample of stars limited to G2 and later would be expected to show a slightly larger fraction of very low Z stars.

e. Supermetallicity. Spinrad and Taylor (1969) originally called attention to a class of relatively old giant stars for which narrow band photometry seemed to indicate metal abundances significantly higher than in the Hyades, $[\mathrm{Fe/H}] \geq 0.2$. Taylor (1970) found the same phenomenon in some nearby G and K dwarfs. The stars affected include members of the oldest known open clusters, M67 and NGC 188. If stars like this are common either in the solar neighborhood or elsewhere in the Galaxy, then our picture of gradual, uniform increase in metal abundance with time will obviously have to be abandoned. It is not, therefore, surprising that their work has been heavily criticized by Eggen (1971), Strom *et al.* (1971), and many others. Some reconciliation of the various points of view is possible.

It is clear that there do exist a certain number of old dwarfs in the solar neighborhood whose abundances of iron and other metals are found to be high by any of the standard techniques. 31 Aql, which has $[\mathrm{Fe/H}] \sim 0.6$ and is at least as old as NGC 188 (25×10^9 yr), and HR 1536 and 4587, which have $[\mathrm{Fe/H}] = +0.2$ and other metals enhanced by slightly larger factors, belong to this class (Hearnshaw 1972), as do HR 7670 and δ Pav with $[\mathrm{Fe/H}] = 0.3$ (Spinrad in NATO, 1974). Stars of this type are, however, a sufficiently small proportion of the population in the solar neighborhood that they do not greatly affect the general trend of $Z(t)$ shown in Fig. 4. The large ages implied by the HR diagram positions of these stars are confirmed by the weakness of their Ca II emission. (Line reversal at the center of calcium H and K is an index of chromospheric activity and, therefore, of stellar age.) These stars tend to have large orbital eccentricities and large velocities in the radial direction in the Galaxy, implying that they really "belong" to a part of the galaxy nearer the center than we are. In light of the strong radial abundance gradients found in external galaxies, there seems to be some sense in all this.

The giant situation is more complicated. About 6% of the local G and K giants have stronger than normal CN bands (Janes and McClure, 1971). They are all either in young moving groups or high velocity stars. CN strength is normally thought of as a general metallicity criterion, but it can also be increased by mixing to the surface of material which has been in a hydrogen burning zone, since C and O are partially converted to N there. Since virtually all giants whose carbon isotope ratios have been studied show enhanced ^{13}C (also a product of hydrogen burning by the CNO cycle), significant mixing cannot be excluded. Harmer and Pagel (1973) find that two of the classic "super metal-rich" giants, ϕ Aur and μ Leo, do have strong CN for their iron-line strength. Strom *et al.* (1971) have pointed out that one of the effects of strong CN lines is to lower the boundary temperature of the stars. The unusual temperature vs optical depth structure has the effect of making elements like Ca, Na, Mn, and V, whose lines are formed high in the atmosphere, look overabundant. These are among the elements that Spinrad and Taylor (1969) based their conclusions on. Extra atmospheric turbulence has a similar effect. Strom *et al.* (1971) conclude that four of the standard SMR giants (ϕ Aur, μ Leo, λ Hya, 20 Cyg) have essentially normal metal abundances, but that μ Aql has $[\mathrm{M/H}] \sim +0.3$. A similar result is obtained for μ Leo by Blanc-Vaziaga *et al.* (1973), who find, using model atmosphere techniques, that Na, Ca, Mn, and Cu look enhanced by factors of 2–3 and Ba depleted by a comparable amount. They conclude that supermetallicity is very selective in its elements. Much the same set of elements has extra enhancement in SMR dwarfs like HR 4587. Williams (1971a,b) prefers to regard these stars as being genuinely metal-rich in preference to overly turbulent, but does not discuss the effect of boundary temperature. The strengths of the weak iron lines in μ Leo appear to be consistent with either normal abundance (R. Peterson, quoted by Spinrad in NATO, 1974) or $[\mathrm{Fe/H}] \sim +0.4$ (Gustafsson *et al.*, 1974), depending on one's choice of surface temperature, the details of the atmospheric model, and whether measured equivalent widths or narrow band photometry is used. We conclude tentatively that the number of metal-rich giants in the solar neighborhood whose surface abundances have not been influenced by mixing is probably small. Stars like μ Leo tend to have high eccentricity orbits, which, as in the case of the dwarfs, may tell us that they belong to the inner parts of the Galaxy.

Even more attention has been given to the old clusters, M67 and NGC 188. Pagel (1974) has concluded that the strengths of the CN bands in both are consistent with metal abundance in the giants between that of the Sun and that of the Hyades, as would be expected from their ages (about 3×10^9 and 5×10^9 yr, respectively). Barry and Cromwell (1974) also find an intermediate metallicity for M67 ($[\mathrm{Fe/H}] = 0.06$) based on low dispersion spectra of dwarfs. Bessell (1972) has interpreted the H-line profiles and Ca II K-line strengths in two F stars in M67 as implying a metal abundance rather lower than solar, and Bond and Perry (1971) found $[\mathrm{Fe/H}] \sim -0.5$ from intermediate band photometry. C. Griffin (in NATO, 1974) also concludes that metals are deficient by factors of 1.5–3 from a differential curve of growth analysis of one red giant (IV-202). The uncertainty is largely a product of the uncertainty in effective temperature; the hotter the star, the larger the metal abundance required to produce the observed line strengths.

Some of the disagreement on the metal abundances in SMR stars has been explained by Oinas (1974), who finds that neutral and ionized metals genuinely give discordant results when analyzed in the conventional way. He concludes that there must be some extra source of opacity in the atmospheres of these stars and confirms the very selective nature of the metal enhancements found by Blanc-Vaziaga *et al.* (1973). Perrin *et al.* (1975) present yet another view of the SMR phenomenon.

Because CN line strengths depend on the C/O ratio and the details of stellar chemistry as well as on the amount of N present, claims of high metal abundance based on CN alone should be viewed with suspicion.

All the data presently available appear to be consistent with the view that stars of any given age have a roughly Gaussian distribution of abundances, whose tail reaches to or above the present average metal abundance in the solar neighborhood, but that there was not an epoch in the past during which the average metal abundance of the stars formed was higher than it is now. There may, however, be SMR *places*.

2. Evolved stars

The discovery of technetium (whose *s*-process isotope, ^{99}Tc, has a half-life of 2×10^5 yr) in giants by Merrill (1952) made it virtually certain that the surface composition of stars can sometimes be influenced by nuclear reactions in the stars themselves. We will consider the abundance anomalies in evolved stars roughly in the order of increasing separation from the main sequence.

a. Carbon stars and other giants. Single giants with nonsolar compositions come in two groups, with and without a roughly normal amount of hydrogen. We discuss the former first; they are somewhat easier to analyze, at least at higher surface temperatures, since the source of the continuous opacity is H^- as usual. The hydrogen-rich stars, which are hot enough for detailed abundances to be determined, can be roughly devided into (1) classical carbon stars of spectral types R and N which seem to be rich in nitrogen and poor in oyxgen resulting in strong CN and C_2 bands (usually with a high $^{13}C/^{12}C$ ratio), (2) S stars, rich in carbon and the heavy elements like Y, Zr, Sr, Ba, and sometimes Tc, which are believed to be produced by the *s* process (the use of capital S for the spectral type and lower case *s*, standing for "slow," for the process is a coincidence), (3) barium stars, again with *s*-process enhancements but somewhat hotter than the S stars, (4) CH stars, both giants and subgiants, with enhancements of CN, C_2, Ba, Sr, La, Ce, Nd, and so forth, superimposed on a weakness of most metallic lines consistent with Pop II Fe abundance. Transition objects between these types and between these types and normal M giants also occur and are called MS, SN, Ba–R, and so forth, stars. The peculiarities which are conspicuous in these objects—variation in C:N:O and $^{12}C/^{13}C$ ratios, and enhancement of *s*-process elements—have also been found, in milder form, in more normal looking giants like Arcturus, α Per, and o Vir.

Very little quantitative chemical information exists for the classical carbon stars. The lore of determining their temperatures, luminosities, masses, and so forth is reviewed by Wallerstein (1973). The fundamental problem is that the surface temperatures are so low that H^- is not the major source of continuous opacity, which is, rather, due to a large number of different kinds of molecules. A good understanding of the continuous opacity and good model atmospheres will eventually make possible complete abundance analyses. The work of Querci *et al.* (1974, references therein, and work in progress) is a major step forward in this regard. Meanwhile, the available information is limited to: (1) oxygen is greatly, and carbon slightly, depleted (so that $C/O > 1$) and nitrogen greatly enhanced (e.g., Querci and Querci, 1970), (2) most metals, from Na to Nd, are roughly normal (Green *et al.*, 1972 for HD 156074) in many of them, but elements like Zr, La, Pr, Nd, Sm, and Gd may be overabundant in others (Fujita and Tsuji, 1964, 1965 for Y CVn), (3) enhanced lithium (up to $Li/H = 10^{-7}$) is often (Merchant, 1967) and Tc sometimes (Peery, 1971) found. The "carbon star" abundance in Table III is that of Querci *et al.* (1974), and is derived from a variety of sources and represents the mean of a number of stars.

More quantitative information can be obtained for the hotter stars. Table III includes an entry for a barium star (HD 116713, analyzed by Danziger, 1965a), and a CH star (HD 198269, analyzed by Lee, 1974). A new class of CH subgiants has been identified by Bond (1974). Their abundances are very similar to the CH giants, and the subgiants are probably the immediate precursor of the giants. The S stars are, again, inconveniently cool for detailed analyses. They differ from the barium stars in that many of them show Tc and enhanced Li, which the Ba's do not (Peery, 1971; Boesgaard, 1970b), and are similar in showing strong lines of the *s*-process elements and in having Zr/Ti enhanced by an order of magnitude or more (Boesgaard, 1970a). The relative abundances of the enhanced *s*-process elements in the barium stars are in good accord with the theory of the process (Warner, 1965; Allen and Cowley, 1974). The barium stars appear to come in both normal carbon and high carbon varieties (Alexander and Branch, 1974), with the types more correlated with how the stars were discovered than with any obvious physical parameter.

The measurement of isotopic ratios in the carbon stars is reviewed by Wallerstein (1973). Enhanced ^{13}C is commonly, but not universally, found in the CH and R and N stars and not in the barium stars. The very low $^{12}C/^{13}C$ ratios found by Wyller (1966) for U CVn and WZ Cas (2.4 and 2.0, respectively) should be taken with liberal amounts of salt (Fujita, 1970), though, contrary to popular superstition, they are not impossible on theoretical grounds, since the CNO cycle does not always have to be in equilibrium. Ratios in the range 4–10 seem well established. Infrared observations of the molecular cloud around the carbon star IRC + 10216 confirm the high ^{13}C abundance found from radio observations and also indicate an enhancement of $^{17}O/^{18}O$ by a factor of at least 25 (Rank *et al.*, 1974, and references therein).

A similar enhancement of ^{13}C has been found in virtually every giant studied, including those which seem otherwise normal. For instance, Griffin (1974) finds $^{12}C/^{13}C = 5.5$–6.0 for Arcturus, and Ridgway (1974) finds $^{12}C/^{13}C = 10$ for four normal K giants. Dearborn (in NATO, 1974) has obtained similar results for a large number of K giants (see also Lambert *et al.*, 1974; Tomkin and Lambert, 1974; Thompson and Johnson, 1974). Some subgiants have nearly "normal" isotopic ratios, $^{12}C/^{13}C = 30$–50 (Dearborn *et al.*, 1975). There is some small amount of sense

TABLE III. Compositions of stars whose surface abundances are thought to reflect nuclear reactions in the stars themselves. The numbers tabulated are logarithmic abundances by number, relative to the Solar System, normalized at Fe, i.e., $\log(N_M/N_{Fe})/(N_M/N_{Fe})_\odot = [M/Fe]$. An Ap star is included for comparison.

Element	Carbon star (mean)[a]	Barium star HD116713[b]	CH star HD198269[c]	Helium star HD168476[d] and HD12448[d]	Hydrogen-deficient carbon star HD182040[e]	R Cr B star RY Sgr[f]	Ap star HD101065[g]
H	0.0	···	+1.56	<−4.6	<−5.3	<−4.5	0.0
He	···	···	···	+0.4	+0.4	+0.4	···
Li	+1.0	···	···	···	···	···	+2.4
C	−1.0	+0.6	+1.92	+0.5	+1.0	+1.5	+0.5
N	+1.2	<+1.2	···	+0.35	···	0.0	···
O	−1.7	···	···	<−0.6	···	−0.6	···
Ne	···	···	···	+0.4	+1.1	···	···
Na	0.0	+0.5	−1.21	···	···	+0.7	−1.3
Mg	0.0	0.0	−0.45	+0.2	···	−0.2	−1.1
Al	0.0	+1.0	···	+0.4	···	+0.1	+0.1
Si	0.0	+0.1	+0.53	−0.35	···	−0.2	−1.3
P	···	···	···	+0.7	···	···	···
S	···	···	···	−0.35	···	+0.1	+0.4
Ar	···	···	···	0.0	···	···	···
K	0.0	0.0	···	···	···	0.0	+0.9
Ca	0.0	+0.1	+0.07	0.0	+0.2	−0.5	−0.7
Sc	···	+0.1	+0.84	+0.3	+0.4	+0.2	+2.1
Ti	0.0	0.0	+0.58	+0.8	+0.1	+0.2	+1.1
V	···	+0.1	+0.4	+0.6	0.0	+0.4	+2.1
Cr	0.0	+0.3	+0.24	−0.4	+0.2	0.0	+0.5
Mn	−0.5	0.0	−0.93	−0.7	−0.5	0.0	+0.9
Fe	0.0	0.0	0.0	0.0	0.0	0.0	0.0
Co	···	+0.1	+0.13	···	+0.6	···	+1.9
Ni	−1.5	+0.1	−0.15	−0.6	−0.1	−0.1	+0.2
Cu	···	+0.3	···		···	···	+1.4
Zn	···	+0.2	+0.12		−0.9	−0.2	−0.8
Ge	···	+0.3	···		···	···	···
Rb	···	+0.7	···		···	···	+3.6
Sr	···	+0.6	···		···	···	+2.0
Y	···	+0.9	+0.48		+0.5	−0.5	+3.9
Zr	+1.5	+0.8	···		+0.3	+0.4	+3.6
Nb	···	+0.3	···		···	···	···
Mo	···	+0.7	···		···	···	+3.3
Ru	···	+0.8	···		···	···	+4.5
Rh	···	···	···		···	···	+4.4
Cd	···	···	···		···	···	+2.3
Sn	···	···	···		···	···	+4.8
Ba	···	+1.2	+2.07		+0.1	−0.5	+2.9
La	···	+0.9	+1.39		+0.2	−0.3	+4.2
Ce	···	+1.0	+1.65		−0.1	+0.2	+3.8
Pr		+0.7	+1.41		···	···	+3.9
Nd		+0.8	···		0.0	···	+4.2
Sm		+0.7	+1.21		+0.8	+0.3	+4.4
Eu		+0.3				0.0	+4.6
Gd		+0.8					+4.7
Tb		···					+4.8
Dy		···					+4.9
Ho		···					+5.0
Er		···					+5.0
Tm		···					+5.7
Yb		+0.4					+4.5
Lu		···					+4.5
W		+1.1					+4.1
Re		···					≥+4.8
Os		···					+3.8
Hg		···					≥+2.9
Tl		···					+3.5
Pb		···					+4.1

[a] From Querci *et al.* (1974), Merchant (1967), Boesgaard (1970a,b).
[b] Danziger (1965a).
[c] Lee (1974).
[d] Hill (1965), modified for modern *f* values.
[e] Warner (1961).
[f] Danziger (1965b).
[g] Wegner and Petford (1974).

to it all; for instance, α Herc, which is probably going up the red giant branch for the first time and so shows only the results of main sequence processing, has a higher $^{12}C/^{13}C$ ratio (30) than α Sco ($^{12}C/^{13}C = 10$), which is ascending the red giant branch for the second time, and so reflects more processing (Wollman, in NATO, 1974).

α Herc and α Sco both appear to have $^{17}O/^{16}O$ enhanced by large factors. The theory of ^{17}O enhancement by means of CNO cycle hydrogen burning out of equilibrium is discussed by Dearborn and Schramm (1974). Various people have looked for correlations of the C:N:O elemental ratios with isotope ratios, but saturation effects in the lines of the more abundant isotopes make it very easy to fool oneself (Faÿ, 1974). Scalo (1974) has interpreted such correlations as implying the existence of two separate kinds of mixing in giants, bringing CNO cycle (Sec. III.C.1) and triple alpha process (Sec. III.C.2) products, respectively, to the surface.

Finally, something must be said about the case of the strange supergiant FG Sge. It is not safe to write down a spectral type for it, since it was a B4 star in 1955, B9 in 1960, F0 in 1969, and F6 (and still cooling at 250 K/yr) in 1972. By the time you read this, it may well be a G star, or even have turned around and headed back toward the territory of the normal nuclei of planetary nebula (it is surrounded by one, which was ejected about 6000 years ago). Even more striking, the abundances of the *s*-process elements like Y, Zr, Ce, and La, which were roughly normal in 1965, started increasing shortly thereafter and are now up to 25 times the solar value, while Fe, Cr, and Ti remain constant and normal. From an observational point of view, FG Sge is evidently turning rapidly into a barium or an S star, but it is not yet possible to tell which. A close look should be kept for the appearance of Li or Tc or high $^{13}C/^{12}C$, which could clarify the situation. The observations are given and discussed by Langer *et al.* (1974). The suggestion that the temperature and chemical behavior of FG Sge could be understood in terms of helium-shell-burning flashes in a highly evolved, low-mass star and the *s* process was first made by Paczyński (1971a), who did not, however, think the situation worth writing a whole paper about (1972, private communication). A detailed theoretical model for the star has been suggested by Sackmann-Christy and Despain (1974) and criticized by Ulrich (1974b), who formulates some more general theoretical considerations and draws attention to the serious difficulty that the cause of the mixing which brings the processed material to the surface in this and other evolved stars is still not well understood (but see Sec. III.E.1 for Iben's work on this). Rose (1974) has pointed out that some of the observed effects could be due to material from the planetary nebula falling back onto the star. It does not seem possible to decide whether FG Sge is a 1–2 $\mathfrak{M}_\odot$ star or a 3–5 $\mathfrak{M}_\odot$ star.

b. Helium stars, R CrB variables, and other hydrogen-deficient stars. Single stars whose surfaces are significantly depleted in hydrogen occur over an exceedingly wide variety of temperatures (05–G), luminosities, and masses. It is somewhat a matter of taste whether they are regarded as a dozen different phenomena or only one. Observers tend to favor the former view and theorists the latter. The depletion ranges from H/He ~ 1 to $<10^{-4}$. The missing hydrogen seems to have been converted to helium; the evidence for this is strong helium lines in the hotter members of the classes and the consistent model atmospheres that can thereby be obtained for the cooler ones (Myerscough, 1968). Abundances can, therefore, often be obtained with respect to He. Carbon is sometimes wildly overabundant, making up about 50% by mass of the atmosphere in some cases. Analyses which do not take into account the effect of carbon on the continuous opacity are, therefore, fairly useless.

Most abundance analyses of these stars have been done by differential curve of growth methods on the assumption that the average metal abundance was that of Pop I. Wallerstein (1973) has suggested that Pop II metallicity may be more appropriate. The classical hot helium stars have been reviewed by Hill (1969) and the cooler hydrogen-deficient carbon stars and R CrB variables by Warner (1967). Table III includes entries for a typical He star (HD 168476, analyzed by Hill, 1965), a hydrogen-deficient carbon star (HD 182040, analyzed by Warner, 1967), and an R CrB variable (RY Sgr, analyzed by Danziger, 1965). Even greater carbon excesses occur in, e.g., XX Cam, for which [C/Fe] = 2 (Orlov and Rodrígues, 1974). In some cool hydrogen-deficient carbon stars, the carbon appears to be pure ^{12}C (Wallerstein, 1973).

The enhancements of Na in the various helium stars are probably real, but the nitrogen abundances are very uncertain. In view of the difficulties of the analyses, Warner (1967) regards all abundances beyond Na in the hydrogen-deficient carbon stars as being consistent with solar values. This may also be the case for the hotter helium stars, when allowance is made for the ups and downs of the f values of iron-peak elements that have occurred in the past few years. Some, but not all, of the stellar determinations used f values found from other stars rather than laboratory values.

It seems clear that these stars show the results of hydrogen and helium burning and, perhaps, associated reactions involving ^{14}N, ^{18}O, ^{22}Ne, and the like (Wallerstein, 1973). The critical question of how they arise from hydrogen-rich progenitors has not been conclusively answered, although the subsequent evolution is reasonably well understood. Neither the mixing of surface hydrogen down to the point where it can be burned nor the loss of the unprocessed envelope is entirely free of difficulties. Helium star evolution tracks have been calculated by Paczyński (1971b) and Dinger (1970, 1971, 1974, and work in progress).

The final entry in Table III is for the exceedingly peculiar Ap star HD 101065 (Wegner and Petford, 1974), also known as Przybylski's star. It has unusually large values of the rare earth and other enhancements characteristic of the classic Ap stars, and some additional high metal abundances of the type shown by some "super metal rich" stars. Although most workers in the field now believe that the Ap peculiarities are to be attributed to atmospheric diffusion (see Sec. II.A), this very extreme star will constitute a critical test of the theory and may eventually be regarded as showing evidence of some nuclear processing as well.

c. Algol-type and other binaries. The non-main-sequence components of evolved close binaries sometimes show great deficiencies of hydrogen and anomalous ratios of C, N, and O. Typical examples are HD 30353 and ν Sgr (Wallerstein, 1968) and the Wolf-Rayet stars that belong to binary systems (Castor and Van Blerkom, 1970; Castor and Nussbaumer, 1971; Paczyński, 1973). The abundance anomalies here can be presumed to reflect the fact that the outer layer of the primaries have been transferred onto the secondary, revealing regions in which hydrogen burning, and perhaps other reactions, have taken place. There do not seem to be any particular difficulties in understanding the surface abundances of these stars as the products of mass transfer, although other aspects (particularly of the Wolf-Rayets) may be the subject of much disagreement. Many of the details of the mass transfer or mass exchange process have been reviewed by Paczyński (1971).

The primaries (i.e., initially more massive, now evolved components) of some other evolved close binaries show a more worrying effect. Many of them have ultraviolet excesses and anomalies in narrow band photometry of the types normally associated with low metal abundance, which are mild cases of larger similar anomalies shown by wide evolved binaries (Koch, 1972). In the wide pairs, these anomalies are to be attributed to nonstellar hot continuum radiation, presumably from hot gas in the system. Hall (1969) formulated a model for the closer, Algol-type binaries in which the color anomalies were similarly due to a disk of hot gas around the secondary star, fed by mass transfer and heated by impact on the star. Such models must be relevant to those systems like SW Cyg in which the uv excess is variable (Hall and Garrison, 1972), but there is evidence that they are not universal. Bond (1972) obtained a spectrum of one of these stars, S Vel, at light minimum, when the A5V secondary was completely concealed behing the K2III primary. He found the spectrum to resemble that of the high velocity red giant δ Lep, which has weak CN bands and is believed to be genuinely metal deficient. There was no evidence of extra uv continuum, beyond what would be expected from the weak-lined K star. The line strengths implied $[Fe/H] = -0.9$, and there was no evidence for enhanced nitrogen from the strength of the NH bands. Mass transfer in this system has not, therefore, penetrated into a region where hydrogen burning by the CNO cycle has taken place. The total mass of the system is now about $2.5\ \mathfrak{M}_\odot$, so that the time between its formation and the onset of evolutionary effects is at most 6×10^9 years (Bond, in NATO, 1974). The metal abundance is thus quite low for the age of the system. Naftilan (1974, 1975) similarly finds evidence that Z is lower in the giant than in the main sequence component of five out of six close binaries whose spectra he studied and that Z is low for its age in the high mass system U Sag.

These systems may just represent a low-Z tail to the metal abundance distribution, corresponding to the high-Z, "super metal rich" tail (Sec. II.C.1.e). Bond's (1972, and in NATO, 1974) suggestion is that the mass transfer, having uncovered the inner parts of one of the stars, is telling us that the metal abundance is not constant throughout stars, the surfaces being enhanced relative to the interiors.

If Bond's suggestion is right, it clearly has profound implications for our entire understanding of the chemical evolution of the Galaxy, stellar evolution, and so forth. A similar suggestion has been made by Joss (1974), who takes the surfaces of stars to be metal-enriched by the infall of comet-like material. Tinsley and Cameron (1974) discuss the implications of this suggestion for galactic evolution. Such surface metal enrichment in the Sun would go part of the way toward solving the notorious solar neutrino problem (Bahcall and Sears, 1972), but an interior metal reduction by a much larger factor than that implied by the S Vel data is required to help significantly (Ulrich, 1974a).

The amount of surface metal enrichment which is likely to be produced by the Joss mechanism in stars with convective envelopes is very small (R. Ulrich, private communication to Fowler, 1975). The data on binaries do not all favor this picture either. Spinrad (in NATO, 1974) has pointed out that the spectrum at minimum light of another Algol binary, AR Lac, has line weaknesses which are most pronounced at short wavelength. This is consistent with an extra source of uv continuum partially filling in the lines. Detailed composition analyses (curve of growth or model atmosphere) of both the primaries and the secondaries of a number of these evolved close binary systems are badly needed.

d. White dwarfs. Because white dwarfs represent the cores of highly evolved stars, we might hope to gain some information from them on the products of the later phases of nuclear burning. Their spectra are, however, discouraging, the very high pressures broadening the lines greatly. In addition, a single star never seems to show evidence of more than two or three elements—H + He; H + Ca; Ca + Mg + Fe; C, H, or He alone; and so forth. More detailed studies confirm this unpromising picture. The DA's (whose spectra show only the hydrogen Balmer lines) have $He/H \leq 0.01$, and no evidence for any heavier elements at all. All other types that have been analyzed show a great dominance of helium over everything else, from $He/C = 100$–1000 in the λ 4760 (C_2 band) white dwarfs (Bues, 1973), to $H/He = 10^{-4}$ and metals/He $= 10^{-9}$ in vMa 2 and Ross 640, whose spectra are dominated by Ca (Wegner, 1972; Hammond, 1974). The DB's (which show only He lines) also have very low metal and hydrogen abundances (Bues, 1970). Apparently diffusion processes in the very strong gravitational fields of these stars have won out over the nuclear processes.

e. Supernovae. A star that supernovas presumably represents the most highly evolved object possible, in which even the outermost layers might be expected to show the results of nuclear reactions. Abundance determinations from supernova spectra would, therefore, by extremely interesting. Only very tentative and preliminary remarks can be made; after the quarter-century of bewilderment that followed Minkowski's (1939) initial detailed description of a supernova spectrum, there is just beginning to be general agreement on the basic character of the spectrum and on the identification of a few features in both Type I and Type II spectra (Searle, 1974). The chief difference between the two types is the presence of hydrogen lines in Type II's, along with HeI, FeII, Ca H and K, Mg, and numerous, unidentified features which are probably blends. Patchett and Branch (1972) have interpreted the spectrum near maximum light of SN 1961v (a II-like type V) as showing lines of H and FeII consistent with a normal Fe/H ratio. Type I spectra, on the other hand, as represented by the scans of Kirshner *et al.* (1973), show no evidence of hydrogen lines at maximum light, although Hα and Hγ appear after about 15 days. Mustel (1974) has concluded that the presence in these spectra of features attributable to He and N, and the absence of features attributable to H and C, reqiure N/H>1 by number. Gordon (1973) believes that H/He ≤ 0.1 in Type I's, and Searle (1974) has suggested that the number of metal ions per free electron is larger in Type I's than in Type II's, which also implies relative hydrogen deficiency and, perhaps, heavy element enhancement. Kirschner *et al.* (1973) interpret their own data as being consistent with solar composition for Type II supernovae and significant He enrichment in Type I's. In addition, Kirshner (1975) has interpreted the spectrum of SN 1972e well after maximum light as showing lines of [FeII], whose strengths would imply [Fe/H] $\sim +1.3$ and the ejection of about 0.1 $\mathfrak{M}_\odot$ of heavy elements from the supernova. Similar SN I spectra are interpreted by Mustel and Chugay (1975) as showing silicon lines with [Si/H] $\sim +2$. Supernova spectra have been reviewed by Oke and Searle (1974).

The analysis of supernova spectra (probably by means of spectrum synthesis) is clearly an important and promising field for the future. The most that can be said at present is that Type I's probably show evidence of abundance anomalies consistent with their highly evolved state. Some complementary information is available from studies of supernova remnants (Sec. II.D).

D. The interstellar medium

The interstellar medium contains several discrete phases with differing density and temperature, some associated with stars and some not. We will discuss here abundance determinations in (1) HII regions surrounding young stars, (2) ionized regions associated with old stars (planetary nebulae), (3) ionized regions associated with deceased stars (supernova remnants), and (4) the general interstellar medium between us and various nearby stars.

1. HII regions

Abundances of the commoner elements can, in principle, be derived from their emission lines in ionized regions. The ratio He/H is the most readily and accurately obtained. Since both are represented by their recombination spectra (from the singly ionized to the neutral state; radio observations tell us that He^{++} does not occur in normal HII regions), the abundance ratio follows directly from the observed line intensity ratio if the recombination coefficients and the kinetic temperature are known. When proper allowance is made for the possible difference in size of the hydrogen and helium Strömgren spheres (see Sec. II.A), results for all regions studied appear to be consistent with He/H$\sim$0.1 (Peimbert and Costero, 1969; Peimbert, 1974).

TABLE IV. A comparison of elemental abundances in the Solar System, in the Orion Nebula (presumably representative of the interstellar medium now) and in the average of several planetary nebulae. The abundances are normalized to hydrogen. Most of the differences are probably not significant.

Element	Solar System[a]	log$N$ Orion Nebula[b]	Planetary Nebulae[c]
H	12.00	12.00	12.00
He	10.9	11.04	11.23
C	8.6	8.37	8.7
N	8.0	7.63	8.1
O	8.8	8.79	8.9
F	4.6	...	4.9
Ne	7.6	7.86	7.9
Na	6.3	...	6.6
S	7.2	7.47	7.9
Cl	5.5	4.94	6.9
A	6.0	5.95	7.0
K	5.5	...	5.7
Ca	6.4	...	6.4

[a] Data as in Table I.
[b] Aller (1972).
[c] Aller and Czyzak (1968).

The other elements present a more difficult problem. They are represented by forbidden (usually) lines arising from only one or a few ionization levels for each element, and there is a great deal of temperature, density, and ionization stratification through the emission nebulae. This means that unobserved ions must be allowed for in a very non-LTE situation, requiring the computation of the degree of ionization from the balance of collisional and radiative processes under circumstances where it is never certain whether or not the value of n_e or T that you derive from one ion is applicable to the next. The difficulties, which are discussed in greater detail by Osterbrock (1970), have not kept people from trying. The Orion Nebula is the object most frequently studied because of its brightness and convenience for northern observers. The values for the Orion Nebula in Table IV are those of Peimbert and Costero (1969), slightly modified by Aller (1972) to allow for improvements in the understanding of the effects of fine structure. Values which differ significantly from these have not been found for other ordinary HII regions. The abundances are essentially solar but with slightly higher Z, [metals/H]$=+0.4$ on the average, which is not unexpected and is consistent with results from young stars. Simpson (1973) and Dopita (1974, 1973) favor [metals/H]$\sim$0.0 in HII regions.

It is not absolutely clear whether the isotope ratios are always terrestrial or not. Clark *et al.* (1974) have observed radio lines of $H^{12}C^{14}N$, $H^{13}C^{14}N$, and $H^{12}C^{14}N$ in the Orion Nebula. On the one hand, they find that the intensity ratio $H^{13}C^{14}N/H^{12}C^{15}N$ (which are weak lines and can be expected to be unsaturated) is consistent with the terrestrial ratio $(^{13}C/^{12}C)/(^{15}N/^{14}N)=273/90$. This would seem to imply either that both C and N have the terrestrial isotope ratios or that a surprising coincidence has occurred. On the other hand, they believe that the lines of $H^{12}C^{14}N$ are also optically thin, and, on that basis, estimate ratios of the numbers of molecules $H^{12}CN/H^{13}CN=6$ and $HC^{14}N/HC^{15}N=27$, both much smaller than the terrestrical ratios. There are two serious problems in calculating ratios of this type. First, the lines may be saturated, and it is not always possible even to tell whether they are, let alone to allow for it. Second, the effective excitation temperature is not necessarily the same for molecules containing different isotopes, especially when photon trapping is important. Wannier *et al.* (1974) believe that they have shown the HCN lines in Orion to be optically thick, removing the evidence for a high $^{13}C/^{12}C$ ratio. In light of these problems, a nonterrestrial isotope ratio should probably not be taken seriously until the same ratio is found for the same source from two different molecules (Solomon, in NATO, 1974). There are, so far, no such cases published. The data for several molecules in several interstellar cloud regions are reviewed by Bertojo *et al.* (1974). None of them meet Solomon's criterion. It therefore seems reasonable to continue to assume that, at least in the solar neighborhood, interstellar clouds and the stars newly formed from them have terrestrial isotope ratios. The variations in $^{13}C/^{12}C$ found in evolved stars are then really to be attributed to nuclear reactions in the stars themselves.

2. Planetary nebulae

The analysis of the composition of planetary nebulae proceeds in the same way as for ordinary H II regions—relatively straightforwardly for the recombination lines of hydrogen and helium, and with difficulty for the other elements. The object most frequently studied is NGC 7027 because of its brightness and unusually rich spectrum, which makes possible abundance determinations for a larger number of elements than in most objects. Most objects studied give rather similar results. The data in Table IV are for a mean of several planetary nebulae studied by Aller and Czyzak (1968) and are meant to be representative of planetaries associated with disk-population stars.

Relative to the Orion Nebula, studied by the same techniques, helium seems overabundant, as do C, N (especially), and O. This is roughly what we would expect from material ejected from a highly evolved star, since it shows evidence of hydrogen burning by the CNO process and, perhaps, some helium burning. If the carbon enhancement is real, the helium burning in the low mass stars which give rise to the planetaries produces more C than O. It is difficult to know what to make of the large discrepancies for Cl, Ar, and S between the planetaries and Orion; and, unfortunately, the solar abundances of these are so poorly determined that they hardly enable us to choose one of the two values as representative of unprocessed material. The good agreement for neon and the rather mild enhancements of C and O in the planetaries would seem to suggest that the differences for Cl, Ar, and S should not be taken seriously.

Not all planetaries have identical compositions (cf. Boeshaar, 1975). There is one (K648) in a globular cluster (M15) where we expect, and find, a generally low metal abundance (Aller and Czyzak, 1973), Ne and O being underabundant by a factor of about 30 relative to disk-population planetaries. Helium, on the other hand, is significantly enhanced, with He/H = 0.18 (O'Dell *et al.*, 1964), and it is probable that N is enhanced (relative to C and O) as well (Peimbert and Torres-Peimbert, 1971). At least one other high galactic latitude planetary nebula seems also to have a generally low metal abundance (Aller and Czyzak, 1973). Shields (1975) finds that [Fe/H] = −1.3, but that the ligher metals have normal abundances, in NGC 7027. The observations of planetary nebulae and their role in galactic evolution has been reviewed by Miller (1974).

3. Supernova remnants

Perhaps the most convincing evidence we could have that heavy element synthesis is occurring today in massive stars would be to see significant overabundances of metals in young supernova remnants. But life is not this simple. The problems in observation and analysis are formidable. First, there are the problems found in H II regions—stratification in temperature, density, and ionization, and the difficulty of deciding what values of the parameters belong to a particular ion. This is made worse by the physical separation of most remnants into small filaments, which undoubtedly differ in density, kinetic and ionization temperatures, and perhaps in composition as well, and whose velocities differ by so much that an integrated spectrum of the whole remnant is likely to have [N II] and Hα smeared into one broad line. Next, most of the available objects have rather low surface brightnesses and large distances, and several hours at the telescope are likely to yield one a very small number of measurable emission lines. Finally, the analysis is made more difficult by the fact that ionization and excitation are due to some combination of collisions and (probably) synchrotron ultraviolet photons. Danziger (in NATO, 1974) has identified 55 lines of He, C, N, O, Ne, Mg, S, Ca, A, and Mg in spectra of high surface brightness SNRs in the Large and Small Magellanic Clouds. The analysis of this data can be exepcted to increase greatly the available information.

Meanwhile, most of the data available is consistent with normal, solar abundances in most supernova remnants (Woltjer, 1972). The apparently high ratio of [N II]/Hα, for instance, is the effect of ionization and excitation by collisions (instead of blackbody photons), rather than of high nitrogen abundance. These normal abundances are not unexpected in the cases of older supernova remnants, like those studied by Parker (1967), including the Cygnus loop and IC 443, which consist mostly of swept up interstellar matter anyway.

Exceptions to the normal abundance pattern occur in the two young remnants which have been reasonably well studied, the Crab Nebula and Cas A. For the Crab Nebula, Woltjer (1958) found He/H = 0.45 from the intensities of the emission lines. Davidson and Tucker (1970) and Davidson (1973) concluded the ratio might be even higher, He/H = 1–2, on the basis of the ionization structure in the nebula. The available data on heavier elements (C, N, O, Ne, S) is consistent with their being present in normal proportion to the sum of H and He (Davidson, 1973).

Cas A has two components, knots moving at about 3000 km/sec and quasistationary flocculi. The former presumably represent material expelled in the supernova explosion, and the latter material from the presupernova envelope. The data of Peimbert and van den Bergh (1971), which have also been discussed by Searle (1971) and Peimbert (1971), seem to indicate that the moving knots have the ratios of H and N to O, Ar, and S down by a factor of 30–100 compared to solar abundances, while the stationary flocculi have N/O enhanced significantly.

The relative normalcy of SNR composition cannot be said to be inconsistent with any particular theory unless we know the total mass of the parent star, the fraction of its mass which was thrown off into the nebula, and whether that fraction is large enough that it should have penetrated down into the region of the star where heavy elements are produced. Even for the Crab Nebula, the total mass and fraction thrown off are probably not known to within a factor of 5–10 (Woltjer, 1972), so no contradiction is possible. Still, it would be nice to have direct evidence that supernovae really do make a range of heavy elements.

4. The general interstellar medium

The predominantly neutral portions of the interstellar medium do not produce detectable emission lines, apart from high quantum number radio recombination lines, which are difficult to interpret quantitatively, the 21 cm line of HI, and various molecular lines. Atomic abundances in these regions must, therefore, be determined from absorption lines, typically in the spectra of OB stars within a few kiloparsecs of the Sun.

The first identification of a line produced by gas between us and a star was made by Hartman (1904). He noted that there was a narrow component in the CaII K line in the spectrum of the spectroscopic binary δOri whose velocity did not change in phase with the orbit. Struve (1928, 1929) demonstrated that the narrow components were truly interstellar rather than circumstellar by discovering the correlation of K line strength with distance to the stars involved. By 1948, Jentzsch and Unsöld (1948) had realized that there was something peculiar about the abundance ratio Ca/Na as revealed by the CaII H and K lines and the NaI D lines. Attempted correlations of optical absorption lines with 21 cm hydrogen observations in the same parts of the sky made it clear (Habing, 1969) that [Ca/Na] was low (that is, either Ca underabundant or Na overabundant) and that [Ti/H] was low, despite the difficulties in allowing for the hydrogen beyond the stars observed optically. A few rocket uv observations of the interstellar Lyα absorption demonstrated that it was the Ca which was greatly depleted (Savage and Jenkins, 1972). A typical modern result for average interstellar abundances determined for a number of stars by optical methods is $[K/H] = -0.3$, $[Na/H] = -0.5$, and $[Ca/H] = -2.5$ (Hobbs, 1974). The amount of Ca depletion varies significantly from place to place. These elements, Ti, and Fe are the only ones with cosmic abundances $N/H \geq 10^{-8}$ whose resonance lines are available to ground-based observers. Interstellar Li has recently been found at a level $Li/H = 3\text{–}6 \times 10^{-10}$ (Traub and Carleton, 1973; Van den Bout and Grupsmith, 1973), which is about a factor of 2 lower than the ratio in young stars.

The Copernicus satellite, by opening up the ultraviolet region down to the Lyman limit, has made possible the observation of the resonance lines of the most abundant ions of about a dozen more elements. Because the Lyman lines can be observed, abundances with respect to hydrogen can be determined in a fairly straightforward fashion although there is some difficulty about the most suitable choice of a curve of growth when only one line of an ion is visible. The data and its analysis have been presented in a series of six papers (York *et al.*, 1973, and references therein; Spitzer and Jenkins, 1975). The over-all picture is one of depletion with respect to hydrogen of most of the elements studied in most of the stars observed, in some cases by factors up to 100, thus confirming the general trend of the optical results (Morton, 1974). Lower resolution ultraviolet observations from the ESRO satellite confirm the general picture of heavy element depletion (Grewing, in NATO, 1974). Only argon is exempted, and that only if the cosmic abundance is the low one of Crawford *et al.* (1975). The depletion of S and Zn is small.

Since the discovery of the calcium/sodium discrepancy, the standard explanation for interstellar abundance anomalies has been "depletion on grains." This has an intrinsic probability, since 2%–3% of the mass of interstellar matter will be in elements heavier than hydrogen, and the ratio of dust mass to gas mass in typical regions of the interstellar medium is about 0.01. (See Aannestad and Purcell for an exceedingly comprehensive review of the properties of interstellar grains.) Unless the grains are practically pure hydrogen and helium (which seems *a priori* improbable) the heavy element abundances cannot, therefore, be unaffected. The depletion-on-grains picture has been rendered very probable by the analysis of Field (1973), who finds a close correlation between the amount of depletion of an element and the highest temperature at which one of its molecules can condense. He envisions materials with condensation temperatures above about 500°K forming small solid cores in nebulae or the outer atmospheres of cool stars. These cores then accrete mantles of more volatile materials in the general interstellar medium. Thus, the most refractory elements are the most depleted, and the most volatile, the least depleted. This bears a certain resemblance to the picture of meteorite formation described in Sec. II.B. Since argon only condenses at about 20°K, its relatively low interstellar abundance would seem to be an argument in favor of low cosmic abundance.

The amount of dust known to exist from optical extinction measurements can account in this way for the observed depletion of Mg, Fe, Si, and the like. It does not, however, fully account for the depletion of C, N, and O (Greenberg, 1974). The most probable hiding places are complex molecules (each of which could have a very low abundance since there are many species possible) or large grains, which produce little optical extinction or reddening. There is evidence that the CNO atoms are really hiding in some such way and not genuinely missing. Ryter *et al.* (1974) have shown that the correlation of x-ray and 21 cm absorption in the directions of several supernova remnants is consistent with a normal ratio of heavy elements to hydrogen. Since the x rays are largely absorbed by photoionization of *K*-shell electrons, they do not discriminate against molecules or grains. The same correlation in the direction of the galactic center could be consistent with either a slightly enhanced CNO abundance there or with the presence of considerable molecular hydrogen. Neither seems intrinsically improbable.

The interstellar ultraviolet absorption lines also provide a direct measure of D/H from their Lyman lines, which are separated by about 0.25 Å. Even though strong chemical fractionation effects occur for deuterium, atomic hydrogen is so much more abundant than molecules along the line of sight to unreddened stars that there does not seem any way for the measured value, $D/H = 1.4 \times 10^{-5}$ (Rogerson and York, 1973) to be severely in error. This is very slightly smaller than the Solar System value from Jupiter, D/H $= 2 \times 10^{-5}$ (Trauger *et al.*, 1973). If the difference could be taken seriously, it would provide some evidence that deuterium is being gradually destroyed rather than created by stars in the Galaxy.

The general conclusion to be drawn, therefore, seems to be that the general interstellar medium has the same composition as the smaller volumes illuminated as HII regions, and that the apparent deficiency of heavy elements is to be attributed to the formation of interstellar grains and molecules. The uv observations have also revealed a rather low average level of ionization of the heavy elements. The small amounts of CIII and NIII observed appear to be inconsistent with the large fluxes of low energy x rays or cosmic rays sometimes suggested as heating sources for HI regions (Meszaros, 1973).

G. Steigman (private communication) has cast serious doubts on much of this picture on the grounds that the uv absorption lines measured by Copernicus are produced in HII regions around the stars and tell us very little about the general interstellar medium.

The interstellar medium also has a significant molecular component. CN and CH have long been known from their optical absorption lines to occur quite widely, as does OH with its 18 cm radio emission line. H_2, though long anticipated, was only discovered in 1970 from rocket ultraviolet observations (Carruthers, 1970). It is now known from Copernicus results to constitute about 10% of the hydrogen along the line of sight to typical stars with B–V excesses of 0.1 to 0.3. But even this underestimates the importance of molecular interstellar gas. The CO surveys by Scoville and Solomon (1974) and Scoville *et al.* (1974) reveal strong emission coming from a 100 parsec radius ring near the galactic center and an annulus 4–6 kparsec from the center. In the central region, if the CNO/H ratio is normal, the amount of mass in H_2 must be much greater than that in atomic hydrogen. In the solar neighborhood $n(H_2) = 1\text{–}2 \times n(H)$ even if CO uses up all the available carbon and oxygen. Thus, averaged over the Galaxy, the mass in interstellar gas must be at least twice what is estimated from HI alone. The positions of these molecular rings are well correlated with the regions of the Galaxy which show large numbers of giant HII regions and enhanced gamma ray flux (Puget and Stecker, 1974). This is not unexpected if massive stars are presently being formed there from dense interstellar clouds.

Finally, the presence of a wide variety of molecules (about 25 to date) in compact ($R \sim 50$ parsec), cool ($T_{\rm kin} \sim 60\text{–}90°$K), dense [$n(H_2) \sim 10^{3-4}$] clouds with total masses of 10^{4-5} solar masses is well known. Understanding the observed composition of these clouds is a problem in chemistry rather than nucleosynthesis, but their existence and properties are clearly relevant to the problem of star formation and will be discussed briefly in that connection in Sec. IV.

E. Cosmic ray source composition(s)

Cosmic rays first revealed themselves as an extra source of ions in the atmosphere which discharged gold-leaf electroscopes (Elster, 1900; Geitel, 1900) and made tracks in cloud chambers (Wilson, 1900). That the source was overhead rather than underfoot was demonstrated by Hess (1911, 1912) and Kolhörster (1913) by means of low-altitude, manned balloon flights. The senior experimenters themselves made the flights, a clear model for the space program. The very great penetrating power of the rays made reasonable the assumption that they were photons of even more than laboratory gamma ray energies. The first evidence that this might not be so came from the observation of a slight latitude dependence of the intensity of cosmic rays (Clay, 1927). Latitude dependence would imply that the incoming particles could be deflected by the Earth's magnetic field and must, therefore, be charged. The effect was not immediately confirmed by other observers, but the charged particle nature of the rays was demonstrated in a different way. Bothe and Kolhörster (1928, 1929) found that several Geiger counters could be simultaneously triggered, even when they were separated by 4 cm of gold. It did not seem possible that gamma rays could scatter a sufficient number of electrons with sufficient penetrating power to account for these observations. The effect of the Earth's magnetic field was again important in revealing the sign of the charge on the particles. Johnson (1935, and references therein), observing from a high-altitude site in Mexico, found that about 10% more cosmic rays came from the west than from the east. The effect, which was quickly confirmed by B. Rossi, L. Alvarez, and others implied that the primary particles must have positive charges at the time they encountered the Earth's field.

The discovery of the positron and the various mesons almost confused the situation beyond repair, opening up a large number of new possibilities for the secondary and primary particles, but the nature and process of formation of showers and secondary particles were clarified in a shower of papers in 1937 (Carlson and Oppenheimer, 1937; Bhabha, 1937; Oppenheimer and Serber, 1937; Heitler, 1937). The dependence of the east–west effect on altitude of observation seemed to imply a high mass-to-charge ratio, and Johnson (1939) first suggested on this basis that the primary particles must be protons.

Finally, an analysis of the altitude dependence of the hard and soft cosmic ray fluxes and the lead-penetrating ability of the high-altitude particles led Schein *et al.* (1941) to the conclusion that the primary cosmic ray particles must be protons, thus establishing the fundamental fact about the composition of cosmic rays as they are now understood.

A cosmic ray component of heavier nuclei was established by Freier *et al.* (1948) from the large amounts of ionization that the particles could produce in photographic emulsions. The originally postulated electrons and gamma rays were finally found in 1962 by Meyer and Vogt (1962) and Kraushaar and Clark (1962), respectively. The electrons make up about 1% of the cosmic ray primaries.

1. The arriving and "source" composition at intermediate energies

In recent years, scintillation detectors, Cerenkov counters, and the like have been flown on high-altitude balloons and satellites by a very large number of groups to measure

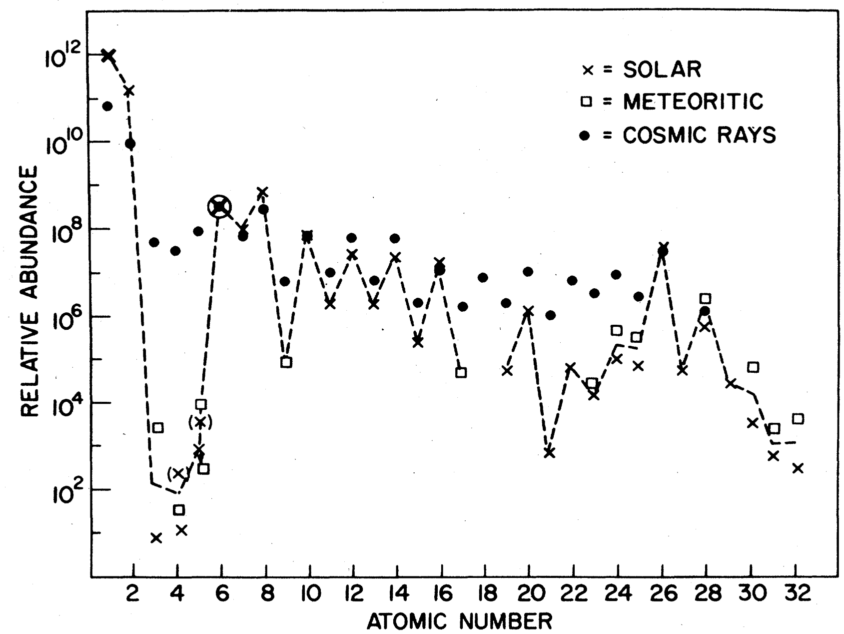

FIG. 6. A comparison of the composition of cosmic rays arriving at the top of the atmosphere with the observed Solar System abundances, normalized at carbon. The data are those of Shapiro and Silberberg (1974), who explain the details of how the arriving composition is determined. The most significant features are the relative depletion of H and He in the cosmic rays, and the smoothing-out of the peaks and valleys of the Solar System distribution. This latter is largely the effect of spallation in the interstellar medium.

the charge composition of cosmic rays at various energies. The largest amount of work has been done roughly in the range 10^7–10^{10} eV/amu, which is high enough to be relatively free from solar modulation effects and low enough that the flux is still large enough to measure with reasonable statistical accuracy. There is sufficient agreement among the various observers that it is possible to build up a reasonably consistent picture of the arriving composition over this energy range. The data shown in Fig. 6 is from the compilation by Shapiro and Silberberg (1974). The main features of the arriving cosmic ray composition have been discussed in Sec. II.A. Many of the differences from the Solar System abundances are due to spallation in the interstellar medium. These effects of the intervening material must be corrected for if we are to discover the composition of the material that left the sources of the cosmic rays (whatever they may be). The problems involved in doing this are also discussed in Sec. II.A. The resulting "source" composition is still subject to some uncertainties, due, among other things, to poorly known cross sections and a certain underdetermination of the problem (since the isotopic ratios are not known, *a priori*, for each element at the source and are generally not measured in the arriving cosmic rays).

The cosmic ray "source" composition shown in Table I and Figs. 7 and 8 is that of Shapiro and Silberberg (1974). It has been arbitrarily normalized to the Solar System composition at carbon. The path length distribution used is the rather complicated one,

$$dN/d\lambda = [1 - \exp(-2.8\lambda^2)] \exp(-0.23\lambda),$$

which seems to be consistent with a wide variety of observations of secondary to primary ratios in the relevant energy range. With this normalization, the cosmic ray sources appear to be deficient in H, He, N, and O with respect to the Solar System and overabundant in nearly all heavier elements, provided that the solar flare values for Ar and S are used rather than the higher interpolated values of Cameron (1973). It should be kept in mind that "source" composition, in this context, means the composition after the acceleration to relativistic energies has already taken place, and that the solar flare particles have already told us that differential acceleration of the elements occurs. Kristiansson (1974) has suggested that enhancements of elements in the "source" composition are correlated with their cross sections for ionization by fast-moving charged particles, and that this tells us something about the effects of acceleration.

In addition, the results for elements heavier than iron are extremely uncertain. The total number of ultraheavy particles collected to date in plastics, ionization chambers, and similar detectors is only a few thousand, the total exposure time available having been roughly doubled by the recent experiment on Skylab (Price and Shirk, 1974). In particular, the "source" composition is not well enough determined to distinguish definitely pure r-process material from some Solar-System-like combination of r- and s-process material. Most of the available data seems to show a larger peak near Pt than near Pb, indicating that the r process is the more important. The mean path length traversed is lower for the very heavy nuclei than for the lighter ones (0.7 g/cm^2 or so, vs 4–6 g/cm^2), implying a travel time of 10^{5-6} yr since acceleration. The relative absence of transuranic nuclei, on the other hand, suggests an elapsed time of 3×10^7 yr or more since the synthesis of the r-process component. No nuclei with $Z \geq 110$ were seen by the Skylab experiment. The picture that comes out of this preliminary Skylab data is then of material, either from the general interstellar medium or somewhat enhanced in r-process material from a nearby supernova, being accelerated some time after the nucleosynthesis took place. Since the travel time is so short, all observed ultraheavy nuclei may well come from one source, so that the conclusions need not be generally applicable. The spectrum from 0.1 to 10 GeV/amu of the nuclei above $Z = 70$ is, however, the same as that of iron in the same energy range, to within the experimental errors, suggesting that the ultraheavies may not be untypical of cosmic rays in general.

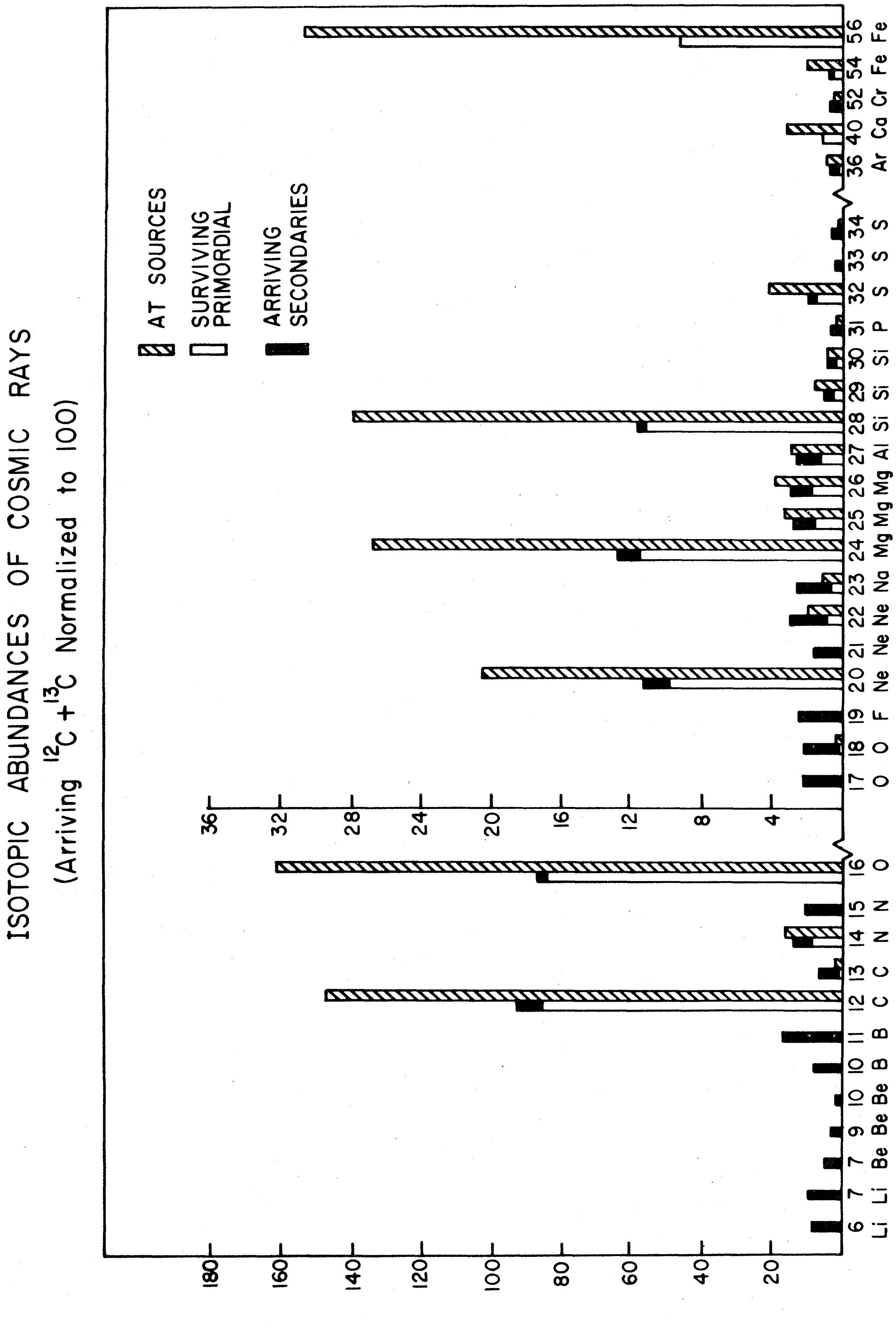

FIG. 7. A comparison of the composition of the cosmic rays from $A = 6$–56 as they arrive at the top of the atmosphere and as they left their sources (after acceleration to relativistic velocities but prior to spallation in the interstellar medium). The data are from Shapiro and Silberberg (1974), who discuss how the corrections for spallation are made. This is also discussed in Secs. II.A and II.E.

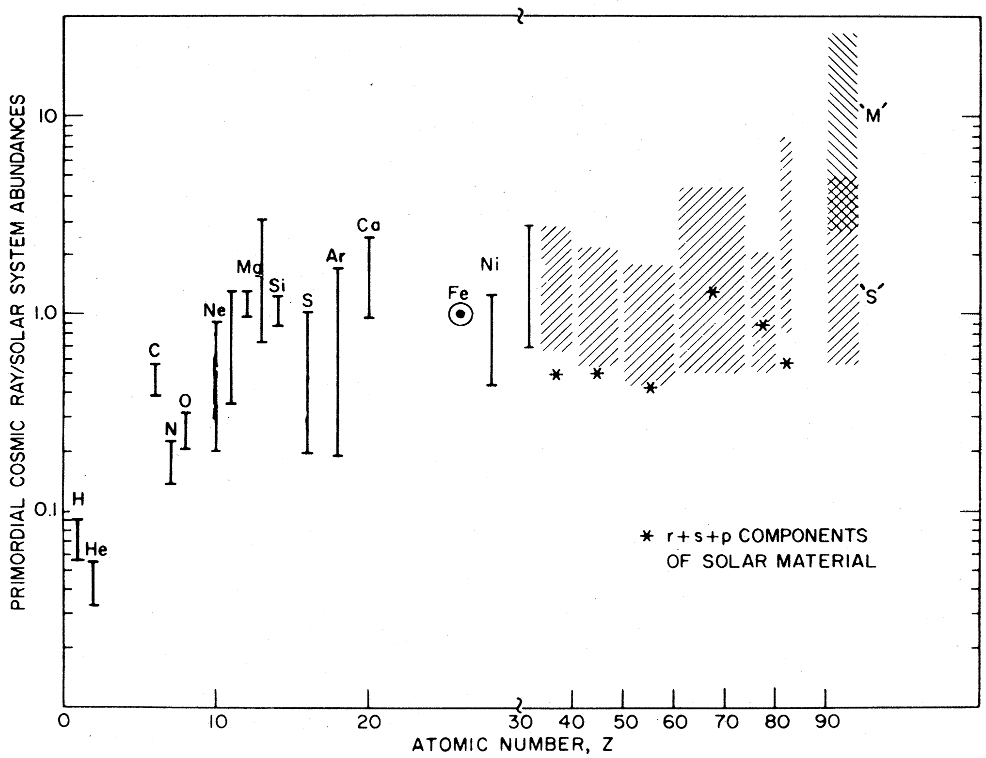

FIG. 8. A comparison of the composition of the cosmic rays as they left their source objects (whatever these may be) with that of Solar System material. The data is that of Shapiro and Silberberg (1974), who explain in detail how the "source" composition is calculated. The large error bars on Ne, S, and Ar reflect the large uncertainties in the Solar System abundances of these elements. The normalization is at Fe. For the heavier elements, the large cross-hatched areas represent primarily uncertainties in the observed cosmic ray abundances. The stars indicate what would be observed if ordinary Solar System material had been propagated through the interstellar medium. There is some evidence for extra enhancement of heavy nuclides (especially those produced in the *r* process, q.v.) over and above that of Fe. The hatched regions 'M' and 'S' in the uranium region represent the meteoritic and solar abundances of Th, respectively.

2. Isotopic composition

The resolution in charge and mass of cosmic ray observations has only recently improved to the point where the measurement of isotopic ratios, apart from D/H and $^3He/^4He$, is possible. The observed ratios (Webber *et al.*, 1973a,b) are quite different from the terrestrial ones, the various isotopes of a given element tending, generally, to be more equally represented. As in the chemical composition, however, we must allow for the effects of spallation in the interstellar medium. Tsao *et al.* (1973) have adopted the approach of assuming the "source" chemical composition and terrestrial isotopic ratios, and then calculating the expected arriving isotopic composition after passage through an exponential path length distribution. The predictions are made for slightly higher particle energies than the observations were made at, but this should not be a serious problem (Shapiro and Silberberg, 1974). For the most part, there is good accord between the calculations and the observations. The one surprise is the presence of a fair amount of ^{58}Fe ($^{58}Fe/^{56}Fe \sim 1.0$), which is not common terrestrially or in the products of any of the likely nucleosynthetic processes (Truran, 1972), nor is it a promising spallation product, since there are no sufficiently abundant heavier nuclei. The resolution in mass of the measurements (Webber *et al.*, 1973a,b) is poor enough that the ^{58}Fe excess is not firmly established and could be due to spillover from ^{56}Fe. Further observations with better statistics and better resolution can be expected to provide additional information on either the isotopic composition of the cosmic ray sources or on errors in our spallation cross sections.

Several radioactive isotopes are of particular interest because they provide measures of either the time since the cosmic ray nuclei were synthesized or the time since they were accelerated. The transuranic elements belong to the former class, and the evidence from their absence for a relatively long time between synthesis and arrival ($\gtrsim 3 \times 10^7$ yr) has already been mentioned. The ratio of Th and U is a similar clock. ^{10}Be, on the other hand, is only formed in transit and decays continuously. Its ratio to the other Li, Be, B isotopes, therefore, provides a measurement of the time since acceleration which is independent of the density of the region in which the cosmic rays are confined. Webber *et al.* (1973a) found a few ^{10}Be nuclei and concluded from them that the time since acceleration was about 3×10^6 yr. (This implies confinement in a relatively high density region, such as the disk of the Galaxy.)

Finally, there are a few nuclei which can be expected to be produced by explosive nucleosynthesis and are unstable primarily to electron capture or have radioactive daughters primarily unstable to electron capture. This means they will decay rapidly from synthesis to acceleration and slowly or not at all after acceleration. ^{53}Fe, ^{54}Mn, ^{55}Co, and $^{56,57,59}Ni$ are typical examples. The observations do not yet have sufficient resolution in nuclear mass to use these clocks unambiguously, but the total Ni/Fe ratio suggests that either acceleration did not occur very soon after synthesis, or that the total travel time was longer than that implied by the ^{10}Be data (Reeves, in NATO, 1974).

3. The higher energy cosmic ray composition

We have already noted in Sec. II.A that the ratio of secondary to primary nuclei declines at energies above 10–20 GeV/amu and that there are associated changes in C/O, Fe/(C + O), etc. as the various primaries are eroded away with differing efficiency by spallation. The comprehensive data on these ratios by Juliusson (1974) shows evidence of several effects. First, the spallation cross sections are simply lower at higher energies. Secondly, the average path length traversed seems to decline with energy, so that many ratios approach their "source" composition values. Finally, C/O falls to below, and Fe/(C + 0) rises to above, its calculated source value. It is possible that this reflects an error in the calculations, since the high energy value of C/O is quite close to the "cosmic" value. If the "source" value of C is increased to reflect the high energy C/O ratio, then the high energy Fe/(C + 0) ratio is also equal to the modified "source" value. [Fe/C] is still significantly greater than unity in this revised cosmic ray source composition. The only difficulty with this picture is that C/O and Fe/(C + O) do not seem to be leveling off at the energies (50–100 GeV/amu) observed by Juliusson (1974), unlike most of the other ratios measured.

4. The low energy cosmic ray composition

The experiment by the Goddard–University of New Hampshire group carried aboard Pioneer 10 showed clear evidence for a new, low energy component in the cosmic rays, whose origin, on the basis of its intensity gradient with distance from the Sun, is apparently extrasolar (McDonald *et al.*, 1974). Below about 30 MeV/amu, the ratios of O and N to C gradually rise, reaching about 20 times their higher energy values at 8 MeV/amu. There is also a turnup in the spectra of H and He at very low energies with He/H being enhanced by a factor of about 5 below 100 MeV/amu (see e.g., Shapiro and Silberberg, 1974). The extra He is pure ^{4}He (Simpson, in NATO, 1974). There is no evidence for corresponding anomalies in other nuclei (e.g., Mg, Si, Fe, B, Ne) whether primary or secondary, in this energy range. There is presently no data available whlch bears on the question of how far this component has travelled or for how long. Explanations of it have focused on either the electronic or the nuclear properties of the elements involved (He, N, O), Fisk *et al.* (1974) remarking on the high ionization potential which might allow them to leak preferentially into the solar cavity and then be accelerated, and Hoyle and Clayton (1974) suggesting that they are the sort of thing likely to be made preferentially in novae. Under these circumstances, nothing intelligent can be said about whether or not the composition of this component of the cosmic rays resembles that of its sources.

F. Other galaxies and QSOs

Abundances in external galaxies are normally found either from the integrated spectra of old stars (for Elliptical, S0, Sa, and Sb galaxies) or from the emission lines of HII regions (for Sc, Sd, and Irregular galaxies). An important step still to be taken is the determination of anything but helium abundance for the same regions of the same galaxies by both methods. Extragalactic composition measurements are therefore subject to additional difficulties over and above those normally inherent in interpreting either absorption or emission spectra of single objects, caused by the need to deal with light coming from groups of either stellar atmospheres or nebulae which are exceedingly inhomogeneous in T, n_e, and, perhaps, composition. In addition, we will be looking for rather subtle effects. No galaxies are known with abundance anomalies on anything like the scale of, for instance, the Ap or R CrB stars. $-0.6 \leq [\mathrm{Fe/H}] \leq +0.6$ seems to cover the entire range of observations of average metal abundance except in dwarf galaxies, and $-0.9 \leq [\mathrm{N/O}] \leq +0.9$ the range of differential effects (Searle, 1973), while no statistically significant variations in He/H may have been found at all.

1. Helium

We have seen that the helium abundance in our own Galaxy is quite uniform at $Y \sim 0.3$, higher values being found only in evolved objects (R CrB stars, planetary nebulae, and the like) and only slightly (if at all) lower values in unevolved Pop II objects. The majority of extragalactic determinations are also consistent with this value. Typical examples are the work of Aller (1972) on 30 Doradus (in the Large Magellanic Cloud) and NGC 604 (in M33), of Peimbert and Spinrad (1970a,b) on NGC 4449, 5461, 5471, 6822, and 7679 (some of which are outside the Local Group), of Searle and Sargent (1972) on the dwarf blue compact galaxies I Zw 18 and II Zw 40, of Shields (1974) on the Seyfert 3C 120, and of Williams (1971) on QSOs, all of whom find He/H ~ 0.1 from ratios of emission lines. Iben and Tuggle (1974) have also found $Y = 0.29 \pm 0.06$ for Andromeda and the Large and Small Magellanic Clouds from an analysis of the period–luminosity relation for Cepheids. It is not improbable that this represents the correct value for all unevolved objects.

There is, however, some evidence for either low or high helium in a few places. For instance Peimbert and Spinrad (1970c) found He/H = 0.13 in the active galaxy M82, and Martin (1974) found values as high as 0.14 for some Seyfert nuclei. There is no particular reason to distrust these emission line results, since the problems discussed in Sec. II.A tend to make one underestimate the amount of helium present. The regions involved are also undoubtedly ones where lots of nuclear reactions and explosions have taken place and some *in situ* production of He would not be out of place. The value He/H = 0.13 found by Peimbert and Spinrad (1970a) for NGC 604 is higher than that of Aller (1972), but the difference may not be significant.

The helium abundance may be low in some galaxies where the metal abundance is also low. Peimbert and Spinrad (1970a) found He/H ~ 0.08 in the Small Magellanic Cloud, and Peimbert and Peimbert (1974) confirm the low value, also from emission line work. They also find a low value (He/H = 0.085) for 30 Doradus in the LMC, while a rather similar analysis by Dufour (1973) gives He/H = 0.1 for the same object. Danziger (in NATO, 1974, and 1974) may have resolved the discrepancy for 30 Dor by showing that the derived He/H ratio varies with position in the nebula and is correlated with the reddening at each point in the sense that would be expected (low He with high reddening) if the He Strömgren sphere were in fact smaller than the H one as discussed in Sec. II.A. Discordant values found earlier may, then, be the result of taking spectra at different places in the nebula, and the actual He abundance be normal. Hartwick and McClure (1974) argue for a low helium abundance in the Draco dwarf spheroidal galaxy on quite different grounds, including the morphology of the horizontal branch, the high mean period of the RR Lyrae variables, and the high mass of the giant stars (as determined from spectroscopic values of temperature and electron pressure in the galaxy). Their data can also be explained if the stars are younger than globular cluster giants or have a different opacity law. Searle (quoted by Peimbert, 1975) finds that Y may be as low as 0.185 in II Zw 40.

As in the case of our own Galaxy, it does not seem possible to exclude either the null hypothesis that He/H = 0.1 ($Y = 0.28$) everywhere, or the view (favored by, e.g., Peimbert and Peimbert, 1974) that the helium abundance is systematically lower in galaxies in which the low metal abundance indicates that a minimum of nuclear processing has taken place, and, perhaps, systematically higher in very active regions.

2. Mean metal abundance

Again as in the case of our own Galaxy, when we come to the average abundance of elements heavier than hydrogen and helium, there is no doubt that real variations occur between galaxies and between places in a single galaxy. Burbidge and Burbidge (1970) point out that we can arrange objects in three groups in order of increasing metal abundance: (1) intergalactic globular clusters, low mass extended galaxies (like Draco), and the extreme Pop II of our own Galaxy, all with $[\mathrm{Fe/H}] \lesssim -1.0$; (2) regular dwarf ellipticals (like NGC 205 and M32), the outer regions of giant ellipticals, the outer disk and globular clusters of M31, and the disk and central globular clusters of our own Galaxy, with $[\mathrm{Fe/H}] = -0.5$–$+0.4$; and (3) the nucleus of M31, the nuclei of giant elliptical galaxies, and probably many S0 and Sa and some Sb galaxies, with $[\mathrm{Fe/H}] \gtrsim +0.4$.

The lowest well documented average metal abundance for a galaxy is probably that of the Draco dwarf spheroidal, for which Hartwick and McClure (1974) find $[\mathrm{Fe/H}] = -2.8$ from intermediate band photometry of several giants. A number of other galaxies also have low average metal abundances. Typical examples include the small irregular galaxy NGC 6822 in the Local Group, for which Peimbert and Spinrad (1970b) find $[\mathrm{N/H}] = -0.78$ and $[\mathrm{O/H}] = -0.23$; the Sculptor dwarf spheroidal which has a metal abundance about like that of the globular cluster M3 (Hodge, 1965); NGC 185 (a dwarf elliptical companion to M31) which has $[\mathrm{O/H}] \sim -1.0$ on the basis of one planetary nebula (Jenner *et al.*, 1974); and the Magellanic Clouds, for which Dufour (1973, 1975) finds that N is depleted by a factor of about 11 and O and Ne by 2–3 in the SMC and somewhat less in the LMC. The general trend $Z(\mathrm{SMC}) < Z(\mathrm{LMC}) < Z(\mathrm{galaxy})$ is also indicated by the statistics of red and blue supergiants (Bertelli and Chiosi, 1974), the properties of pulsating variables (Robertson, 1973), and the rather low dust to gas ratio in the SMC (Hodge, 1974). McClure and van den Bergh (1968) have remarked that dwarf ellipticals systematically have weaker CN absorption (and, presumably, lower metal abundance) than giant elliptical galaxies. The compact blue dwarf galaxies I Zw 18 and II Zw 40 studied by Searle and Sargent (1972) also have significant metal depletion, with $[\mathrm{Ne/H}]$ and $[\mathrm{O/H}] \sim -1.0$ in the former and ~ -0.5 in the latter. These low Z galaxies are all characterized by much lower total masses than that of the Galaxy, independent of morphological type. Notice also that, where several elements have been measured, N is always the most deficient. This is consistent with the trend $[\mathrm{N/Fe}] \simeq [\mathrm{Fe/H}]$ in galactic stars advocated by Pagel (1973a).

There do not seem to be galaxies whose overall metal abundance is larger than that in the solar neighborhood [although the globular clusters in Andromeda have a significantly higher Z than those in the Galaxy, according to van den Bergh (1969)], but there are undoubtedly metal-rich regions—the nuclei of the galaxies whose abundance gradients are discussed in the next section. Metal abundances in QSOs seem to be essentially normal (Osterbrock, 1969).

3. Abundance gradients in galaxies

Systematic variations in emission line ratios with position in external galaxies have been known for more than 30 years (Aller, 1942, on [OIII]/Hβ; Burbidge and Burbidge, 1965, on [NII]/Hα), but were long attributed to variations in the temperatures of the exciting stars, ionization, or excitation temperature. Data indicating that the centers of galaxies are systematically redder than the outer regions have also been around for some time (Tifft, 1963, and references therein). But it was not until 1968 that Peimbert (1968) more or less unequivocally demonstrated that no consistent picture could be made out of the line intensity variations in the Sc galaxies M51 and M81 just by varying ionization or excitation conditions, and that abundance variations must play a role. He concluded that N/H must be 2 to 6 times higher in the center of M51 than in its outlying regions and that $[\mathrm{N/O}] \sim +0.6$ in the nucleus.

In the ensuing six years, a great deal of evidence has accumulated to indicate that both spiral and elliptical galaxies have gradients of metal abundance, Z always being higher in the nucleus. M31 has been the most thoroughly studied. Measurements of emission line intensities enabled Rubin and Ford (1971) to conclude that N/H varies significantly in the gas over the inner 500 parsecs, while McClure and van den Bergh (1968) found similarly for the stars (from CN bands) that the semistellar nucleus has the highest metal abundance, the nuclear bulge next, and the inner halo and outer disk roughly solar abundance. The inner halo and outer disk neighborhood of M31 are like the solar neighborhood in having very few stars with extreme metal deficiencies (van den Bergh, 1969); and the companion galaxy, NGC 205, resembles our companions (the Magellanic Clouds) in having a significantly lower metal abundance than the "parent" galaxy. M31 differs from our galaxy in having no correlation of metallicity with position for its globular clusters. Scanner observations by Spinrad *et al.* (1971) showed radial gradients in CN, the lines of NaI and color (especially U–B) which they interpreted as reflecting high metal abundance at the center, which drops to about the solar value 200–500 parsecs from the center, and somewhat lower in the halo (but not as low as for Pop II objects in our Galaxy). Their results for the giant elliptical galaxy

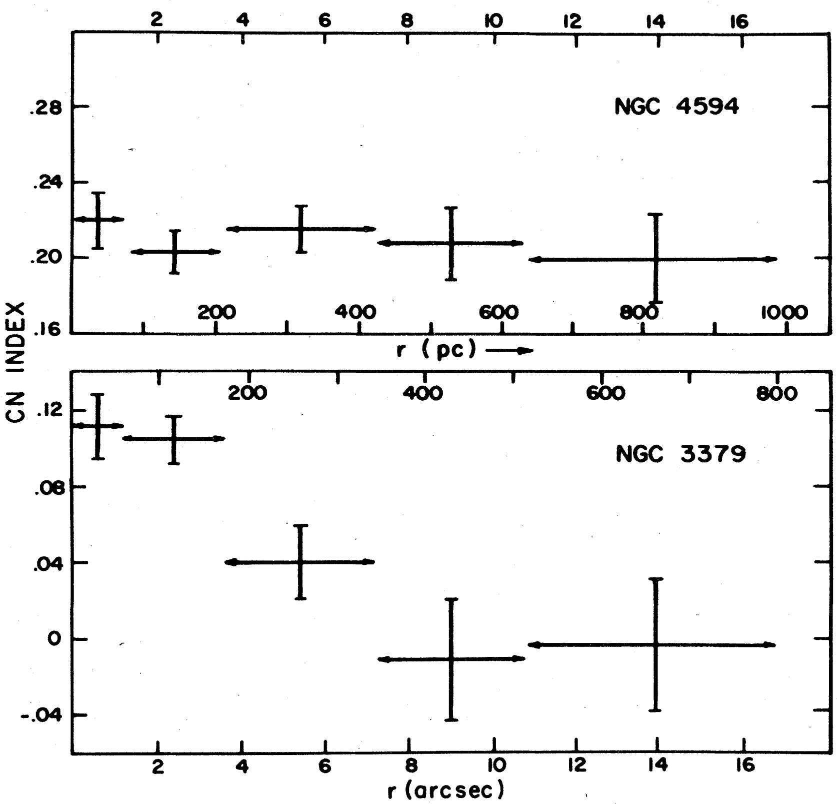

FIG. 9. Gradients of CN line strength (and, therefore, presumably in total metal abundance) with position in two galaxies, NGC 4594 (a flattened Sa spiral) and NGC 3379 (a spheroidal E0). The gradient is much stronger for the spheroidal type. Other galaxies can display stronger gradients perpendicular to their planes than in them, also suggesting that the gradient is primarily due to a halo or spheroidal population. The data is that of Spinrad *et al.* (1972)

NGC 4471 were similar to those for M31. A recent model by Joly (1974) suggests that more than half the luminosity of the nucleus of M31 comes from super metal rich stars ([Fe/H] $\gtrsim$ 0.6). She also concludes that the nucleus probably has a dwarf-enriched main sequence and must have some young B stars. Her model for the nucleus of NGC 3031 (Sb) also requires both SMR and hot stars (Joly, 1974b). Searle (1971; see also Shields, 1974a) has shown that the existence of composition gradients across disks is a general phenomenon in late type spirals.

A number of studies of CN strength and other metal abundance indicators have also demonstrated the existence of Z gradients in elliptical and tightly wound spiral (Sa and Sb) galaxies. McClure and van den Bergh (1968) concluded that several giant ellipticals had super metal rich cores and metal-poor halos, the average Z seeming to be higher (given the same total luminosity) for galaxies in clusters. McClure (1969) remarked that the SMR core composition gave way to a roughly solar disk composition immediately outside the core for a number of spirals and ellipticals. Welch and Forrester (1972), using indicators of CN, G band, MgI, MgH, and NaI line strengths, suggested that the cores of two ellipticals (NGC 3115 and 4472; the latter also studied by Spinrad *et al.*, 1971) were enriched in both metals and the proportion of late-type dwarfs.

Spinrad *et al.* (1972) have found variations between galaxies in the amount of gradient which are correlated with other properties. Spherical galaxies (which are also the ones with the most rapid radial drop in surface brightness) show the steepest gradient in metal abundance, and highly flattened systems (for which surface brightness drops less rapidly with radius) little or no gradient. Figure 9 shows the CN gradients in two typical galaxies—the flattened Sa, NGC 4594 (the "Sombrero" galaxy) and the spheroidal EO, NGC 3379.

The SO galaxy NGC 3115 presents an interesting intermediate case, having little or no abundance gradient along its major axis, but a strong one along its minor axis (that is, perpendicular to its galactic plane). The systematics therefore suggest that the Z variations are associated with the nuclear and halo populations of these galaxies, rather than their disk populations.

Following the pioneering work of Peimbert (1968) on M51, other loosely wound spiral (Sc, Sd) and low-dust irregular (IrI) galaxies have also been found to show radial metal abundance gradients. M101, for instance, has [N/O] $\geq$ 0.5 and [O/H] $>$ 0.0 in its core and [N/O] $= -0.5$ and [O/H] $= -0.4$ at the position of the HII region NGC 5471, 13.4 kiloparsecs from its center. The analysis has been extended to larger radii by Smith (in NATO, 1974), who finds that [O/H] continues to drop to -0.6 at 23 kiloparsecs from the center, but that [N/O] is relatively flat in the outlying regions.

It would be surprising if our Galaxy did not display similar gradients. The correlation of globular cluster metal abundances with position is well known (van den Bergh, 1969). Evidence within the galactic plane is harder to come by, since we do not have an unobstructed view to the galactic center. The evidence for higher metal abundance toward the center includes the slightly higher ratio of x-ray absorption to HI there (Ryter *et al.*, 1974); the high velocities of the "genuine" super metal rich stars like 31 Aql (Sec. II.C above); the strong metal lines of nonvariable late M giants in the nuclear bulge (McClure and van den Bergh, 1968); gradients in the properties of pulsating variables (van den Bergh, 1968); and some intermediate band photometric criteria (papers by Grenon and by Janes and McClure in Cayrel de Strobel and Deplace, 1973). On the other hand, van den Bergh and Herbst (1974) find that there are very few SMR stars in the nuclear bulge. In view of the inaccessibility of the regions closer to the center than about 600 parsecs and the steepness of the gradients in many external galaxies, the data may, in fact, all be consistent. As in the case of single stars, galactic abundances and gradients based on CN line strengths alone must be regarded with some caution.

Helium abundance does not seem to vary systematically across galaxies (Peimbert and Spinrad, 1970b; Smith, in NATO, 1974), although Ford and Rubin (1969) have reported some contrary evidence from the ratio HeI $\lambda5876/\mathrm{H}\,\alpha$ in M31.

III. NUCLEAR PROCESSES, THEIR SITES AND PRODUCTS

The total problem of accounting for the synthesis and distribution of all nuclides in all types of objects is clearly an exceedingly complex one and well beyond the reach of present day astronomy and physics. Considerable progress has, however, been made in three areas of the total problem. The first of these is accounting for the general trends of metal abundance with time in our own Galaxy, with position in our own and other galaxies, and with other average properties of galaxies. This belongs to the realm of "galactic evolution" and will be discussed in Sec. IV. Second, we have gained some understanding of the origins of the abundance anomalies seen in some of the peculiar evolved stars mentioned in Sec. II.C. These stars tend to show evidence of mixing to the surface of the products of the processes which are discussed below as hydrostatic hydrogen burning by the CNO cycle, hydrostatic helium burning, and the *s* process. Finally, theories of nucleosynthesis have had considerable success in accounting for the general trends of the abundance distribution in the Solar System and many of its details. The main thrust of the discussion in the sections which follow will be a comparison of the products of various synthesis processes with the observed Solar System abundances.

Early work on nucleosynthesis tended to try to produce the entire range of Solar System abundances from a single equilibrium or quasiequilibrium process under a single set of conditions. None of these attempts was very successful. The alternative approach of considering a number of different kinds of processes occuring at different times and/or in different kinds of objects has proved much more fruitful. The variety of necessary processes was first delineated, in short form, by Hoyle *et al.* (1956) and by Cameron (1957). Burbidge *et al.* (1957, hereinafter called B^2FH) calculated the various processes in greater detail and this is the fundamental basis on which virtually all modern work on nucleosynthesis in stars is built. B^2FH identified eight necessary processes, which they called hydrogen burning, helium burning, and the α, e, s, r, p, and x processes. The first two of these are self-explanatory. The others are: α—capture of alpha particles by ^{16}O and ^{20}Ne to build $A = 4n$ nuclei up to ^{48}Ti; e—equilibrium among nuclei, free protons, neutrons, and alphas which makes the iron peak nuclides; s—the slow (i.e., slow compared to beta decay rates) capture of neutrons by iron peak and other seed nuclei to make the less abundant isotopes in the $A = 23$–46 range and many of those heavier than the iron peak; r—rapid (compared to beta decay rates) capture of neutrons on seed nuclei, making a large number of the isotopes in the range $A = 70$–209 as well as uranium and thorium, and, perhaps, a few lighter things (because of the rapid time scale, nuclides are formed that could not be made by the s process); p—capture of protons or photoexpulsion of neutrons to make the rare, proton-rich isotopes; x—the unknown process which makes the fragile light nuclides D, Li, Be and B. Cameron (1957) identified many of the same processes: hydrogen and helium burning "in the orderly evolution of stellar interiors"; statistical equilibrium in both supernovae and presupernovae; neutron capture on a slow time scale and on a fast time scale; proton capture and photonuclear reactions (on both slow and fast time scales); and a nonstellar process to make D, Li, Be, and B. In addition, he saw a need for explosive burning of hydrogen and helium in supernovae and for heavy ion reactions under hydrostatic conditions. In the discussion that follows, we will generally use the nomenclature of B^2FH, but will find it necessary to include Cameron's additional processes, and explosive heavy-ion reactions as well.

A. Cosmological nucleosynthesis

The earliest modern speculations on the origins of the elements appear to have been of a cosmological type, Clarke (1889) writing on a possible formation of elements from a primordial "protyl" and Vernon (1890) on formation from a "primordial atom." The fundamental fact that made it possible to put these speculations in quantitative form was not discovered for another 25 years. This is that the abundances of the elements are tied to their nuclear rather than their chemical properties (Oddo, 1914; Harkins, 1917). The ensuing rather complicated history has been sorted out with some care by Kuchowicz (1967). In retrospect, we can list the important points established in roughly chronological order: (1) the need for high temperatures if helium and hydrogen were to be brought into equilibrium in their present abundance ratio (Tolman, 1922); (2) the importance of barrier penetration in fusion as well as fission reactions (Atkinson and Houtermans, 1929), which keeps the requisite temperatures from being completely ridiculous, and also suggests the possibility of cyclic or catalytic synthesis of helium (Atkinson, 1931); (3) the impossibility of accounting for all abundance ratios in terms of equilibrium at a single temperature and density (Urey and Bradley, 1931; Pokrowski, 1931), which requires that light and heavy elements be produced under different circumstances; (4) the need to include the effects of "freezing out" of nuclear processes as matter gradually cools from the synthesis temperature (Farkas and Harteck, 1931); (5) the possibility of successive neutron captures followed by beta decays as an alternative to charged particle reactions (Walke, 1934; Gamow, 1935); (6) the suitability of the early universe as a site for such captures simultaneously with cooling (Gamow, 1946); (7) the strong resemblance between a smoothed version of the observed abundances and what would be produced in an expanding universe consisting initially of an "ylem" of neutrons (Alpher *et al.*, 1948; Gamow, 1949), provided that the product of the expansion time scale and the neutron density was about 1.3×10^{-6} gram sec cm^{-3}; (8) the possibility of starting with radiation, protons, neutrons, electrons, and positrons in equilibrium rather than with pure neutrons and getting roughly the observed He/H ratio (He/H $= \frac{1}{6}$) largely independent of initial temperature, provided only that it is very high (Hayashi, 1950); (9) the hopelessness of the whole endeavor in view of the absence of stable nuclides with masses 5 and 8 (Fermi and Turkevitch, 1949) [The problem is that, by the time the universe is cool enough for deuterons not to be destroyed by thermal photons (a precondition for the beginning of nucleosynthesis), it is also sufficiently low in density that the rate of either three-body reactions or the formation of tritium (which might bridge the gap by ^{4}He + T $\rightarrow$ ^{7}Li) is much too small to carry the observed 1% of matter across the gap.]; and, (10) the impossibility of going all the way with the equilibrium concept and starting with a universe symmetric in matter and antimatter, because all the matter annihilates (Alpher *et al.*, 1953). Other work from the same period which, in retroscpect, does not seem to have been part of the main stream includes (1) the suggestion of synthesis at densities of 10^{11-12} g/cm^3 but at low temperature (van Albada, 1946); (2) the discussions of synthesis under equilibrium conditions at a range of densities established by gravitational forces (Klein, 1946; Beskow and Treffenberg, 1947); (3) the calculations of heavy-element production by fission of large neutron condensations (Mayer and Teller, 1949); and (4) the synthesis of deuterium in an explosion of a proton gas (Murgai, 1952). Hoyle and Taylor (1964) did the first modern calculation of cosmological nucleosynthesis, based on the "universality" of He/H ~ 0.1. We defer to Sec. III.C the history of element synthesis in stars.

Very little further progress on cosmological synthesis was made until after the discovery of the isotropic, microwave background radiation (Penzias and Wilson, 1965), which bore a strong resemblance to a sea of 3°K blackbody photons, such as should remain from a hot, dense, early phase of the universe (Dicke *et al.*, 1965). This prompted Peebles (1966a,b) to re-examine the problem, in light of the newly known temperature, which can be uniquely extrapolated back to any time and, therefore, density in the past, given a long list of assumptions, to give the temperature and density at which synthesis occurred. He concluded that the Big Bang is, indeed, good for making all the elements up to helium, including small amounts of ^{2}H and ^{3}He. Wagoner *et al.* (1967) covered the same ground more thoroughly, finding that a small amount of ^{7}Li was also produced, and that helium production could be reduced to arbitrarily low levels if the universe contains about as many neutrinos (or antineutrinos, but not both) as it does photons. They also showed that heavy elements could be made in high density universes or regions.

The most recent calculations are those of Wagoner (1973, 1974, and in NATO, 1974). He starts with a standard model with the following assumptions: (1) The universe was initially in thermal equilibrium at some temperature above 10^{11}°K (this does not exclude the Hagedorn equation of state, for which the limiting temperature is about 10^{12}°K), so that the initial ratio of protons to neutrons is set by the equilibrium. (2) The baryon number is greater than zero (i.e., the universe is not matter–antimatter symmetric). (3) The lepton number is not large enough (positive or negative) to make the neutrinos degenerate. (4) There are no unknown particles (quarks, superbaryons, etc.). (5) General relativity is the right theory of gravity. (6) The universe (or the region of it under consideration) is homogeneous and isotropic. (7) The reaction rates (for which, see Fowler *et al.*, 1975) involving nuclides with $A \leq 7$, which have all been measured in the laboratory to factors of 2 or better in the relevant energy range, are known. Under these circumstances, synthesis occurs during a fairly brief time as the temperature of the universe reaches 10^9°K (so that deuterons stay bound). The mass of the causally connected universe (the mass inside $R = ct$, where t is the time since the expansion started) at this time is about $10M_\odot$, so that synthesis takes place in unconnected regions which are very much smaller than the masses of galaxies. The one remaining free parameter in the model is the entropy per baryon, or the ratio of the number of photons to the number of baryons. This ratio determines the density at which the critical temperature is reached and does not change with time. Since we know the present photon density fairly precisely from the temperature of the background radiation (2.7°K; Peebles, 1971), the free parameter is, in effect, the present density of matter in the universe. The present density of the universe (if GR with cosmological constant $\Lambda = 0$ is the right theory of gravity) is surely in the range $0.01\rho_c \leq \rho \leq 6\rho_c$, where $\rho_c = 3\mathrm{H}^2/8\pi\mathrm{G}$ is the density required to close the universe. It has the value 6×10^{-30} $(\mathrm{H}/55)^2$ g cm^{-3}, where H is Hubble's constant in units of km sec$^{-1}\cdot$megaparsec^{-1}. The lower limit is set by the mass observable in the form of galaxies (Shapiro, 1970), and the upper limit by the fact that we do not see large amounts of deceleration of the expansion of the universe as we look back to large redshifts (Sandage, 1972; see Gunn and Oke, 1975, for a still tighter upper limit).

The amount of helium made in the standard model is then given by $Y = 0.229 + 0.094\ (\rho/\rho_c)$ for densities greater than $0.02\rho_c$. Since our discussion of helium reached the conclusion that, where little nuclear processing had occurred (in either our own or external galaxies), $Y = 0.23$–0.30, we do not put any further constraints on the model. On the other hand, the similarity of the standard model helium and the observed "primeval" helium may give us some confidence in the essential rightness of that model. ^{3}He is similarly produced in suitable amounts over a fairly wide range of density. Deuterium is more sensitive. Unless the density is relatively low, virtually all the deuterons find other particles and burn through to helium. Thus, if the Big Bang is to produce D/H $\geq 1.4 \times 10^{-5}$ by number (the present interstellar value), the present density of the universe must be less than $0.1\rho_c$, implying an infinite, open universe. The possibility of producing deuterium by other, noncosmological, processes is discussed in Sec. III.B. It is not very encouraging. But it does seem to be possible to put together a coherent picture of deuterium and helium synthesis, the observed values of H and the mass in galaxies, and the age constraints from nucleocosmochronology (discussed in Sec. II.B.2) and the ages of globular clusters, provided that $\rho \sim 0.03\rho_c$ (Gott *et al.*, 1974). The standard model produces rather less ^{7}Li than is observed for this low a density, but having a small fraction of the universe in higher density fluctuations will make adequate ^{7}Li without disturbing any of the other nuclides (Wagoner, 1973 and in NATO, 1974).

It is possible also to discuss the effects of changing the assumptions that led to the standard models. Since synthesis depends, in fact, only on the values of three parameters at the time $T = 10^9$°K, if we can evaluate those parameters under any set of assumptions, then nucleosynthesis under those assumptions will be the same as in the standard model which happens to have the same values of those parameters. The three parameters are the baryon density, the expansion rate of the universe, and the ratio by number of neutrons to protons. Each of the numbered paragraphs which follows is devoted to the effects of violating the correspondingly numbered assumption of the standard model, as outlined by Wagoner (in NATO, 1974).

(1) In the absence of thermal equilibrium at temperatures above the meson rest mass, there are no neutrons and, therefore, no synthesis.

(2) In a matter–antimatter symmetric universe (even if enough material survives annihilation to form galaxies, which is doubtful) synthesis is greatly inhibited because the neutrons are free to cross the matter–antimatter region boundaries and annihilate. Thus, no helium is produced, and, probably, nothing else either, as argued forcefully by Steigman (1973) (but see Aly *et al.*, 1974 for a dissenting view).

(3) A nonzero lepton number which is sufficiently large that the neutrinos are degenerate during synthesis affects things in two ways. First, the density is higher at a given temperature, which increases the expansion rate. Second, and more important, the proton–neutron equilibrium is forced either toward protons (by a sea of neutrinos) or neutrons (by a sea of antineutrinos). Thus, a positive lepton number inhibits all synthesis, while a slightly negative one, by increasing the n/p ratio, increases the amount of helium produced. If the lepton number is very negative, then the neutrons decay so late that the baryon density is too low for particles to find each other easily, and only deuterium (up to 10%) is produced in any quantity. In order for these effects to be important the number density of neutrinos must be comparable with the number density of photons, i.e., $L/B \geq 10^{+9}$ or $\leq -10^{+9}$ (Fowler, 1971).

(4) The effect of new kinds of particles is typically also to inhibit synthesis by delaying the time when protons and neutrons are available to begin reacting. Consider, for example, the statistical bootstrap model of Carlitz *et al.* (1973), in which a single hadron with mass $= 10^{38}\,M_\pi$ fills the entire observable universe at the time when the observable universe has a radius equal to the pion Compton wavelength. The amount of nucleosynthesis then depends on the time scale on which these superbaryons decay to protons and neutrons. If this time scale is greater than about 10^4 sec, the universe will be too cool for helium to form by the time nucleons are available, but lots of deuterium will be made. The decay time favored by Carlitz *et al.* (1973) of 10^7 sec (which has the advantage that galaxies are easily formed because of the large density inhomogeneities associated with the massive hadrons) is, therefore, inconsistent with cosmological synthesis of helium.

(5) If general relativity is not the right theory of gravity, then the expansion rate at a given temperature and density will be different from that in the standard model. This affects the time at which the neutron/proton ratio is frozen out since freeze out occurs essentially when the beta decay time scale and the expansion time scale are equal. For instance, in a theory with a slightly faster expansion rate, freeze out will occur at a higher temperature, with a higher n/p ratio, so that more than the standard amount of helium is formed. Conversely, a slightly slower expansion rate gives less helium. Changing the rate by a factor of 2 changes Y by about 10% either direction. Deuterium and lithium are similarly affected. For expansion rates very different from the standard one, no synthesis occurs. If the rate is too rapid, there is no time for reactions to occur before the universe gets too cool; and if the rate is too slow, there are no neutrons left by the time 10^9°K is reached (remember freeze out occurs near 10^{10}°K). Deuterium survives to the fastest expansion rate, and ^{3}He (since it is proton rich) to the lowest expansion rate.

(6) Inhomogeneities and anisotropy are likely to be the most important violations of the standard assumptions in the real universe. They are also the ones whose effects are most complicated and the hardest to estimate. There are several possibilities. If there are density fluctuations, less deuterium is made in the high regions, but, averaging over all volume elements, we get more deuterium for a given average density of the universe than in the standard case. For a maximally lumpy universe, the limit on the density now, if $D/H \geq 1.4 \times 10^{-5}$, is raised to $\rho \leq \rho_c$. Turbulence will have the opposite effect. If the early universe was turbulent, the synthesis occurred under conditions with entropy per baryon lower than we now see it, the turbulence having necessarily dissipated in the meantime since we don't see it now. Lower entropy per baryon mimics synthesis at a higher density; thus, if the deuterium we see is to have been made cosmologically, the limit on the present density of the universe is even lower than in the standard model. The helium production varies in the opposite sense to the deuterium production as the various parameters are changed, but by smaller amounts. Finally, curvature inhomogeneities, rotation, shear, and other anisotropies will all affect the expansion rate of the universe and, therefore, the nucleosynthesis in the ways mentioned in paragraph (5). Curvature fluctuations, for instance, can reduce helium production by about 30% at a given density (Gisler *et al.*, 1974). Hoyle (1975) presents a more extreme view of the possible effects of inhomogeneities.

We can conclude that, if $Y = 0.23$–0.30 is to be made cosmologically, then something fairly close to the standard model must be correct, at least statistically. If the deuterium is also to be cosmological, then the average density of the universe must now be less than about $0.1\rho_c$. In addition, in the models which make both D and ^{4}He, $^3\mathrm{He}/^4\mathrm{He} \geq 0.7 \times 10^{-5}$, which does not conflict with any of the estimates of its abundance discussed in Sec. II, and $^7\mathrm{Li/H} = (0.04$–$0.61) \times 10^{-9}$, which is only a little less than the value found in very young stars, $\mathrm{Li/H} = 10^{-9}$.

Nucleosynthesis in a standard, low density hot Big Bang universe seems to be in good accord with the observed abundances of H, D, ^{3}He, ^{4}He, and ^{7}Li. All other nuclei are produced in amounts which are too small by many orders of magnitude.

B. Cosmic ray spallation and other versions of the *x* process

In their classic 1957 review of nucleosynthesis in stars, Burbidge, Burbidge, Fowler, and Hoyle (1957) were unable to account for the synthesis of Li, Be, B, and D in terms of any processes occurring in the interiors of normal stars. Rather, these nuclides were always destroyed there, proton bombardment converting them rapidly to helium at any temperature above about 10^6°K. They therefore attributed the formation of these elements to an "x process," which necessarily occurred in a fairly low temperature, low density region, and discussed as possible sites of such a process the atmospheres of stars and bright gaseous nebulae for all four, and, in addition, helium burning zones in evolved stars for lithium. The reactions suggested were spallation in the first two cases and ^{3}He (α,γ) ^{7}Be in the third, the former originally advocated by Hayakawa (1955) and Fowler *et al.* (1955), and the latter by Cameron (1955).

No one process now seems capable of producing all the "difficult" light nuclides, ^{2}H, ^{3}He, ^{4}He, ^{6}Li, ^{7}Li, ^{9}Be, ^{10}B, and ^{11}B. We have already seen the promising role of processes in the early universe in the production of the hydrogen and helium isotopes and ^{7}Li. The other candidates for portions of the x process which we will discuss below are (1) supermassive objects, (2) hydrogen shell burning, (3) novae, (4) supernova shock waves, and (5) cosmic ray spallation. One other candidate, spallation by particles emitted very early in the history of our own and other Solar Systems by active young suns (Fowler *et al.*, 1962) can now be excluded on energetic grounds (Ryter *et al.*, 1970), although Canal (1974) has recently revived the idea in connection with lithium production in young stars.

1. Supermassive stars

A prestellar stage in which nuclear reactions occurred in supermassive objects was originally suggested to account for the nonzero metal abundance in even the oldest galactic stars. The objects are of the same sort as advocated by Hoyle and Fowler (1963) to account for the energy of strong radio sources. Synthesis in such objects, as studied by Wagoner (1969, 1973) and Hoyle and Fowler (1973), is capable of producing interesting amounts of D, ^{3}He, and ^{7}Li (along with plenty of ^{4}He, and heavier elements to the amount $Z = 0.0001$ or so), but only with rather special choices for the mass, n/p ratio, and explosion strength of the supermassive object. There are several difficulties with such a picture. First, it seems that $Z = 0$ objects do not in fact explode in the required fashion (Fricke, 1973). Second, the ratios of C, N, and O isotopes produced (Audouze and Fricke, 1973) in this fashion do not seem to agree with the available observations for very old stars (Cohen and Grasdalen, 1968). And third, the great uniformity of the observed He/H ratio does not seem likely to result from what is essentially local production.

2. Shell burning and shell flashes

An exception to the rule that the light nuclides are only destroyed in stars may occur during the phase when hydrogen is burned in a thin shell around a helium core (in which helium burning may also be taking place), particularly if the products can be mixed to the surface quickly. The abundance of ^{3}He builds up appreciably because of the relatively low temperature at which the hydrogen shell burning takes place, and mixing of that ^{3}He down into the helium region can result in substantial ^{7}Li production through ^{3}He (α,γ) ^{7}Be (e^-,ν) ^{7}Li. This was suggested by Cameron and Fowler (1971) as a way of accounting for the very high lithium abundances (Li/H $= 10^{-7}$) observed in some red giants (Boesgaard, 1970b). This mechanism could also contribute to the general supply of ^{3}He and ^{7}Li through the action of stellar winds. The amount of the contribution is difficult to assess, but obviously in order to build up Li/H $= 10^{-9}$ virtually all the matter in the Galaxy would have had to have been through stars that blew off at least 1% of their mass in such a lithium-enriched stellar wind. Some stars undoubtedly have winds this strong, but it seems difficult to process enough mass through them as the Li is probably destroyed again before the star has evolved much further (Iben, in NATO, 1974).

3. Novae

The promising role of novae in producing the odd-A isotopes of C, N, O, and F will be discussed in Sec. III.C, but Hoyle and Clayton (1974) have recently suggested that reactions at the surface of white dwarfs during nova explosions may also make significant amounts of D and ^{3}He through reactions of the type ^{14}C (p,n) ^{14}N, ^{13}C (p,n) ^{13}N, followed instantly by n (p,γ) D, and then things like ^{2}H (p,γ) ^{3}He, etc. The energy requirements to make interesting amounts of deuterium and ^{3}He are rather severe, amounting to about 3×10^{44} erg per nova outburst, which is an appreciable fraction of the total energy of the event.

4. Supernova shock waves

As the shock wave caused by the detonation of a supernova (whatever its cause) reaches the surface of the object, it will accelerate smaller and smaller amounts of material to higher and higher energies. This suggests that high energy protons might then produce the various light nuclides by spallation under circumstances where they would survive. Colgate (1973, 1974) has suggested this as a possible way of making deuterium. There are two kinds of problems with this sort of model. First, the same high energy protons that make D by spallation of ^{4}He also make Li, Be, B by spallation of C, N, O. At first glance, this sounds like a good thing (Colgate, 1974), since these nuclides are also difficult to come by. Unfortunately, the process makes Li, Be, and B so copiously that if enough material has been through supernova shock waves of the Colgate type to produce the observed deuterium, then we should be swimming in ^{7}Li, ^{9}Be, and ^{11}B (Epstein *et al.*, 1974). ^{7}Li is particularly serious since it is harder to destroy than deuterium and is made from ^{3}He and ^{4}He even if the initial CNO abundance is very low in the supernova material. Further work (Epstein *et al.*, 1975; W. D. Arnett, private communication) indicates that Li and B are always overproduced relative to D for any shock which could make D, but Colgate (cited by Schramm in NATO, 1974) has responded with a model in which the low energy part of the shock (below 5 MeV/nucleon) is quickly cooled by photon diffusion, which reduces Li, Be, B production to an acceptable level, while the higher energy part of the shock produces deuterium. Epstein *et al.* (1975) reply that Li and B are overproduced relative to D for any shock energy, provided only that the matter involved has a Pop I composition ($Z \gtrsim 0.01$). A second difficulty has been pointed out by Weaver and Chapline (1974). In order to get any spallation at all, most of the energy of the shock must be in the ions and not in electrons or photons. This was the case in the Colgate (1973, 1974) model, but Weaver and Chapline find that a self-consistent solution for the propagation of the shock has the ion temperature remaining below about 100 keV, even for very energetic shocks, so that no spallation occurs. The next two stages of this disagreement (as reported by Schramm, in NATO, 1974) have Colgate (1975) pointing out that the realistic case has several ion species (rather than the single one considered by Weaver and Chapline) which interact with the electrons in different ways, resulting in high relative ion velocities for ions of different species (i.e., the required shocks) and Weaver and Chapline responding that Coulomb friction between ion species couples them so that relative ion velocities contain less than 5% of the total available energy, and there are no shocks and no synthesis. It is clear that the last guns of this battle have not yet been fired.

Production of deuterium in supernova shocks does not, therefore, seem terribly promising, but it is perhaps of some interest that ^{7}Li and ^{11}B are produced in their cosmic ratio if the boron abundance is the high one observed in C2 meteorites (Cameron *et al.*, 1973) rather than the low one found for the interstellar medium, the Sun, and Vega. ^{9}Be is still overproduced relative to everything else, but it is more easily destroyed than boron. ^{3}He, ^{6}Li, and ^{10}B are made in less than their cosmic ratio to deuterium (Epstein *et al.*, 1974). Arnould and Nørgaard (1975) also find that ^{7}Li and ^{11}B are made directly from ^{3}He and ^{4}He under explosive conditions, such as might obtain in the outer parts of a supernova. Finally, on the basis of recent measurements of the proton spallation cross sections of ^{13}C and ^{14}N at energies from threshold to about 20 MeV (Jacobs *et al.*, 1974; Oberg *et al.*, 1975), it is possible to say that all of the observed Li, Be, B ratios (including either high or low total boron) by element and by isotope can be matched by the spallation production ratios for some spectrum of low energy projectiles (Bodansky, in NATO, 1974). Thus we cannot exclude the hypothesis that all of the Li, Be, B isotopes were made in a single kind of spallation event with projectile energies of 10–20 MeV and a target containing C, N, and O with terrestrial isotope ratios. The difficulty is to find a suitable site for such a process.

5. Spallation by cosmic rays

This source is an *a priori* promising one. The cosmic rays themselves contain Li, Be, and B made by interactions of relativistic CNO with protons and helium nuclei in the interstellar medium. The relativistic Li, Be, B nuclei will for the most part be lost to the Galaxy on the cosmic ray confinement time scale ($\sim 10^7$ yr), but there must also be corresponding nuclei produced at low energy in the interstellar medium by the interaction of cosmic ray protons and alpha particles with ambient CNO. The importance of this mechanism was established by the calculations of Reeves *et al.* (1970) and Mitler (1970), and a more thorough examination by Meneguzzi *et al.* (1971) established that roughly the required amounts of ^{6}Li, ^{9}Be, and ^{10,11}B (for the lower boron abundance) could be made in this way over the age of the galaxy. Spallation of Fe in the interstellar medium also makes roughly the observed amount of ^{50}V, which is the rarest isotope between carbon and iron, and does not seem to be readily produced by any other process. ^{138}La and ^{180}Ta may also be spallation products (Hainebach *et al.*, 1974a). Most calculations of the total production of spallation product nuclei are made on the assumption that the production rate has been constant in time. Since, on the one hand, the supernova rate (and thus, according to the conventional wisdom, the cosmic ray flux) was probably higher in the past, and, on the other hand, the interstellar CNO abundance was undoubtedly lower in the past, the assumption that the production rate (which depends on cosmic ray flux times CNO abundance) has been constant may well be fairly realistic. Evidence from cosmic ray tracks in meteorites and from the metal abundance in old disk stars suggests that neither thing has varied a great deal in any case. Since the total amount produced is rather uncertain, it is encouraging that the ratios of the nuclides advertized as being produced by cosmic ray spallation are also roughly right. This is shown in Table V (taken from Reeves, in NATO, 1974). Audouze and Tinsley (1974) have included the effects of possible time variations and of Li destruction in stars.

Our present understanding of the "x process" can then be summarized by saying that a low density, hot Big Bang universe is capable of producing the observed amounts of ^{2}H, ^{3}He, ^{4}He, and perhaps, ^{7}Li, while spallation by galactic cosmic rays can be responsible for the remaining light nuclides, ^{6}Li, ^{9}Be, and ^{10,11}B (provided that the cosmic abundance of boron is low). No known process seems capable of making the large amounts of ^{10}B as well as ^{11}B required by Cameron *et al.* (1973); nor does there seem to be any way to make the high Be/B ratio for the Sun implied by the observations of Shipman (1974). Additional contributions, which may not be negligible but are hard to estimate quantitatively, may come to ^{3}He and ^{7}Li from stellar winds, to ^{7}Li and ^{11}B from supernova shock waves, and to all the light nuclides from any process which can fire 10–20 MeV protons at a dilute gas containing CNO. These conclusions are in essential agreement with those of Audouze and Tinsley (1974), Reeves (1974), and other recent model builders.

TABLE V. A comparison of observed values of the abundance ratios of the isotopes of Li, Be, and B with the production ratios for nuclides made by cosmic ray spallation (data from Reeves in NATO, 1974).

Ratio	Observed value	Where observed	Calculated value	Conclusion
^{11}B/^{10}B	4	Meteorites	2–3	OK
^{6}Li/^{9}Be	$\frac{1}{3}$	Meteorites	$\frac{1}{5}$	OK
B/Be	10	Vega	10–20	OK
	100	Meteorites		
	1?	Sun		
^{7}Li/^{6}Li	12.5	Meteorites	2	^{7}Li made elsewhere (cosmological?)

C. Hydrostatic processes in stars

1. Hydrogen burning and the hot CNO cycle

It is now quite generally believed that stars generate most of their energy for most of their lives by the conversion of hydrogen to helium. This was not always the case; worrying about the sources of stellar energy long preceded knowledge of the physics necessary to solve the problem. The chief difficulty is to get energy out for long enough. Ordinary chemical processes fail by many orders of magnitude; for instance, a sun initially made of carbon and oxygen in ratio $C/O = 0.5$ could liberate the solar luminosity (about 4×10^{33} erg/sec) for rather less than 10 000 years by converting the carbon and oxygen to CO_2. This is in good accord only with Archbiship Ussher's estimate of the epoch of creation (4004 BCE).[2] Gravitational contraction from infinite radius to the present size of the Sun provides enough energy for a lifetime of 10^7 years. This time scale suffices to bring the oceans up to their present level of salinity, if the average flow of rivers has always been what it is at present (it hasn't), but fails badly if time is to be allowed for Darwinian evolution, for forming the English Channel, and for bringing various astronomical systems (globular clusters and the like) into the equilibrium configurations they seem to possess by means of gravitational interactions. No process, like accretion, which adds energy from the outside can maintain the thermal gradient necessary for hydrostatic equilibrium in normal stars (Eddington, 1927), however adequate the amount of energy in the suggested sources (e.g., infall of meteoritic material). Complete annihilation of stellar matter was repeatedly suggested by Jeans (1929, and references therein), who thought it the most probable mechanism throughout his life. Complete annihilation of the Sun would, in accordance, with $E = mc^2$, supply the present solar luminosity for 15×10^{12} yr. Jeans was misled firstly by the then-current view that stars evolved from high to low total mass and secondly by some of his own calculations, which seemed to imply the need for time scales of 10^{11-12} yr for binary stars to acquire their very circular orbits and similar dynamical processes. No particular process by which the proposed annihilation might occur was suggested. The necessary time scale was finally put on a firm footing by Rutherford (1929), who estimated the age of the Earth as about 3.4×10^9 yr from the ratio of ^{235}U to ^{207}Pb, but, believing Jean's estimate of the solar age, interpreted it to mean that the Sun was presently producing uranium.

Meanwhile, the correct answer—the building up of heavy elements from light ones—was hit upon by Perrin (1919, 1921). The general idea of "subatomic" energy in stars was widely discussed in the early 1920s, building on the base that "what is possible in the Cavendish Laboratory may not be too difficult in the Sun" (Eddington, 1920). The necessity that the process, whatever it was, be stable but not overstable was emphasized by Russell (1919), along with other conditions which the subatomic source must fulfill. The arguments against helium formation as a primary source of stellar energy, as summarized by Eddington (1926), included the improbability of the hydrogen abundance being high enough and the seeming impossibility of cramming four protons and two electrons close enough together for anything to happen. The chief argument for hydrogen burning (which Eddington himself clearly favored) was that the energy available from it was much larger than for any other sort of subatomic process. This follows from the masses of the nuclides, and the careful measurements of Aston (1920, 1927) with his mass spectrograph (the first) seem to have been particularly influential. Eddington's main objections were taken care of first by Russell's (1929) realizing that the Sun really is mostly hydrogen, and second by Atkinson and Houtermans' (1929; Atkinson, 1931) calculation of the effects of barrier penetration on charged particle reactions. Their picture of fusion reactions, which Atkinson called "regenerative," bore considerable resemblance to the CNO cycle. Von Weizsäcker (1937, 1938), on the other hand, emphasized that, somewhere, sometime, the reaction $^1H + {}^1H \rightarrow {}^2H + e^+$ must have taken place, in order to get the whole process started.

The decisive step in establishing the modern theory of hydrogen burning was taken by Bethe (1939; Bethe and Critchfield, 1938). They considered the reactions among a wide variety of light nuclei (including some now known to be unstable) and concluded that only a very limited selection of reactions could occur at ordinary stellar temperatures ($\sim 2 \times 10^7$ °K) at rates sufficient to account for observed stellar luminosities. Bethe showed that Li, Be, and B were rapidly destroyed ($^7Li + p \rightarrow 2\,{}^4He$, and so forth), contributing relatively little energy, since their abundances are low. Proton captures by elements heavier than fluorine, on the other hand, occur so slowly as to be negligible, because of the high Coulomb barriers, while proton capture by helium is prevented by the absence of a stable mass-5 nucleus. Three-particle charged particle reactions (like the triple-alpha reaction) were similarly very slow at ordinary stellar temperatures and densities. They were left with two kinds of reactions whose rates were large enough to account for stellar luminosities. In Bethe's words, "The first mechanism starts with the combination of two protons to form a deuteron with positron emission, viz., $H + H = D + e^+$. The deuteron is then transformed into He^4 by further capture of protons; these captures occur very rapidly compared with process (1). The second mechanism uses carbon and nitrogen as catalysts, according to the chain reaction

$$\begin{aligned} &C^{12} + H = N^{13} + \gamma, \qquad N^{13} = C^{13} + e^+ \\ &C^{13} + H = N^{14} + \gamma, \\ &N^{14} + H = O^{15} + \gamma, \qquad O^{15} = N^{15} + e^+ \\ &N^{15} + H = C^{12} + He^4. \end{aligned}$$

The catalyst C^{12} is reproduced in all cases except about one in 10,000, therefore the abundance of carbon and nitrogen remains practically unchanged (in comparison with the change of the number of protons)." The rate of the reactions with the CN catalysts was found to depend more steeply on temperature than the rate of the proton–proton interaction. Thus the p–p chain should dominate in low mass stars (which have lower central temperatures) and the CN cycle in high mass stars. Bethe (1939) found the crossover point to occur at, or slightly below, the mass of the Sun. Fowler (1954, 1960) showed that the Sun operates on the p–p chain, and more recent cross-section data confirm that the transition, in fact, occurs near $1.5\,\mathfrak{M}_\odot$, for main sequence stars with Pop I abundances of CNO.

[2] And, if it is right, the end of the world is at hand.

Hydrogen burning, therefore, has at least three effects on the composition of material in stars. First, the Li, Be, B, and deuterium will be burned through to ^{4}He in all regions where the temperature is above about $10^{6\circ}$K, and the ^{3}He as well at slightly higher temperatures. Thus, the amounts of these fragile nuclides remaining in the interstellar medium provide important constraints on how much of the medium has been in stars (provided the production mechanisms, discussed in Sec. III.B, are understood). The tendency for surface Li (and perhaps B and Be as well) abundances to be lower in old main sequence stars than in young ones (the solar surface lithium abundance is about 1% of that in young stars) then also suggests mixing of the surface material into a hot region at some stage, resulting in rather complete destruction of these nuclides throughout the star, and more severe constraints on "astration."

The second effect is the obvious one of increasing the helium abundance. The influence of this on "cosmic" abundances does not appear to be as great as one might at first think. We are led by several lines of evidence to conclude that the supply of helium in the Galaxy has not been much increased by hydrogen burning in ordinary stars. First, as discussed in Sec. III.A, the Galaxy probably started life with nearly its present He/H ratio as a result of nucleosynthesis in the early universe. Second, to have made most of the helium by stellar hydrogen burning we now see would have resulted in galaxies being very much brighter in the past than they are now, to the point where their red-shifted luminosity would now probably violate the observed upper limits to the fluctuations in the night sky infrared background (Partridge, 1974; Davis and Wilkinson, 1974), but Kaufman (1975) suggests that the energy may come out at unobservable wavelengths. Third, strangely enough, stars in the end really don't make very much helium. The point is, that by the time they have advanced to a state of evolution where the return of large amounts of mass to the interstellar medium is probable (second ascent of the giant branch, presupernova, or whatever), the hydrogen-burning and helium-burning shells are very close together, so that only a few percent of the star's mass is in the form of helium (Paczyński, 1970, and many other model calculations). Just how much helium is returned for a given amount of mass put into stars will depend on the details of the stellar models; which mass ranges blow up completely and which shed only part of their mass (and how much); the fraction of mass put into high and low mass stars; and so forth. Current estimates (Arnett, in NATO, 1974, and see further in Sec. III.F below) suggest that each generation of stars will increase the interstellar helium abundance by not much more than it does the interstellar heavy element abundance ($\Delta Y = 1\text{–}2\Delta Z$). Thus, if the metals we now see ($Z \sim 0.03$) were all made in normal stars, the expected change in Y is 0.03 to 0.06. This is compatible with the difference in helium abundance between old and young objects in our own Galaxy (or between low- and high-metal galaxies) according to some of the evidence discussed in Secs. II.C and II.F. It is also compatible with no detectable changes in Y having occurred over the history of the Galaxy or between galaxies, as suggested by some other parts of the observational evidence.

The third abundance change results only from hydrogen burning by the CNO cycle and affects the relative amounts of the catalyst nuclides, the isotopes of C, N, and O (and, under special circumstances, F, Ne, Na, and Mg). The general principle is, that as hydrogen burning continues, an equilibrium situation will be gradually built up, in which each reaction in the cycle occurs at the same rate. Thus the abundances of the individual nuclides must be inversely proportional to their proton capture cross sections at stellar temperatures, entirely independent of the initial proportions of C, N, and O or their isotopic ratios. This situation has long been believed to be responsible for the production of ^{14}N from the C and O made by helium burning in previous generations of stars (Sec. III.C.2 below); hence the statement in Sec. II that nitrogen is a secondary element, whose abundance can be expected to vary more strongly from place to place than that of things that can be made from hydrogen or helium in a single generation of stars.

The relatively small proton capture cross section of ^{14}N was known to Bethe (1939) and to B^{2}FH (Burbidge *et al.*, 1957). Precise estimates of the ratios of all the CNO nuclides to be expected in equilibrium require good laboratory cross sections. Recent progress in this area includes the remeasurements of ^{15}N (p,γ) ^{16}O and ^{17}O (p,γ) ^{18}F by Rolfs and Rodney (1947a,b) which imply increases in the rates of both these processes relative to the competing ones. Using these new cross sections, we find that CNO hydrogen burning operates through the following tri-cycle, with the branching ratio at ^{15}N being about 880:1 in favor of (p,α) and the ratio at ^{17}O very possibly of order 1:1:

$$
\begin{array}{lll}
{}^{12}\mathrm{C}(p,\gamma)\,{}^{13}\mathrm{N} & & \\
{}^{13}\mathrm{N}(e^+\nu)\,{}^{13}\mathrm{C} & & \\
{}^{13}\mathrm{C}(p,\gamma)\,{}^{14}\mathrm{N} & & \\
{}^{14}\mathrm{N}(p,\gamma)\,{}^{15}\mathrm{O} & & \\
{}^{15}\mathrm{O}(e^+\nu)\,{}^{15}\mathrm{N} & & \\
{}^{15}\mathrm{N}(p,\alpha)\,{}^{12}\mathrm{C} \text{ or } & {}^{15}\mathrm{N}(p,\gamma)\,{}^{16}\mathrm{O} & \\
 & {}^{16}\mathrm{O}(p,\gamma)\,{}^{17}\mathrm{F} & \\
 & {}^{17}\mathrm{F}(e^+\nu)\,{}^{17}\mathrm{O} & \\
 & {}^{17}\mathrm{O}(p,\alpha)\,{}^{14}\mathrm{N} \text{ or } & {}^{17}\mathrm{O}(p,\gamma)\,{}^{18}\mathrm{F} \\
 & & {}^{18}\mathrm{F}(e^+\nu)\,{}^{18}\mathrm{O} \\
 & & {}^{18}\mathrm{O}(p,\alpha)\,{}^{15}\mathrm{N}.
\end{array}
$$

The resulting equilibrium ratios of the nuclides involved are shown in Table VI, along with the observed values of

TABLE VI. A comparison of the observed ratios of the CNO isotopes in Solar System material with those calculated for material which has acted as a catalyst in CNO tricycle burning of hydrogen under equilibrium conditions. CNO hydrogen burning is believed to be responsible for the production of ^{14}N from ^{12}C and ^{16}O. It appears also to make an adequate amount of ^{17}O and about half the observed ^{13}C. ^{15}N, ^{12}C, and ^{16,18}O must be made elsewhere.

Ratio	Observed[a]	Calculated[b]
^{12}C/^{13}C	90[c]	4
^{13}C/^{14}N	0.036	0.018
^{15}N/^{14}N	0.0036	0.0002
^{17}O/^{14}N	0.022	0.020
^{18}O/^{14}N	0.012	Very small
^{14}N/^{16}O	0.17[d]	60

[a] Cameron (1973).
[b] C. A. Barnes (in NATO, 1974) and Truran (1973).
[c] But 5–10 in some evolved stars (Sec. II.C).
[d] But > 1 in some evolved stars (Sec. II.C).

the ratios. The great enhancement of ^{14}N confirms hydrogen burning as a likely mechanism for its production (in addition, no other process seems to make it). ^{17}O and perhaps ^{13}C are also made in suitable amounts, relative to ^{14}N, by this process. These nuclides must be returned to the interstellar medium (presumably by stellar winds and planetary nebulae) prior to any explosion of the star which made them since they will be destroyed again in the regions which explode. The observed enhancements of ^{13}C, nitrogen, and ^{17}O in some evolved stars (where mixing down to the hydrogen burning region is likely to have occurred) substantiate the production mechanism for these nuclides. The simultaneous occurrence of nitrogen and helium enhancements in planetary nebulae is also encouraging. ^{15}N and ^{18}O do not seem to be adequately produced by the CNO cycle in normal stars. They will be discussed later in this section (and ^{18}O also under helium burning).

In the outer regions of a star's hydrogen burning zone, the temperature will be sufficiently low that equilibrium is not reached, but significant amounts of ^{13}C will be formed due to that reaction in the cycle which has the lowest Coulomb barrier. Thus mixing that extends part way down into the hydrogen-burning zone can be expected to increase the surface ^{13}C/^{12}C ratio without otherwise affecting abundances. We apparently see evidence of this in the otherwise normal-looking giants (like Arcturus; see Sec. II.C) which show ^{13}C enhancements. Slightly deeper mixing will bring up extra ^{17}O in amounts which may depend critically on the mass of the star (Dearborn, in NATO, 1974 and Wollman, 1973). Other implications of the CNO tri-cycle cross sections will be considered by Fowler *et al.* (1975).

Some doubt has been cast on the entire scheme of hydrogen burning in stars by the failure of R. Davis to detect any neutrinos coming from the Sun (see, for instance, the review by Bahcall and Sears, 1972). Neutrinos with energies above the threshold of Davis' experiment (0.814 MeV) are produced by the CNO cycle (which does not matter, since the Sun does not operate on the CNO cycle), but also by three sub-branches of the p–p chain, which should occur in the Sun:

$$
\begin{array}{llllll}
{}^1\mathrm{H}+{}^1\mathrm{H}\rightarrow{}^2\mathrm{D}+e^++\nu & \text{or} & {}^1\mathrm{H}+{}^1\mathrm{H}+e^-\rightarrow{}^2\mathrm{D}+\nu\ * & & & \\
{}^2\mathrm{D}+{}^1\mathrm{H}\rightarrow{}^3\mathrm{He}+\gamma & & & & & \\
{}^3\mathrm{He}+{}^3\mathrm{He}\rightarrow{}^4\mathrm{He}+2\,{}^1\mathrm{H}+\gamma & \text{or} & {}^3\mathrm{He}+{}^4\mathrm{He}\rightarrow{}^7\mathrm{Be}+\gamma & & & \\
 & & {}^7\mathrm{Be}+e^-\rightarrow{}^7\mathrm{Li}+\nu\ * & \text{or} & {}^7\mathrm{Be}+{}^1\mathrm{H}\rightarrow{}^8\mathrm{B}+\gamma & \\
 & & {}^7\mathrm{Li}+{}^1\mathrm{H}\rightarrow{}^8\mathrm{Be}^*\rightarrow 2\,{}^4\mathrm{He} & & {}^8\mathrm{B}\rightarrow{}^8\mathrm{Be}^*+e^++\nu\ * & \\
 & & & & {}^8\mathrm{Be}^*\rightarrow 2\,{}^4\mathrm{He}. &
\end{array}
$$

The three critical neutrinos are those emitted in the starred reactions. The rates of the latter two reactions are very dependent on the central temperature of the Sun, and a variety of changes in solar model can reduce their predicted fluxes (though no single change suffices to bring the predicted flux of about 5 SNU below the observed upper limit of 1 SNU; Iben, 1972; Ulrich, 1974a). A SNU is a Solar Neutrino Unit, and represents 10^{-36} captures per target nucleus per second. The theory behind all this is explained in some detail by Bahcall and Sears (1972). The cross sections and rates for the nuclear reactions involved are all known adequately (Kavanagh, 1972) and cannot be blamed for the discrepancy. The neutrinos generated by the $p+e+p$ reaction do not depend on temperature very much, and their flux tests only the hypothesis that the Sun is presently generating its observed luminosity by the p–p chain. Their predicted flux is less (but only a little less) than Davis' upper limit. The only single hypothesis thus far published that reduces the predicted flux below the observed upper limit by changing the solar model is the suggestion by Fowler (1972a) that the Sun is subject to an instability which occasionally expands and cools its interior, reducing its neutrino production rate. This can be done by mixing fresh hydrogen into the center. All the excitement must be over in about 10^7 yr since the photons take that long to random walk their way from the solar interior to the surface, while the neutrinos (if any!) come straight out. The physical process which causes the required mixing is not understood, and the resulting evolution of the Sun may violate other things we think we know about stellar evolution. Hypotheses which require changing the form of the weak interaction so that the neutrinos somehow do not reach us, also present a variety of difficulties. Perhaps we should concur with Feynman (1974) in saying that a 2σ effect is not worth a new theory, but the apparent absence of solar neutrinos is at present a serious problem, which may yet require us to rethink our entire picture of stellar energy generation and nucleosynthesis. Another neutrino detection experiment, done by some technique other than ^{37}Cl, would be a great help in clarifying the nature of the problem.

At a sufficiently high temperature, hydrogen can also burn by a Ne–Na cycle, similar to the CN one (Marion and Fowler, 1957): $^{20}\mathrm{Ne}(p,\gamma)$ $^{21}\mathrm{Na}(e^+\nu)$ $^{21}\mathrm{Ne}(p,\gamma)$ $^{22}\mathrm{Na}(e^+\nu)$ $^{22}\mathrm{Ne}(p,\gamma)$ $^{23}\mathrm{Na}(p,\alpha)$ ^{20}Ne. There is a small amount of leakage via $^{23}\mathrm{Na}(p,\gamma)$ ^{24}Mg which makes it, in fact, a Ne–Na–Mg bi-cycle. Marion and Fowler (1957) and B^{2}FH thought it probable that substantial amounts of ^{21}Ne might be produced by this process. It requires that the initial CNO abundance be very low, so that there is time for the slower-burning Ne to get involved. A good estimate of the isotope ratios to be expected from the Ne–Na cycle in or out of equilibrium must wait for adequate measurements of the cross sections involved. A start on this has been made by Rolfs and Winkler (1974), whose result for $^{20}\mathrm{Ne}(p,\gamma)$ ^{21}Na shows that it is the first measured charged particle cross section to be dominated by a subthreshold resonance.

The CNO cycle (and/or Ne–Na–Mg cycle) can also occur under very nonequilibrium, explosive conditions, where the cooling time scale is so short that, once a proton capture has occurred, there may not be time for a beta decay and another proton capture before the material has cooled, resulting in enhancements of the odd-A nuclei. Possible sites for such a process are the surface of a neutron star which is accreting matter (Thorne and Zytkow, 1974), supermassive objects (Audouze and Fricke, 1973, and references therein), and novae (Starrfield *et al.*, 1972; Hoyle and Clayton, 1974). The reactions occur at very high temperatures under these circumstances, facilitating barrier penetration in the heavier reactants. The relevant reactions and their rates have been discussed by Audouze *et al.* (1973). Interesting amounts of ^{15}N, ^{17}O, ^{19}F, and ^{21}Ne appear to be produced. In order to reach roughly solar abundance of these nuclei, it is necessary that about 1% of the mass of the Galaxy have been through the hot CNO cycle in a region where the abundances of the capturing nuclei (^{12}C, ^{14}N, ^{16}O, ^{20}Ne) already have roughly their solar value (Andouze, in NATO, 1974). This requirement is somewhat difficult to meet in any of the three sites, but perhaps least so for the novae. At least 0.02 $M_\odot$ must, on average, be processed through the hot CNO cycle in each nova explosion. Observations of ^{13}C/^{12}C and ^{15}N/^{14}N in the nova DQ Herculis near maximum light provide some weak support for novae as a site of the hot CNO cycle (Sneden and Lambert, 1975).

At the moment, then, the CNO cycle or one of its variants appears to be the most promising source of all isotopes of C, N, O, and F (except ^{12}C, ^{16}O, and ^{18}O) and of ^{21}Ne. The other effects of hydrogen burning are a drastic reduction in Li, Be, B, D, and ^{3}He and a mild increase in ^{4}He.

2. Helium burning

The energy released by hydrogen fusion balances the luminosity radiated from a star's surface and maintains the star in a roughly constant configuration until the hydrogen is exhausted in a core containing about 10% of the mass of the star. The reason for the existence of some such critical core mass was first explored by Schönberg and Chandrasekhar (1942). The length of time which a star can spend on the main sequence therefore depends on both its mass and its luminosity, and is shortest for the most massive stars, because it is observed (and can be understood theoretically; see e.g., Eddington, 1926) that the luminosity, L, scales with the mass, M, about as $L/L_\odot = (M/M_\odot)^{3.5}$. The main sequence lifetime for the Sun, a G star, is about 10^{10} years, and it increases or decreases by roughly a factor of 10 for each spectral type away from G (10^9 yr at F; 10^7 yr at B; 10^{11} yr at K; etc.). Following the exhaustion of hydrogen in a star's core, its center contracts (to supply, from gravitational potential, the energy being lost from the surface) and its outer layers expand. It therefore moves into the red giant region of the Hertzsprung–Russell diagram. The position of a single star on the HR diagram and the appearance of the diagram for a cluster of stars are, therefore, age indicators. The ages of the oldest stars in our Galaxy found in this way are in reasonably good accord with other things we think we know about the age of the universe (Gott *et al.*, 1974), and the appearance of HR diagrams of clusters of various ages can be understood in considerable detail (see, for instance, Cox and Guili, 1968). Since the contraction of the core of the star inevitably heats it, it seems natural now to consider the possibility of additional nuclear reactions occurring at the higher temperatures.

The synthesis of elements beyond helium in stars was first attempted in a coherent way by Sterne (1933). His calculation was an equilibrium one, as he believed the nonequilibrium situation (in which reactions and their converses do not occur about equally often) considered by Atkinson and Houtermans (1929) would probably be overstable. What he seems to have had in mind was a gradual liberation of nuclear energy as a star shrank and heated and the equilibrium composition shifted toward heavier elements; he believed that elements at least as far as Zn would thereby be obtained in roughly the right proportions. The modern history of synthesis of the heavy elements in stars begins with the work of Hoyle (1946), who considered the situation in a star which collapses following hydrogen exhaustion. He also believed that in at least some cases high enough densities ($\sim 10^7$ g cm^{-3}) and temperatures ($\sim 4 \times 10^9$°K) would be reached that a statistical equilibrium would be established among the nuclides. Subsequent rotational instability would then return the products to the interstellar medium. The right idea—a sequence of reactions, gradually building heavier elements—was finally grasped and sold to the astronomical community by Salpeter (1952, 1953), who showed that the triple-alpha reaction ($3\,{}^4\text{He} \rightarrow {}^{12}\text{C}$), originally suggested by Bethe (1939), would produce useful amounts of energy at temperatures near 2×10^8°K and calculated its rate under stellar conditions. The critical point is that ^{8}Be is unbound by less than 100 keV, so that at high temperatures there is a small equilibrium concentration of it (about one part in 10^{10} of the ^{4}He at 10^8°K). The capture of a third helium nucleus to give ^{12}C is then exothermic. Salpeter (1952) also considered qualitatively the capture of further alpha particles by the ^{12}C and the reactions among the carbon, oxygen, neon, etc., nuclei themselves at temperatures near 10^9°K. Simultaneously, Alpher and Herman (1953) remarked that "element synthesis in the interior of common stellar types does not appear possible."

Salpeter (1953) considered helium burning in a star which had been well mixed throughout its main sequence life, and which only left the main sequence when virtually all the hydrogen in the entire star had been converted to helium. Such well-mixed stars never become red giants, leaving a large hole in our understanding of stellar evolution. The role of a composition discontinuity between a helium core and a hydrogen-rich envelope in producing the distended red giant structure had been elucidated by Öpik (1938), Hoyle and Lyttleton (1942, 1949), Li and Schwarzschild (1949), and Bondi and Bondi (1950). Later calculations of helium burning (B^{2}FH, and references therein) took account of this. Meanwhile, Öpik (1951) had studied the rates of alpha particle reactions in the right sort of stars, ones with composition discontinuities, and found that the triple-alpha reaction as well as ^{12}C$(\alpha,\gamma)\,^{16}$O, ^{16}O$(\alpha,\gamma)\,^{20}$Ne and so forth would result in the building up of significant amounts of heavy elements in ordinary stars before the white dwarf stage was reached. There seem to be two reasons that his work was not as influential as, in retrospect, it looks like it ought to have been. First, he was not aware of the resonances in the reactions forming the ^{8}Be and ^{12}C (see B^{2}FH) which considerably increase the rate of the reactions, so that his triple-alpha reaction required a temperature of $(4\text{–}6) \times 10^8$°K to produce significant carbon, while carbon is already being destroyed by ^{12}C$(\alpha,\gamma)\,^{16}$O near 2×10^8°K. Second, very few people read the Proceedings of the Royal Irish Academy.

Helium burning releases only about 10% as much energy per gram as hydrogen burning. This follows directly from the measured masses of the nuclei involved. In addition, stars are typically more luminous (by factors of about 10) in the helium burning phase than on the hydrogen-burning main sequence so that the helium-burning lifetime is only about 1% of the hydrogen-burning one. Core helium burning nevertheless makes a major contribution to the luminosity of the stars like the Sun during the "horizontal branch" phase (so-called because the stars occupy a horizontal strip on the Hertzsprung–Russell diagram) and to the luminosity of more massive stars during the red giant phase, the chief difference between the two being that in low-mass stars helium is not ignited until the center has already contracted to the point where the electrons are highly degenerate, resulting in a sort of explosion. The rest of the luminosity is generated by hydrogen burning in a shell around the helium core. The reasons for all these things are explained in some detail by, e.g., Cox and Guili (1968).

The products of helium burning are clearly observed on the surface of some highly evolved stars, for instance the hydrogen-deficient carbon stars and the R CrB variables (Sec. II.C) whose atmospheres appear to be about 50% helium and 50% carbon. More subtle manifestations occur in other stars with moderate carbon enhancements. These stars must reflect either mixing or the loss of an outer envelope (or both), but it is not clear which effect dominates.

As usual, understanding of the details of stellar energy generation from helium burning and of the composition of its products has had to wait for accurate laboratory measurements of the cross sections involved. The two main products of helium burning are surely ^{12}C and ^{16}O, but the proportions of each depend on the balance of the rates of $3\,^4He \rightarrow {}^{12}C$ and $^{12}C\,(\alpha,\gamma)\,^{16}O$ as a function of temperature. Qualitatively, it is clear that at the lowest temperatures, all the alpha particles will fuse (very slowly!) to form carbon, without any of them ever having sufficient energy to penetrate the Coulomb barrier and make oxygen, while at very high temperatures, ^{12}C will burn on up to ^{16}O as fast as it is made. The energy generation rate does not depend very much on which product dominates, but a knowledge of the C/O ratio at the end of helium burning is crucial for understanding which nuclear reactions occur next, and under what conditions. B^{2}FH believed that ^{20}Ne would also be a major product of helium burning, but this is no longer thought to be the case (Truran, 1973), due to the absence of levels with suitable spin and parity in the compound nucleus.

The measured cross sections are still not quite as accurate as the situation requires. The triple-alpha reaction rate is probably adequately known and is about a factor of 2 slower than the value implied by Fowler *et al.* (1967). Recent measurements have been made at Caltech and at the University of Washington (Barnes and Nichols, 1973; Chamberlin *et al.*, 1974). The $^{12}C\,(\alpha,\gamma)\,^{16}O$ rate is less satisfactory. The main contribution to the reaction at stellar temperatures comes from the tail of a 7.12 MeV excited state of ^{16}O (the ground state, as usual, has the wrong spin and parity), and experimental results are often parametrized in terms of the reduced alpha width of this state as a fraction of the single particle limit. The parameter is called θ_α^2 and is likely to fall between zero and unity. A small value corresponds to a slow rate for $^{12}C\,(\alpha,\gamma)\,^{16}O$, and, therefore, mostly carbon in the product, and a large value, to almost pure oxygen in the product. Experimental values of θ_α^2 in the past 5 years have ranged from 0.02 to 0.80, and stellar evolution calculations have frequently and arbitrarily assumed that carbon and oxygen are produced in equal amounts by weight (Truran, 1973). The most recent and, with luck, the most accurate laboratory measurement of the reaction rate (Dyer and Barnes, 1974), when interpreted in terms of hybrid optical-model-*R*-matrix theory (Doonin *et al.*, 1974), corresponds to $\theta_\alpha^2 = 0.053$ (Fowler *et al.*, 1975).

The resulting carbon:oxygen ratios as a function of stellar mass have been calculated by Dyer and Barnes (1974), based on Arnett's (1972) models of the evolution of helium cores. The final composition is about 70% carbon for stars less massive than $15\,M_\odot$, dropping to 50% at about $40\,M_\odot$. The usual half-and-half assumption was therefore both a good and a lucky guess. Helium shell burning flashes in low-mass stars probably produce a larger fraction of carbon than helium burning under other circumstances (Christy-Sackmann and Paczyński, 1975).

During helium burning, reactions like $^{14}N(\alpha,\gamma)\,^{18}F(e^+\nu)\,^{18}O(\alpha,\gamma)\,^{22}Ne(\alpha,n)\,^{25}Mg$ will also occur due to the high ^{14}N abundance left behind by CNO cycle hydrogen burning. Reactions of this type have been studied in massive stars by Peters (1968), by Couch *et al.* (1974), and by Gallino (in NATO, 1974), and have three significant effects. First, they appear capable of supplying the observed amounts of ^{25}Mg, ^{18}O, and ^{22}Ne. Second, some of the products are useful seed nuclei for the URCA process as discussed in Sec. III.D.1 below. And, third, they provide a source of neutrons which can interact with ambient iron-peak nuclei, building up heavier elements by the *s* process (Sec. III.E). The flux of neutrons made available in this way is not adequate to account for all the synthesis of heavy elements which has occurred by this process over galactic history, and other sources will be discussed in Sec. III.E.

Helium burning can also occur explosively when a supernova shock wave passes through the helium-rich region of a massive, evolved star. The situation is rather similar to explosive hydrogen burning. ^{14}N (the most abundant heavy nucleus, as a result of the CNO cycle) captures one or two alpha particles, sheds protons or neutrons, and the zone cools before beta decays have a chance to occur. Synthesis under these circumstances has been considered by Howard *et al.* (1971) and by Arnould and Beelen (1974). The products depend somewhat on the initial values of the temperature and density as the shock passes through the region, and on the cooling time scale. Howard *et al.* (1971) found that, under some circumstances, ^{15}N, ^{18}O, ^{19}F, and ^{21}Ne were made (as ^{15}O, ^{19}F, ^{19}Ne, and ^{21}Ne) in roughly their solar ratio, and suggested explosive helium burning as the most probable source of these nuclides. No ^{13}C or ^{17}O was made (being bypassed in $^{14}N + {}^4He$), but these are just the nuclei we saw were abundantly made by explosive hydrogen burning. ^{14}N and ^{22}Ne were not made by this process either, but they are major products of hydrostatic hydrogen and helium burning, respectively, and can be returned to the interstellar medium either peacefully in stellar winds or, unprocessed, from the outer regions of supernovae. Arnould and Beelen (1974) have found that after the shock wave has passed through the outer layers of an evolved star (from the helium shell burning region to the surface) the totality of the matter involved has roughly the solar ratios of all C, N, O, F, and Ne isotopes except ^{16}O and ^{20}Ne, which must be made elsewhere (helium burning and carbon burning, respectively), for reasonable values of the initial parameters of the model. One of their models produced neon having isotope ratios very much like that of the peculiar Ne–E meteoritic component (Sec. II.B.2; $^{20}Ne/^{22}Ne \lesssim 3.4$; $^{21}Ne/^{22}Ne \lesssim 0.1$). The associated oxygen had an unusually low ^{18}O abundance, which may also be relevant to the ^{16}O-rich meteoritic component.

Helium burning is, then, a significant source of stellar luminosity during reasonably well understood evolutionary phases. It results in the production of ^{12}C and ^{16}O in roughly equal amounts, some of which will be built up to heavier nuclides in subsequent processes (to be discussed in the following sections) and some of which, once returned to the interstellar medium, will be converted to other isotopes of C, N, and O through hydrogen burning (Sec. III.C.1).

Helium burning, either static or explosive, also appears to produce exactly the rarer isotopes of C, N, O, F, and Ne which are not made by hydrogen burning. Some of the products of hydrostatic helium burning can also be returned directly to the interstellar medium via stellar winds since we see them enhanced on the surfaces of some evolved stars. The astrophysical implications of the helium-burning cross sections will be further discussed by Fowler *et al.* (1975).

3. Hydrostatic carbon, oxygen, and silicon burning

B²FH originally proposed that, after the exhaustion of helium in the core of massive stars, further contraction would occur, accompanied by an increase in central temperature up to about 10^9°K. At this temperature, the photons are sufficiently energetic to promote $^{20}Ne\ (\gamma,\alpha)\ ^{16}O$. Since their helium-burning calculations had resulted in a neon abundance comparable with that of carbon and oxygen, large numbers of alpha particles thereby became available, and were recaptured by other neon (as well as carbon and oxygen) nuclei, building up heavy elements. The net result of $^{20}Ne\ (\gamma, ^{16}O)\ \alpha\ (^{20}Ne, \gamma)\ ^{24}Mg$ is the liberation of about 4.6 MeV. At gradually increasing temperatures up to 3×10^9°K, the synthesis of ^{28}Si, ^{32}S, ^{36}Ar, and ^{40}Ca would also occur, each new nuclide being formed from the previous one by the capture of an alpha particle released by photo disintegration. Under these circumstances, the excess abundances of these "alpha nuclei" relative to their neighbors and the gradual decrease in abundance from ^{20}Ne to ^{40}Ca could readily be understood. B²FH recognized the possibility of reactions among the ^{12}C and ^{16}O nuclei, but believed these would not become important until temperatures near 3×10^9°K were reached. Subsequent cross section measurements have demonstrated that $^{12}C + ^{12}C \rightarrow ^{20}Ne + \alpha$, etc., begin to occur at about 8×10^8°K, before photodisintegration becomes important. $^{12}C + ^{12}C$ and $^{16}O + ^{16}O$ were first correctly treated by Hoyle (1954), and hydrostatic carbon and oxygen burning were studied by Hayashi *et al.* (1958) and by Reeves and Salpeter (1959).

Nuclear reactions of carbon, oxygen, and heavier elements release still less energy per gram at each stage than helium burning, while the stars involved continue to be much more luminous than they were on the main sequence. Evolutionary stages in which stellar luminosity is derived from heavy element reactions therefore proceed very rapidly, so rapidly that we do not expect to "catch" any stars in these phases unless they call attention to themselves by some sort of violent behavior. The HR diagram thus ceases to be a useful tool for following the stars' evolution. All models, however, indicate that the stars should remain very distended and red, as well as luminous, during these late phases. These stars must, therefore, be sought among the red giants, but they will not necessarily look very different from red giant stars undergoing shell hydrogen or helium burning. Because the late stages proceed so rapidly, very little change occurs in the outer layers of the star during them (for instance, the hydrogen-burning shell does not have time to work its way very much farther out). It is, therefore, possible to approximate such a star by a "helium" star, whose mass is equal to the mass of the core in which hydrogen was exhausted on the main sequence. Ignition of carbon and heavier fuels can occur either hydrostatically or explosively. The explosive case is discussed in Sec. III.D. It can occur either due to the passage of a shock wave or because a fuel has begun to burn in a degenerate core. All results of the calculations of hydrostatic heavy element burning must be treated as provisional until the new reaction rates of Fowler *et al.* (1975) have been incorporated into them.

Evolution of stars sufficiently massive that they surely ignite carbon before their cores become degenerate ($M \geqslant 9\,\mathfrak{M}_\odot$, Paczyński, 1971c; $M \gtrsim 6\,\mathfrak{M}_\odot$, Arnett, 1974a) has been studied by Sugimoto (1971, and references therein), by Ikeuchi *et al.* (1972, and references therein), and by Arnett (1974b, and references therein; Couch and Arnett, 1972). The results, including the compositional structure as a function of mass at various evolutionary epochs, seem to be in good agreement among the three groups. This is encouraging and means that the results are not too sensitive to reactions rates, treatment of convection, and other input physics which were somewhat different in the three studies. Our discussion here follows the work of Arnett (1974b; in NATO, 1974, 1975; and assorted preprints), which has carried the widest variety of core masses (4, 8, 16, 32, 64, and $100\,\mathfrak{M}_\odot$) to the end of their hydrostatic evolution. Preliminary study of 3 and $2.25\,\mathfrak{M}_\odot$ cores, representing 12 and $10\,\mathfrak{M}_\odot$ main sequence stars, suggests that they follow the same pattern as the $4\,\mathfrak{M}_\odot$ core. The subsequent hydrostatic evolution has also been followed for some of the cores, including the effects of neutral currents (Schramm and Arnett, 1975; Tubbs and Schramm, 1975). The eventual goal of such studies is to be able to predict how much mass is trapped in a remnant, how much is expelled into interstellar space in the hydrodynamic phases (the answer "none" is not excluded for either of these), and what the composition of the expelled material is, all as a function of the initial stellar mass.

This program is far from complete, but a number of interesting results have already emerged. It has been known for some time that the detailed distribution of the products of hydrostatic carbon (Arnett and Truran, 1969) and oxygen (Woosley *et al.*, 1972) burning in the range $A = 20$–40 does not agree with "cosmic" abundances.

TABLE VII. Comparison of the compositions of massive stars at the completion of hydro-static heavy element burning with the composition of Solar System material and the cosmic ray sources. The 8 $M_\odot$ core corresponds to a 22 $M_\odot$ main sequence star, and represents the mass-weighted average presupernova. The 4 $M_\odot$ core corresponds to a 15 $M_\odot$ main sequence star and is the number-averaged presupernova. Calculated ratios and interpretations of the observations from W. D. Arnett (in NATO, 1974).

Ratio	8 $M_\odot$ core	Solar System	Conclusion
He/O	1.9	25	He cosmological
C/O	0.35	0.41	OK
Ne/O	0.20	0.24	OK
Mg/O	0.15	0.075	Not bad
Si group/O	0.28	0.082	Some Si and Fe left
Fe group/O	0.80	0.140	behind in remnant
Ratio	4 $M_\odot$ core	Cosmic ray sources	Conclusion
C/O	1.1	0.71	OK
Ne/O	0.23	0.24	OK
Mg/O	0.41	0.44	OK

The material that contributed heavy elements to the Solar System must, therefore, have undergone additional nuclear reactions under explosive conditions during the event that returned it to the interstellar medium. The results of these processes are discussed in Sec. III.D.

On the other hand, the general features of the abundance distribution bear interesting resemblances to the observations. The calculated distribution depends, of course, on the core mass. We can pick particular cores which ought to be representative under various circumstances. Consider the case in which all stars more massive than $8\,\mathfrak{M}_\odot$ undergo supernova explosions which return most of their mass to the insterstellar medium, leaving behind a pulsar. The amount of processed material thrown off under these assumptions is a steeply increasing function of stellar mass (see, e.g., Fig. 1 of Talbot and Arnett, 1974), while the number of stars available to do it is a steeply decreasing function of stellar mass. The "average" bit of matter returned will, therefore, have lived in some intermediate sort of star, whose core mass is about $8\,\mathfrak{M}_\odot$ ($M \sim 22\,\mathfrak{M}_\odot$ on the main sequence). The first part of Table VII compares the total composition of the Arnett $8\,\mathfrak{M}_\odot$ core at the end of silicon burning to that of the Solar System. The general distribution among C, O, Ne, and Mg is quite similar, while the extra Si and Fe in the model suggests that quite a lot of the center of the star remains behind in the pulsar. In order to make everything match, $1.47\,\mathfrak{M}_\odot$ of Si and Fe group elements must remain behind in the remnant. This is a not implausible number, but, clearly, detailed comparison of theory and observation must await a better understanding of what makes supernovae explode and what determines the nature and mass of the remnants. Arnett and Schramm (1973) have also suggested that we should compare the cosmic ray "source" composition with the calculations. On the assumption that each supernova leaves a pulsar and that each pulsar can accelerate the same number of cosmic ray particles, we would expect the cosmic ray "source" composition to be given by an average core where the average is taken over the number of stars instead of over mass. The number-average core has a mass of about $4\,\mathfrak{M}_\odot$, corresponding to a main sequence mass of about $13\,\mathfrak{M}_\odot$. Its over-all composition is compared with the cosmic ray source composition in the second part of Table VII. Again, there is reasonable agreement for C, O, Ne, and Mg. The amounts of Si and Fe group elements would again be excessive, implying a stellar remnant.

Figure 10 shows the distribution of the various major constituents in the $8\,\mathfrak{M}_\odot$ core after the silicon-burning flash. Rarer nuclides, which are not shown, are also included in the reaction networks. The large regions with roughly constant composition are in convective equilibrium. The shells where the reactions are occurring are all rather thin by this time, which is one of the many features that makes the evolution difficult to follow (it is worst for the smallest core masses). "Si" represents the elements Si to Ca and "Ni" the iron-peak elements (most of the iron is made as ^{56}Ni, which then decays to ^{56}Fe).

Adequate laboratory cross sections are something of a problem for these heavy element reactions, as they were at earlier stages. $^{12}C + {}^{12}C$ is discussed in Sec. III.D. $^{12}C + {}^{16}O$ has recently been remeasured by Cujec and Barnes (1974), and the oxygen burning reactions by Spinka and Winkler (1972). Work on these and other relevant reactions is currently in progress at several laboratories, and up-to-date results and their implications will be given by Fowler *et al.* (1975).

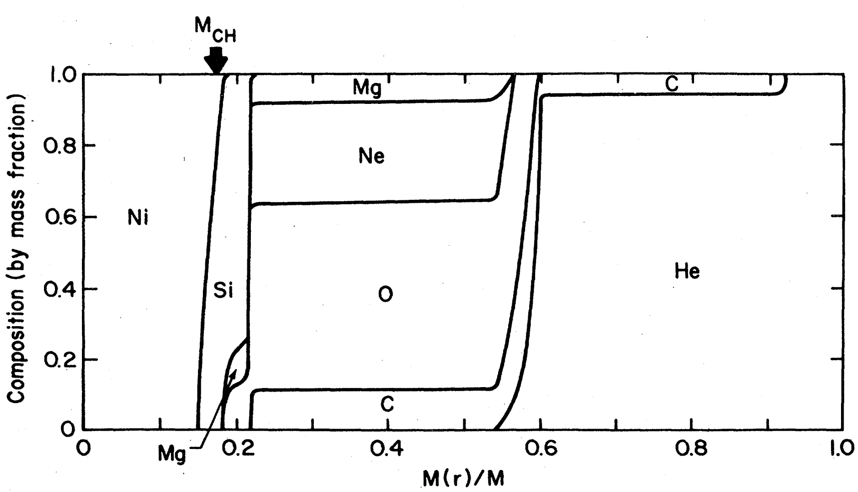

FIG. 10. Composition by mass fraction of the 8 $M_\odot$ helium core studied by Arnett (1974b and references therein) after silicon burning. If the inner 1.4 $M_\odot$ (indicated by an arrow and M_{CH}) remains behind in a remnant, this star will eject into the interstellar medium about 3 $M_\odot$ of intermediate weight elements (C, O, Ne, Mg), a comparable amount of helium, and about half a solar mass of Si and heavier elements. Si represents the group of elements from Si to Ca, and Ni the iron-peak group of elements.

Hydrostatic or mildly hydrodynamic burning (the cross-over occurs when the nuclear time scale becomes less than the sound-travel time across the star) of the elements from carbon to silicon provides the main source of energy during a brief but vital stage in the late evolution of massive stars. During these phases more energy is lost to the star by neutrinos than by photons. The distribution of the products of these reactions bears a general (but not detailed) resemblance to their "cosmic" abundances if the average star, by mass, is considered and to the cosmic ray "source" abundances if the average star, by number, is considered. In addition, these products are the raw material upon which explosive nucleosynthesis (which has been remarkably successful in accounting for the details of the Solar System abundances) must work during the hydrodynamic phase that destroys the star and returns matter to the interstellar medium, and the temperature and density condition in the various zones of the star are initial conditions for the explosion. Most urgently needed are more and better cross sections [measured, for preference, but significant progress is presently being made toward being able to calculate them adequately (Woosley *et al.*, 1975)] and further Arnett-type models of the stars which will include enough of the relevant physics to determine the mass and nature of the stellar remnant as a function of star mass.

D. Violent processes in stars

Explosive burning of a fuel can occur either because it is suddenly heated to a high temperature or because it ignites where the matter is degenerate. In the latter case, the gas pressure is not influenced by the temperature (until the temperature gets high enough to lift the degeneracy), so that the burning region does not expand and cool to maintain hydrostatic equilibrium. Instead it gets hotter and hotter and the reactions occur faster and faster. Sudden heating can also occur because of the passage of a supernova shock wave, as we have already discussed for explosive hydrogen and helium burning, and this will be the main site of explosive heavy element burning. It is clear from two points of view that an explosion must occur. First, the synthesis products must somehow be returned to the interstellar medium from a region where the gravitational binding energy is rather high, and, second, the expansion velocities of typical supernova remnants imply an energy to mass ratio that, if stored as thermal energy, yields a high enough temperature for explosions to occur. The effects of ignition of a fuel under degenerate conditions vary. Helium is so ignited in low-mass stars, producing a "helium flash." The luminosity of the star is not increased by this, but it is sufficiently shaken up to move rapidly from the red giant branch to the horizontal branch in the HR diagram. The effects for neon ignition and oxygen ignition (Arnett, 1974b), which occur when the mass inside the dominant shell-burning source (usually carbon) reaches about the Chandrasekhar mass, are similarly mild. But the result of degenerate carbon ignition (which was first considered by Hoyle and Fowler, 1960) appears to be disastrous. The star may be completely disrupted.

1. Explosive carbon burning

a. Carbon detonation in degenerate cores. Weigert (1966) first demonstrated that neutrino energy losses would cool the cores of intermediate mass stars (his model was for $5\,\mathfrak{M}_\odot$) enough to prevent carbon ignition in the phase of core contraction immediately following helium exhaustion in the center. The problem is that the contraction is rather slow, so that neutrino losses can more than keep up with the energy released from gravitational potential. (In more massive stars, contraction is more rapid, and the net result is heating and peaceful carbon ignition as discussed in the previous section.) One of two things can happen to the star with a central temperature inversion established in this way. Carbon may ignite off-center, thereby heating and expanding the core again, so that it, too, burns carbon peacefully. All will be well in this case. Or, as Arnett (1968, 1969) and Rose (1968, 1969) pointed out, the subsequent growth of the core mass as the helium burning shell works its way out through the star and the increase of central density lead to a slow increase of central temperature and to carbon ignition at the center under extremely degenerate conditions. The mass range in which these occur was estimated by Paczyński (1969, 1970) as 4–8 or $9\,\mathfrak{M}_\odot$. We will come back in a moment to how you tell which happens. In the latter case, there are, once again, two possibilities. Arnett (1969) originally found a violent thermal runaway in which the star was completely disrupted by core carbon ignition when the degenerate core mass had reached about $1.4\,\mathfrak{M}_\odot$. (The similarity to the Chandrasekhar limit is not an accident.) The central temperature climbs so high during this runaway that the entire core is processed up to iron-peak elements, liberating $1.4\,\mathfrak{M}_\odot$ of Fe per such explosion. If this happened very often, we would be swimming in iron (Arnett 1969, 1974a).

A possible alternative was suggested by Paczyński (1972). The degenerate carbon–oxygen core will already contain a small amount of ^{23}Na and similar nuclei, and more will immediately be produced after ignition by $^{12}C + {}^{12}C \rightarrow p + {}^{23}Na$. At densities where the Fermi energy of the degenerate electron gas is higher than the threshold for electron captures, $^{23}Na + e^- \rightarrow {}^{23}Ne + \nu$. In a convective core, the Ne will rapidly be carried to a lower density region and decay back again; $^{23}Ne \rightarrow {}^{23}Na + e^- + \bar{\nu}$.

The process can happen repeatedly, leaving the Na and the electrons unharmed, but releasing two energetic neutrinos each time. These leave the star, and continuously cool the core. ^{21}Ne behaves similarly. Such cycles were first thought of by Gamow and Schoenberg (1940, 1941) and are called URCA processes (not an acronym, but the name of a Brazilian casino, since the process is analogous to the loss of money when gambling). Their rates have been calculated by Tsuruta and Cameron (1970).

It is clearly possible, in principle, for such cycling to cool the center enough for the carbon burning to proceed to completion at a roughly constant temperature, no explosion ever occurring. In some cases it may be possible (Ergma and Paczyński, 1974) to exhaust the carbon in an inner convective core of about $0.4\,\mathfrak{M}_\odot$. The subsequent collapse may then detonate carbon in the outer core, which will disrupt the remainder of the star, leaving the inner core behind as a low-mass neutron star, since the density was already well above the white dwarf range ($\gtrsim 10^9$ g cm^{-3}) during the degenerate carbon burning. Barkat *et al.* (1974) and Buchler and Mazurek (1974) also find that part of the core can survive explosive carbon detonation. Alternatively, quasistatic processes could continue on up to the iron-peak elements, whereupon electron capture would yield gravitational collapse and a more massive neutron star (Arnett, in NATO, 1974). The details of the calculation are rather complicated, and there has been some debate in the literature over whether the URCA processes in fact give any cooling at all, and whether it is really enough to help. The most recent word to date (Couch and Arnett, 1974, and Arnett, in NATO, 1974) is that there is some cooling, but it only delays and does not prevent the thermonuclear runaway, which probably disrupts the star. The inclusion of neutral current effects changes the neutrino emission rates by a negligible factor in these stars (Dicus, 1972; Schramm, in NATO, 1974).

The critical problem is then to establish the mass range over which these "carbon detonation supernovae" occur. There are two kinds of indirect evidence that it must be a rather small range of masses, M_1 to M_2, which disrupt in this way. First, if $M_2 - M_1$ is much larger than $0.5\,\mathfrak{M}_\odot$, the production of iron-peak elements over the history of our Galaxy greatly exceeds the amount of iron that we see around now (Arnett, 1974a). Second, if all the stars in the range 4–8 $\mathfrak{M}_\odot$ or so blow apart and leave no condensed remnant, it is hard to understand where all those pulsars come from, as there may not be enough stars above $8\,\mathfrak{M}_\odot$ to give rise to them, according to Ostriker *et al.* (1974). They suggest that we must at least lower M_2 to around $6\,\mathfrak{M}_\odot$. A discussion by Biermann and Tinsley (1975) of the positions of Type II supernovae in spiral arms leads to the conclusion that their progenitors have masses of $8\,\mathfrak{M}_\odot$ and larger.

From the point of view of stellar evolution calculations, the two limiting masses are set by rather different kinds of processes. We know (or think we know) that stars like the Sun shed some portion of their outer layers in a planetary nebula during their second (helium-shell-burning) ascent of the giant branch and settle down to useful lives as white dwarfs without ever having burned carbon at all. Any star which can shed enough mass to bring its remaining core below the Chandrasekhar limit before carbon ignition occurs will never burn carbon and will end its life as a white dwarf (Paczyński, 1970, and many other calculations). This sets M_1. We therefore "know" that it must be greater than $1.4\,\mathfrak{M}_\odot$, the Chandrasekar limit for helium, carbon, and the like. We can increase M_1 a little further by looking at young galactic clusters. If there are white dwarfs in these clusters, they must surely have been formed from stars more massive than the present main sequence turnoff. The existence of a dozen or more white dwarfs in the Hyades thus allows us to conclude that M_1 is at least 2–2.5 $\mathfrak{M}_\odot$ (Auer and Woolf, 1965; van den Heuvel, 1975) and probably 3–4 $\mathfrak{M}_\odot$ (Tinsley, 1974a; van den Heuvel, 1975a). There is also one white dwarf in the same direction in the sky as the Pleiades. It is not a binary (Greenstein, 1974), its brightness is not inconsistent with its being at the distance of the cluster (Eggen and Greenstein, 1965), and it shares the cluster proper motion (Jones, 1973). If it is indeed a cluster member, then we can raise M_1 to about $6\,\mathfrak{M}_\odot$, on the clearly unjustifiable assumption that all stars of the same mass do the same thing (Woolf, 1974).

M_2, on the other hand, is set by the lowest temperature (and, therefore, the lowest mass star) for which energy generation by $^{12}C + {}^{12}C$ exceeds energy loss by neutrino emission due to the universal Fermi interaction (Beaudet *et al.*, 1967). This, in turn, depends on weak interaction theory and on the cross section for the reactions. Measurements of the $^{12}C + {}^{12}C$ reactions at low energies have been carried out by Patterson *et al.* (1969) and by Mazarakis and Stephens (1972), down to a center of mass energy of 2.45 MeV. Plausible extrapolations of the measurements already differ by an order of magnitude at a center of mass energy of 2 MeV (an effective temperature of about 8×10^{8}°K) and diverge rapidly at lower effective temperatures. The current state of our understanding of heavy-ion nuclear reactions does not permit us to choose an unambiguous "best extrapolation" (see Fowler, 1974). The most optimistic extrapolation for our purposes, the one which gives the largest rate at low temperatures, is due to Michaud (1972). It results in lowering M_2 to about $7\,\mathfrak{M}_\odot$ (Arnett, 1974a).

The range of stellar masses in which disruptive carbon ignition can be expected to occur is, therefore, somewhere between 2.5–9 $\mathfrak{M}_\odot$ and 6–7 $\mathfrak{M}_\odot$, where each of these limits, in turn, has roughly a ± 1 attached to it. Only if something like the narrower range (or even $M_2 - M_1 = 0$) is correct can we easily understand where the pulsars come from and why there isn't a good deal more iron than there is.

b. Explosive carbon burning in massive stars. The central regions of a massive star are capable of undergoing static nuclear reactions all the way up to the iron-peak elements. Beyond that, no further nuclear reactions can be exothermic since these nuclei are the most tightly bound ones of all. A dense, inert iron core gradually builds up. When its mass reaches roughly the Chandrasekhar limit, it becomes unstable to a variety of disasters (reviewed by Arnett, 1973a), all of which have a general tendency to cause rapid collapse of the core and the liberation of large amounts of gravitational potential energy in some form. The news of this disaster (and some of the energy) propagates out through the star in the form of a shock wave. Additional energy may be deposited in the outer layers by neutrinos. All this provides at least a qualitative motivation for considering the effects on the composition of the outer layers of an evolved massive star which would be caused by a sudden heating of the layers, followed by expansion and cooling on roughly the hydrodynamic time scale. The general scheme is usually called explosive nucleosynthesis and has been reviewed by Arnett (1973) and Clayton and Woosley (1974).

It is customary to discuss explosive nucleosynthesis in a parametrized framework. That is, instead of starting with values of density, temperature, and so forth from definite stellar models, one usually selects values of four parameters, and then goes back later and compares them to stellar models. The four parameters are initial temperature T_0; initial density ρ_0; the time scale on which these change, τ, as a fraction of the hydrodynamic time scale $\tau_H = 446\rho^{-\frac{1}{2}}$ sec [thus $\tau = \chi\tau_H$, $\rho = \rho_0 \exp(-t/\tau)$ and $T = T_0 \exp(-t/3\tau)$]; and the excess of neutrons per nucleon, $\eta = (N - Z)/(N + Z)$. The four parameters suffice to determine uniquely the products of the explosion of a particular dominant fuel, provided that all the reaction cross sections are adequately known. This is virtually never the case, partly because the number of nuclear physics laboratories is finite and rather small, and partly because many of the necessary reactions involve target nuclei in excited states or unstable target nuclei, and these are nearly impossible to study in the laboratory. Many of the required cross sections must, therefore, be calculated. In the absence of a complete theory of the nuclear force, a semiempirical approach is used. The most recent efforts in this direction will be discussed by Woosley *et al.* (1975). The general idea is to represent the behavior of nuclei in some formalism (usually the Hauser–Feshbach one) whose parameters are determined empirically and then to use that formalism to interpolate or extrapolate the properties of nuclei for which the experimental data are inadequate. This works well only in a statistical sense. The estimated reaction rates are, therefore, best for (a) heavier nuclei, since complex systems tend to have more states, and (b) high temperatures, since more levels will then be populated. Luckily, these are roughly the conditions under which we want to study explosions anyway. The formalism and some of the problems are further discussed by Fowler (1972). Once the parameters are chosen and the reaction rates (as functions of temperature and density) estimated, differential equations for the time evolution of the composition can be written down. These are numerous, complicated, and nonlinear, and have coefficients which depend very steeply on the (rapidly changing) temperature. Analytic solutions are not possible, and various numerical methods have been evolved. These usually replace derivatives by finite differences over small time steps. The computer then does the rest.

The first reaction network calculation at explosive temperatures was the one by Truran *et al.* (1966) of the approach to equilibrium in pure ^{28}Si heated to $T = (3 \text{ or } 5) \times 10^9\,°$K. Shortly thereafter, Truran *et al.* (1967) followed the evolution of the composition as the matter gradually cooled (freeze out) after being heated to a peak temperature $T_0 = 5 \times 10^9\,°$K behind a supernova shock wave. The final composition hardly depended on the nature of the initial fuel for this high a value of T_0.

Explosive burning of a dominantly carbon fuel at temperatures to be expected in real supernovae ($T_0 \sim 2 \times 10^9\,°$K) was calculated by Arnett (1969a) and was the first such process to give detailed agreement with Solar System abundances for a group of intermediate mass nuclei. Pardo *et al.* (1974) have recently recalculated nucleosynthesis during explosive carbon burning. They find that the Solar System abundances of ^{20}Ne, ^{23}Na, ^{24,25,26}Mg, ^{27}Al, and ^{29,30}S are all produced by an explosion of fuel of about equal parts of carbon and oxygen and the following values of the parameters described above: $T_0 = 2 \times 10^9\,°$K, $\rho_0 = 10^5$ g cm^{-3}, $\chi = 1$, and $\eta = 2 \times 10^{-3}$. The extra neutrons are stored in the form of about 2% ^{22}Ne (the natural product of the ^{14}N left behind the CNO cycle hydrogen burning after it has been through a hydrostatic helium burning region). The temperature is the most critical parameter, but χ and η are also constrained to within a factor of about 2 if the Solar System abundances are to be matched. Density is not very critical. At slightly higher temperatures ($T = (2.2\text{–}2.4) \times 10^9\,°$K), but similar values of the other parameters, explosive carbon burning favors slightly heavier nuclei and produces significant amount of ^{28}Si, ^{31}P, ^{32,33,34,36}S, and ^{35}Cl. The present range of uncertainty in supernova models is such that it is possible that the parts of the star that reach this temperature may still contain significant amounts of carbon, although oxygen seems more probable. Heating to much higher temperatures ($T \sim (2.6\text{–}3.2) \times 10^9\,°$K) overproduces sulfur and ^{35}Cl by such large factors that it is hard to believe that material from a carbon zone heated this much can have made any substantial contribution to the Solar System. This is not surprising in stellar model terms: the regions that are dense enough to get themselves heated this much by the shock contain only oxygen and no carbon. We will use these values of the parameters again toward the end of the next section when we compare the conditions needed for explosive nucleosynthesis to those found in stellar models.

At first glance, there seems to be some fundamental inconsistency between the results quoted here for the isotopes of Ne, Na, Mg, Al, and S and the claim made toward the end of Sec. III.C that the relative proportions of ^{12}C, ^{16}O, ^{20}Ne, ^{24}Mg, and ^{28}Si in a massive star at the completion of hydrostatic burning were in good accord with their Solar System values. In fact, all is well. The explosive conditions convert these common isotopes into the rare ones quite efficiently, so that only a small portion of the ejected stellar material has to be processed in the explosion to bring the rare isotopes up to their solar abundances. Relatively little matter is, in fact, processed because the material cools so quickly. Thus the ratios of the abundant nuclei will not be disturbed by the explosive nucleosynthesis which puts the "finishing touches" on the abundance distribution and makes it match the details of the observed distribution.

In a star with Population I composition, the matter which undergoes explosive carbon burning (ECB) will have about 0.3% of its mass in the form of elements heavier than Mg, which have remained untouched by the hydrostatic stages. These will be exposed to the free protons, neutrons, and alpha particles released during explosive carbon burning. The result is the production of significant amounts of many of the rarer isotopes of elements from S to Ge. Howard (1973) has considered these reactions. The nuclides are not produced in ratios which bear any detailed resemblance to their solar abundances, but the average level of "overproduction" (ratio of mass fraction of the product in the processed material to its mass fraction in "cosmic" material) of these rare nuclides is, typically, factors of 100–200 for ECB at temperatures $T_0 = (2.0\text{–}2.2) \times 10^9\,°$K. This is comparable with the overproduction of the more abundant nuclides like ^{20}Ne, ^{24}Mg, and ^{23}Na which are made by explosive carbon burning. Thus, if the processing necessary to make these has occurred (about 1% of Solar System matter through an ECB zone), interesting amounts of the following will also have been made: ^{40}Ar, ^{40}K, ^{43,46,48}Ca, ^{45}Sc, ^{47,49,50}Ti, ^{50}V, ^{62,64}Ni, ^{65}Cu, ^{67,68,70}Zn, ^{69,71}Ga, ^{73,76}Ge, and ^{75}As. They are made primarily from ^{32}S, ^{36}Ar, ^{40}Ca, Fe, and Ni seeds and include many nuclei which are not readily made by other processes. The conditions of ECB have little effect on nuclei with $A > 70$. In particular, the classic p-process nuclei cannot be produced under the same conditions that produce the Mg isotopes.

We will defer comparing the calculated and observed abundances of ECB nuclei graphically until after the next section, so as to be able to include the effects of oxygen and silicon burning as well. As in the case of hydrostatic carbon and oxygen burning, these calculations must be regarded as provisional until the cross sections of Fowler *et al.* (1975) are included. Some of the problems are discussed by Fowler (1974).

2. Explosive oxygen and silicon burning and the *e* process

Nearer the center of the star than the regions which undergo explosive carbon burning, hydrostatic reactions will leave behind heavier elements, primarily ^{16}O, ^{24}Mg, and ^{28}Si (with some excess neutrons stored in ^{18}O, ^{26}Mg, or ^{30}Si). These regions are also denser and hotter than the carbon-rich ones at the onset of hydrodynamic evolution and will be heated still more by the supernova event itself. Thus it makes sense to consider nuclear processes in matter (consisting of oxygen, magnesium, and silicon in various proportions) which is heated suddenly and then expands and cools on something like the hydrodynamic time scale. Processes of this type have been extensively treated by Woosley *et al.* (1973). Depending on the fuels present, and on the maximum temperature reached (which, of course, determines which things can undergo reactions), the processes can conveniently be called explosive oxygen burning, explosive silicon burning, and the *e* (for equilibrium) process. These processes produce most of the abundant nuclides with atomic weight, A, between 28 and 62 (Si to Ni). Clearly, in any real star, there will be a continuum of processes. There is a continuum in the calculations as well, in the sense that oxygen (or even carbon), if heated sufficiently at high density and then cooled hydrodynamically, will yield much the same final composition as silicon treated the same way.

As in the case of explosive carbon burning, the calculations are done in terms of several parameters, T_0, ρ_0, η, χ, and the initial composition, and the necessary values of the parameters compared with the conditions inside stellar models later. Silicon burning differs fundamentally from the earlier stages in that the dominant reaction is not $^{28}\mathrm{Si} + {}^{28}\mathrm{Si}$, by analogy with $^{12}\mathrm{C} + {}^{12}\mathrm{C}$, etc., but rather the photodisintegration of the silicon into alpha particles which then interact with the remaining ^{28}Si to build various heavier nuclei (Fowler and Hoyle, 1964). The usual problems with inadequate cross sections exist, although not such serious ones as might be supposed, since, at the very high temperatures involved, many reactions are nearly in equilibrium with their inverses, so that the product yield of the nuclei involved depends only on their masses, which are well known. Variations in the values of the important, poorly known cross sections have been included as additional parameters in the calculations of Woosley *et al.* (1973). Cross sections of special interest have been tabulated by Clayton and Woosley (1974).

Two alternatives to this continuum of processes have been explored. B^{2}FH originally thought in terms of the addition of successive alpha particles to ^{16}O, followed by heating to temperatures above $\sim 3 \times 10^9$°K for long enough that all nuclei, free protons, and free neutrons would come into equilibrium. This resulted in the synthesis of all isotopes of the iron-peak elements (V, Cr, Fe, Co, and Ni) in their solar proportions, as those were then understood (Goldberg *et al.*, 1960). Subsequent increases in the photospheric iron and other abundances have badly disturbed that agreement. The ^{56}Fe at the center of the peak was made as iron, and a large value of η, ~ 0.07, was required in this picture. More recently, it has been found that the sum of a number of mass zones that have undergone silicon burning to various extents and achieved a state of equilibrium only for species in the mass range $28 \leq A \leq 62$ (and only with respect to the exchange of protons, neutrons, alphas, and photons) can produce most of the abundant species in the mass range of interest in roughly their solar proportions. This was the original approach of Truran *et al.* (1966). It has since been followed up by Bodansky *et al.* (1968) and by Michaud and Fowler (1972). In this picture, the $A = 4n$ nuclei and the iron peak are made simultaneously, the ^{56}Fe is made as ^{56}Ni that subsequently decays, and a much lower value of η, ~ 0.002 (much closer to what would be expected to come out of the earlier hydrostatic stages) was sufficient.

The approach of Woosley *et al.* (1973), considering a variety of fuels, appears to be closer to what we would expect to encounter in a real star and will be followed here. The complexity of the problem is indicated by the fact that a typical reaction network in one of the explosive nucleosynthesis calculations involves 90–100 nuclides (about half stable and half unstable) each of which can undergo 12 different reactions with n's, p's, α's, and γ's and two beta decays. Of these reactions, Woosley *et al.* (1973) identify about 260 as important in determining the final products and 36 as very important. At the lowest temperatures considered, clusters of nearby nuclei came into equilibrium, while at the highest temperatures, most of the range $A = 28$–62 came into equilibrium at the beginning, and then gradually froze out. The most successful runs had equilibrium over $A = 28$–45. It is characteristic of all the processes considered that a variation in one parameter can be largely compensated by a variation in another, so that roughly "cosmic" abundances can be obtained over a moderately wide range of conditions, such as would be expected in stars with a range of main sequence masses, compositions, and so forth.

Four distinct processes can be regarded as contributing to the Solar System composition. They are treated in the four lettered paragraphs that follow.

(a) Explosive oxygen burning (first considered by Truran and Arnett, 1970) operates on the products of hydrostatic carbon burning. A typical initial composition represents a zone in which helium burning originally produced roughly equal amounts of ^{12}C and ^{16}O which then underwent carbon burning. After sufficient time has elapsed for the ^{20}Ne created during that phase to be photodisintegrated, the composition, by mass, will be 54% ^{16}O, 30% ^{24}Mg, 2% ^{26}Mg (carrying the extra neutrons; this implies $\eta = 0.002$), and 14% ^{28}Si. For this choice of composition and η, $T_0 = 3.6 \times 10^9$°K, $\rho_0 = 2 \times 10^5$ g cm^{-3}, and $\chi = 1$, the explosion produced the Solar System abundances of ^{28}Si, ^{32,34}S, ^{35,37}Cl,

^{36,38}Ar, ^{40,42}Ca, and ^{46}Ti to within about 25%, when normalized to ^{28}Si. ^{33}S, ^{39,41}K, and ^{50}Cr were made in amounts deficient by factors of 2–3. Nothing was overproduced, and the agreement of the calculated and measured isotope ratios of chlorine, argon, and potassium is particularly striking. Woosley *et al.* (1973) explore in considerable detail the permissible variations in the parameters. The only part of the initial composition that matters very much is the neutron excess (but 0.002 is precisely what one would expect for Pop I matter that had been through CNO cycle hydrogen burning and had all its CNO turned into ^{14}N thereby). The other parameters interrelate in a rather complicated way (for instance, if the time scale is longer, the matter spends more time in the highest density and temperature part of its history, and so the peak temperature and density must be decreased to keep the good fit), but, in general, compensations can be made to keep the good abundance agreement over a range in T_0 of about 10%, and ranges in time scale and density of factors up to 10.

(b) Explosive silicon burning (first studied by Truran *et al.*, 1966) operates on the products of either hydrostatic carbon burning or hydrostatic oxygen burning, or some combination thereof. The two standard initial compositions considered by Woosley *et al.* (1973) were pure Si (with 3% ^{30}Si carrying the extra neutrons) and a mixture of O, Mg, and Si in the same proportions as in paragraph a. The two were virtually indistinguishable after 0.001 sec. In fact, any combination of parameters that resulted in roughly equal amounts of ^{28}Si and ^{56}Ni at the beginning of freezeout worked about equally well. Thus, the two defining characteristics of explosive silicon burning are that the material must be heated hot enough that the entire range of atomic weights $A = 28$–62 comes into equilibrium and that the mixture be frozen out well before all the silicon is destroyed. The dominant nuclear process is the "melting" of alpha particles off the ^{28}Si by (γ,α) reactions and the recapture of the alphas by other ^{28}Si nuclei, and whatever other heavy nuclei there are around. A typical successful run had $T_0 = 4.7 \times 10^{9}$°K, $\rho_0 = 2 \times 10^7$ g cm^{-3}, $\chi = 4.45$, and $\eta = 0.002$. None of these (except η) are critical, as long as the two fundamental conditions are met. Under these circumstances, the solar abundances of ^{32}S, ^{36}Ar, ^{40}Ca, ^{52}Cr, and ^{54}Fe are produced, and slightly deficient amounts (factors of 1.5 to 3) of ^{28}Si, ^{48}Ti, ^{50}Cr, ^{51}V, ^{54}Mn, and ^{56}Fe. The most important underproductions are ^{44}Ca, ^{54}Cr, ^{58}Fe, and ^{58}Ni. The last will come from further processes to be discussed, but the first three remain unproduced by any process at low values of η.

(c) Nuclear statistical equilibrium with normal freeze out (NSE1) occurs when the fuel (1) gets hot enough for equilibrium to be established, (2) stays there long enough for all the silicon to be destroyed, and (3) freezes out at high enough density that alphas are captured as quickly as they can be melted off things. The termination of equilibrium during cooling is then caused by a deficiency of alpha particles, all of which are bound into species that are characteristic of statistical equilibrium at a higher temperature. The results do not depend on initial composition except for the value of η, the neutron excess. A typical run, with $T_0 = 8.1 \times 10^{9}$°K, $\rho_0 = 6.2 \times 10^8$ g/cm^3, $\chi = 1.0$, and $\eta = 0.002$ makes the Solar System amounts of ^{52,53}Cr, ^{54}Mn, ^{54,56,57}Fe, and slightly deficient (factors 2–3) amounts of ^{48,49}Ti, ^{50}Cr, ^{51}V, and ^{58}Ni when normalized at ^{56}Fe. Only the value of η is critical, provided that the three conditions mentioned are met.

(d) Nuclear statistical equilibrium with alpha-rich freezeout (NSE2) occurs when the fuel (1) gets hot enough for equilibrium over the range $A = 28$–62, (2) stays there long enough for all the silicon to be destroyed, and (3) freezes out at low enough density that not all the alphas made can find heavier nuclei to capture them. Equilibrium is then terminated by the capture reactions, unlike the case in paragraph c. To meet these conditions it is necessary that ρ_0 meet the criterion:

$$\rho_0 < \min 4.5 \times 10^{-22} T_0{}^3$$
$$< \min 2.5 \times 10^{-31} T_0{}^4 \chi^{-\frac{2}{3}}.$$

A typical run, with $T_0 = 5.5 \times 10^{9}$°K, $\rho_0 = 2 \times 10^7$ g cm^{-3}, $\chi = 4.45$, and $\eta = 0.00153$ yielded Solar System amounts of ^{56,57}Fe, ^{59}Co, and ^{58,60,61,62}Ni. As unual, only η and the general criteria are critical. The original treatment of the equilibrium process by B^{2}FH had a much higher value of η. This is successful in making the otherwise missing ^{58}Fe, but greatly overproduces ^{60}Ni, ^{48}Cr, and the like.

In the innermost parts of the silicon-burning region and of a real star, densities and temperatures are likely to be high enough for significant amounts of electron capture to occur. This suggests that we should perhaps consider the effects of varying η across the NSE1 or NSE2 zone; or, for simplicity, try combining various amounts of several zones with different values of η. A program of this type has been carried out by Hainebach *et al.* (1974) with some success. A two-zone model, with more than 90% of the matter in the usual low-η conditions and a small amount of matter at $\eta = 0.08$, subjected to NSE1 is a considerable improvement. NSE2 conditions have not been as thoroughly explored. The best fit of all was obtained from a three-zone model with 86.3% of the matter at $\eta = 0.0033$, 11.4% at 0.0668, and 2.3% at 0.0769. The exact values of these are not terribly important, but the large gap between the low and high values of η is necessary to prevent overproduction of ^{54}Fe. This model gave Solar System proportions of ^{51}V, ^{50}Ti, all isotopes of Cr and Fe, ^{55}Mn, and ^{58,60,62}Ni. No ^{59}Co or ^{61}Ni were produced. They must come from NSE2. Further increases in the number of zones would not improve things further.

Two more tasks remain to complete our discussion of explosive carbon, oxygen, and silicon burning. First, we must compare the conditions which have been found in the preceding sections to be necessary for explosive nucleosynthesis with the conditions in highly evolved stars, in order to see whether supernova explosions of massive stars are in fact a suitable site for the processes we have been discussing. This is done in Table VIII, which contains data (from Arnett, in NATO, 1974) on the conditions in the various zones of 4 and 8 $\mathfrak{M}_\odot$ cores just prior to the beginning of hydrodynamic evolution. The values of the parameters T_0, ρ_0, and η found to be necessary for the various processes are also given. The parameter "T" in the core models is the temperature that will be achieved

TABLE VIII. Comparison of conditions required for explosive nucleosynthesis with those in the relevant zones of evolved cores of massive stars. T_0 and ρ_0 are the initial temperature and density in region exploding on a hydrodynamic time scale and η is the neutron excess $(n_n - n_p)/(n_n + n_p)$. T and "T" are the temperatures in the stellar zones before and after the addition of the energy from the explosive nuclear processes. The 8 $M_\odot$ core corresponds to a 22 $M_\odot$ main sequence star and is the mass-weighted average presupernova. The 4 $M_\odot$ core corresponds to a 15 $M_\odot$ main sequence star and is the number-averaged presupernova. The conditions in the 8 $M_\odot$ core are in good accord with the requirements, while those in the 4 $M_\odot$ core are not. Thus if each supernova accelerates an equal number of cosmic rays, the cosmic ray abundances of the products of explosive nucleosynthesis could be quite different from the Solar System abundances. NSE1 and NSE2 are nuclear statistic equilibrium with normal and alpha-rich freezeout, respectively. Data from Arnett (in NATO, 1974).

	Explosive nucleosynthesis			
Process	$^{12}C + ^{12}C$	$^{16}O + ^{16}O$	NSE1	NSE2
T_0 (°K)	2×10^9	$(3.1\text{–}3.9) \times 10^9$	$\geq 5 \times 10^9$	$\geq 5 \times 10^9$
ρ_0 (g cm^{-3})	10^{4-7}	10^{5-7}	$\geq 5 \times 10^7$	$\leq 5 \times 10^7$
η	0.001–0.004	0.002–0.004	0.002–0.004	0.002–0.004

Zones of 8 $M_\odot$ core					
Zone	He	C	Ne	O	Si
T	0.14×10^9	1.1×10^9	1.1×10^9	2.8×10^9	4.0×10^9
"T"	...	1.8×10^9	1.8×10^9	4.0×10^9	7.0×10^9
ρ	1.3×10^2	9.4×10^4	1.1×10^5	2.8×10^6	5.9×10^7
η	...	≥ 0.0011	0.0019–0.0027	0.0019–0.0027	$\geq$0.0019–0.0027
Mass (in 10^{33} g)	9.6	3.5	3.5	3.3	2.4

Zones of 4 $M_\odot$ core					
Zone	He	C	Ne	O	Si
T	0.27×10^9	0.65×10^9	1.4×10^9	1.5×10^9	...
"T"	...	1.4×10^9	2.0×10^9	3.0×10^9	...
ρ	8.2×10^2	2.2×10^4	1.0×10^5	6.0×10^6	...
η	...	≥ 0.001	0.0039–0.0047	$\geq$0.0039–0.0047	...
Mass (in 10^{33} g)	3.64	3.43	3.23	3.00	Small

when 3×10^{17} erg/g are added to the matter at the temperature, T, that was reached at the end of hydrostatic evolution. This is the amount of energy that is released by the nuclear reactions in the explosion (notice it is much less than is released by hydrogen burning, or even helium burning). The masses in the various zones of the models are also indicated. The conditions in the $8\,\mathfrak{M}_\odot$ core provide a good match to those required for the explosive processes. This is encouraging since, as we saw in the discussion of hydrostatic processes (Sec. III.C.2), this is the average core (by mass) of the range that can be expected to contribute to nucleosynthesis through supernova explosions. Conditions in the $4\,\mathfrak{M}_\odot$ core are in much less good agreement. This is the median core by number. Thus, if each supernova contributes an equal number of cosmic rays, the details of the heavy element cosmic ray abundances can be expected to differ from the solar values. In particular, an excess of the "high η" isotope ^{58}Fe might be expected (K. Hainebach, private communication, 1975).

Finally, it is necessary to combine the various explosive processes in some sensible proportions and compare the total product yields with Solar System abundances. Hainebach *et al.* (1974, and in NATO, 1974) considered the sum of equal parts of explosive carbon burning (ECB), explosive oxygen burning (EOB), equilibrium with normal freezeout (NSE1), and equilibrium with alpha-rich freezeout (NSE2). This model has 13 free parameters (four temperatures, four densities, $X = 1$, $\eta = 0.002$, and the proportions of the four processes). Solar System abundances of about 30 nuclides in the range $A = 28$–62 are correctly produced. The nuclides which are not made in this scheme include ^{58}Fe, ^{54}Cr, ^{36}S and the lower abundance, neutron-rich isotopes like ^{44,46,48}Ca, and ^{50}Ti. Woosley *et al.* (1973) present results for a combination of equal parts of EOB, explosive silicon burning (ESB), and NSE2. A slightly smaller number of parameters for this scheme reproduces 18 nuclides to within 25% and 11 more (all deficient) to within factors of 2 or 3. Arnett (in NATO, 1974) combines equal amounts of EOB, NSE1, and NSE2 with 1.5 times as much ECB with similar success. All of these schemes have assumed that the argon abundance is the high one of Cameron (1973). Woosley *et al.* (1973) have remarked that if the real abundance of the Si–S–Ar triad falls steeply, as suggested by the solar wind data (Crawford *et al.*, 1975), then it will be necessary to consider a different mix of the various processes. In particular, the combination of silicon-burning zones, with different amounts of silicon remaining in them, treated by Michaud and Fowler (1972) then becomes a more promising approach.

The composite shown in Fig. 11 is more complicated than any of these. It is made up of (1) the combination of equal parts of EOB, ESB, and NES2 considered by Woosley *et al.* (1973), which is normalized at ^{28}Si, (2) the products of carbon burning (Pardo *et al.*, 1974) coming from a zone with twice the mass of the EOB zone, 90% of which was processed with $T_0 = 2 \times 10^9$°K, and 10% with $T_0 = 2.35 \times 10^9$°K, (3) the minor products of carbon burning (Howard, 1973), which are indicated by a lower case "h," produced in an ECB zone with $T_0 = (2.0$–$2.3) \times 10^9$°K, and (4) the products of a high-neutron-excess NSE1 (Hainebach *et al.*, 1974), which are indicated by an upper case "H," produced in an NSE1 zone with $\eta \sim 0.08$, containing 10% as much mass as the NSE2 zone in the Woosley *et al.* composite. Notice that the abundances are given by number rather than by mass fraction. Given the mass fractions of the various zones just described as contributing, the normalization at ^{28}Si is the only one necessary, but there are, of course, numerous free parameters, including the temperatures, densities, neutron excesses, and expansion time scales of the various processes, and the amounts of the various zones. In this composite, reasonable amounts of all nuclides in the range ^{20}Ne to ^{64}Ni are made, except for ^{44}Ca, ^{50}V, and ^{21,22}Ne. It will be recalled from previous sections that ^{21,22}Ne, along with both the rare and common nuclides lighter than neon, are made adequately by other processes (H and He burning under hydrostatic and explosive conditions, spallation, and cosmological synthesis) and that ^{50}V can be made by spallation of ^{56}Fe. ^{44}Ca is made in ESB but is deficient by a factor of 30 or more and does not seem to be made in any other process. Some of the neutron-rich isotopes (^{36}S, ^{40}Ar, ^{46,48}Ca, etc.) may also be made in the *r* process discussed below.

The total effect of explosive nucleosynthesis is, then, to put the finishing touches on the nuclear abundance distribution which was produced by hydrostatic processes. The results of combining "reasonable" amounts of the products of several explosive processes bear a strong, but not perfect, resemblance to the cosmic abundances of Cameron (1968, 1973). It should be kept in mind that many of the Solar System element abundances are not known to better than factors of 3 in any case. Direct evidence for the occurrence of explosive nucleosynthesis in supernovae might be obtained in two ways. First we might directly observe enhancements of its products (Fe, S, etc.) in galactic supernova remnants. This has not happened, as we saw in Sec. II.D, though there is some evidence for enhanced Fe in the spectra of Type I supernovae themselves (Kirshner, 1975). Alternatively it may eventually prove possible to observe discrete gamma ray lines which are emitted when unstable nuclei made in the explosion decay. For instance, the ^{56}Fe is very probably made as ^{56}Ni, which decays (via ^{56}Co) in two stages having half-lives of 6.1 and 77 days, accompanied by the emission of gamma ray lines in the range 0.472–3.26 MeV. This probe for explosive nucleosynthesis has been explored extensively by Clayton (1973, and references therein).

E. Processes which build heavy elements and rare isotopes

As we have seen in the preceding sections, charged particle reactions under hydrostatic and explosive conditions are rather successful in accounting for the observed Solar System distribution of the nuclides up through about ^{64}Ni. Beyond that point, hydrostatic reaction cannot proceed. Because iron has the most tightly bound of all nuclei, further reactions must be endothermic. In addition, the Coulomb barriers become very high. Thus, as Howard (1973) remarks, nuclei beyond A = 65–70 are hardly affected, even by the extreme temperatures of explosive carbon burning. Both B²FH and Cameron (1957) detailed the need for several additional processes which could build the elements beyond the iron peak. Their thoughts were guided by strong indications of neutron shell effects in the abundances tabulated by Suess and Urey (1956). Some low abundance isotopes of lighter elements may also be built in these processes. Because

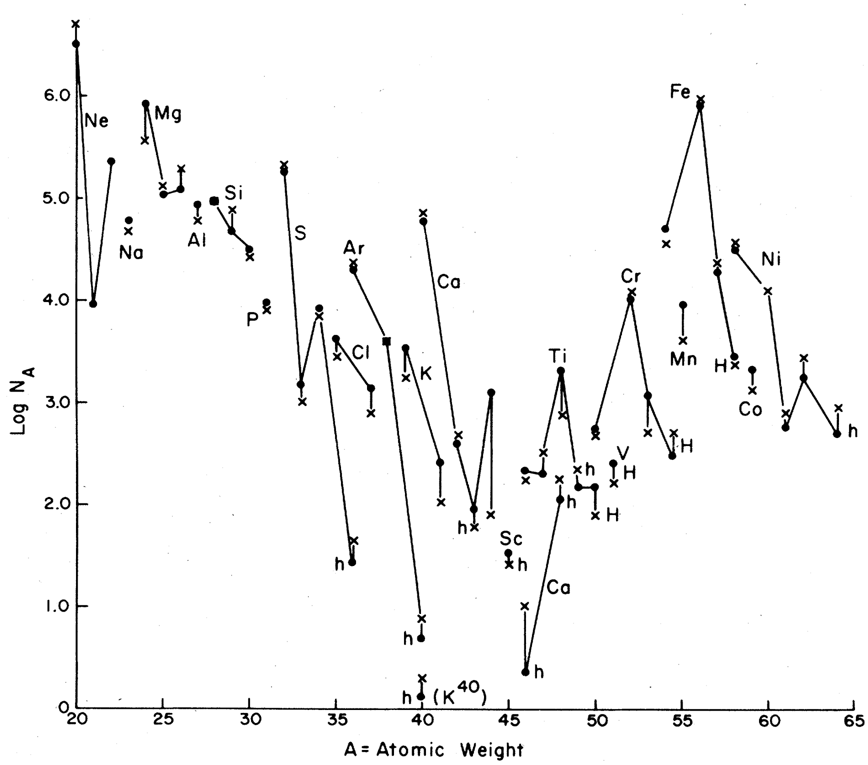

FIG. 11. A comparison of the products of explosive nucleosynthesis with the observed Solar System abundances in the range neon to nickel. All nuclides except ^{44}Ca (and ^{21,22}Ne which are made by helium burning) in this range are produced within a factor of 2–3 of their observed abundances by the composite of processes considered here. The observed abundances are uncertain by roughly this factor in any case. The composite includes products of explosive carbon, oxygen, and silicon burning and nuclear statistical equilibrium in proportions which are in good accord with the sizes of the relevant regions in an evolved stellar core. The proportions and the details of the conditions under which the reactions occur are discussed in the text (end of Sec. III.D.2). The lower case h's indicate the minor products of explosive carbon burning treated by Howard (1973) and the upper case H's the products of nuclear statistical equilibrium in a neutron-rich environment, as studied by Hainebach *et al.* (1974). The cluster of points representing silicon should be raised by 1.0 in logNA.

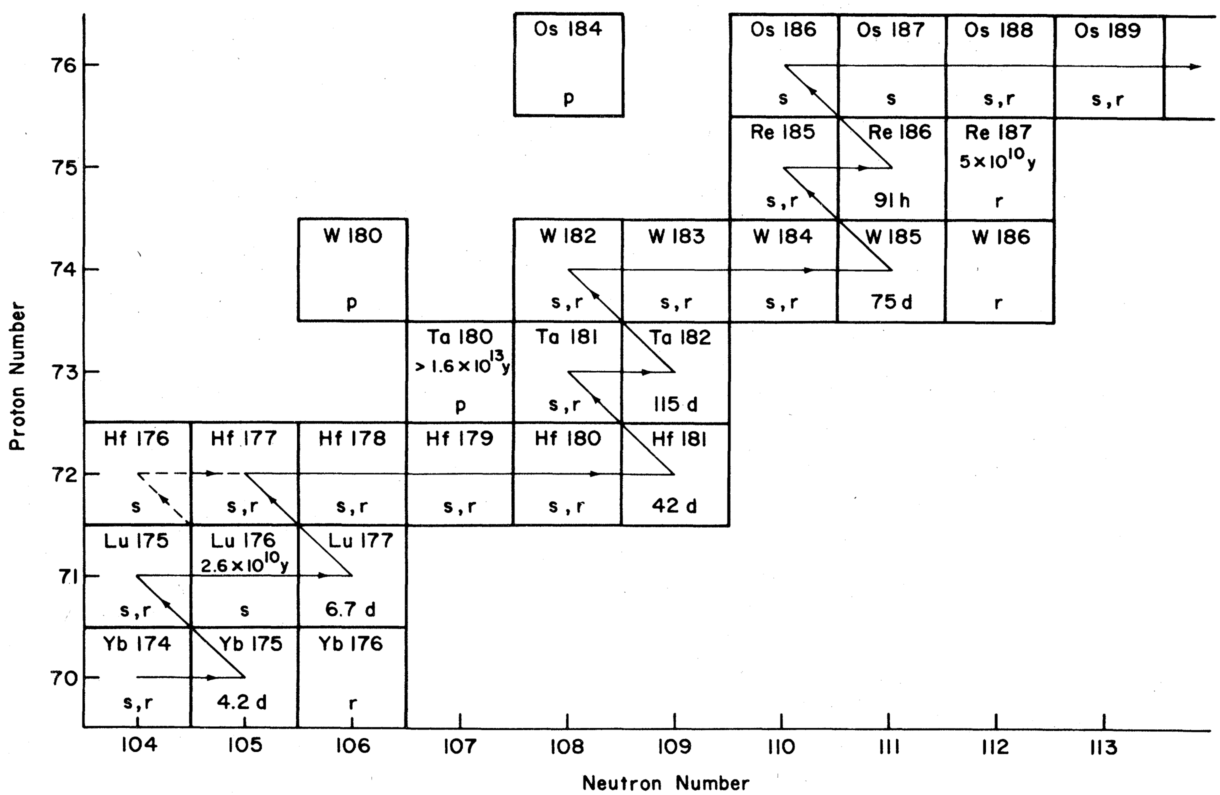

FIG. 12. The path of the *s* process (capture of neutrons on a slow time scale) through the region $A = 174$–189. All nuclides which lie along the solid line indicating the path of the process will have *s*-process contributions. Those to the right of the line can only be made by the *r* process (rapid addition of neutrons) and those to the left of it only by the *p* process (addition of protons or removal of neutrons from existing heavy nuclides). For each nuclide, its name, atomic weight, lifetime of the ground state (if it is beta unstable) and processes contributing to its abundance are given. There is a branch at ^{176}Lu which sometimes beta decays and sometimes captures another neutron, because it has a low-lying excited state with a short lifetime. Because the population of such excited states depends on the temperature and electron density where the synthesis occurs, the abundances of the products produced in such branches can be used to define the conditions under which *s*-process nucleosynthesis must have occurred.

most of the reactions involved will be endothermic, they will have to take place at rather high temperatures where there is lots of energy available. In addition, the high Coulomb barriers suggest that free neutrons will be the most promising building block for the heaviest elements. The short lifetime of free neutrons thus further constrains the sites of these processes. On the other hand, since all of the products are rather rare, it is not necessary for very much matter to have been through those sites. B^{2}FH originally identified three heavy-element building processes, called *s*, *r*, and *p*, while Cameron (1957) considered four, two analogous to *s* and *r*, and two analogous to *p*, in involving (p, γ) or (γ, n) reactions, but occurring hydrostatically and hydrodynamically.

1. The *s* process

Realization of the importance of neutron capture for heavy element building came soon after the discovery of the neutron (e.g., Gamow, 1935), and early work on cosmological nucleosynthesis made clear a general trend for abundances and neutron capture cross sections to be anticorrelated among the heavy nuclides, leading, for instance, to the large abundances of the nuclei with closed neutron shells (the "magic numbers" $N = 50$, 82, 126) relative to their neighbors, as one would expect for nuclides built by successive neutron captures.

The discovery of technetium (whose longest-lived isotope, ^{99}Tc, has a half-life of 2×10^5 yr) in S stars by Merrill (1952; other properties of S stars are discussed in Sec. II.C) made it clear that heavy-element synthesis must be actively occurring in them. Cameron (1954, 1955) and Greenstein (1954) realized that there was a plausible source of neutrons in the helium burning zone of moderately evolved stars, given by ^{13}C (α, n) ^{16}O, and neutron capture reactions could, therefore, occur in stars, even though they were not, in fact, important in the early universe for elements past helium or lithium. Because ^{13}C and other similar sources of neutrons [^{22}Ne (α, n) ^{25}Mg is the most important] release them only on the nuclear time scale of the star involved, a typical nucleus which captures a neutron will have ample time to beta decay (if it is unstable) before capturing another one. (The nuclear time scale of a star is the energy available from the particular reaction it is undergoing during a given evolutionary phase divided by its luminosity during that phase.) This stellar neutron capture process is, therefore, called "slow" or *s*.

Figure 12 shows a typical part of the chain of *s*-process captures and decays. It is reasonable to ask whether each nuclide in the chain can always make the decision "decay" or "stable" unambiguously, and, therefore, whether the chain is unique. A single iron-peak nucleus will capture less than 100 neutrons in building a typical s-process product, and the process occurs over a phase of the star's life determined by its nuclear time scale. Thus the interval between captures will be at least years. Nuclei whose lifetime against beta decay is of this order will then be branch points in the s-process chain. The ratio of the products from the two branches then tells us the time between captures rather precisely. This, in turn, helps to define the temperature, density, and neutron flux conditions under which s-process synthesis must have occurred. B²FH identified two such branch points which could be used to define time scales, at ^{79}Kr and ^{151}Sm. The two defined time scales for a single capture of 10^5 and 10 yr, respectively. B²FH interpreted this to mean that the time scale was set largely by the availability of neutrons, and that nuclides with $A \lesssim 100$ had been synthesized where neutrons were scarce and heavier nuclides where they were more readily available. This general conclusion was strengthened by the trends of abundance vs A. From the iron peak to about $A = 100$, abundances fall rapidly with increasing A, on the average, while beyond $A \sim 100$, the abundance distribution is rather flat. This is also consistent with quite a lot of iron-peak material having been exposed to a limited flux of neutrons, so that only lighter nuclides could be built, and a smaller amount of material having been exposed to a copious flux of neutrons, so that roughly equal amounts of all s-process nuclides were built. The s process terminates at $A = 210$ because the nucleus produced (^{210}Bi) is an alpha-emitter.

Within the framework of this general picture of the s process there are four major tasks. These are (a) to identify the nuclei produced by it, (b) to calculate the expected abundances of s-process nuclides as a function of the conditions under which the process occurs, (c) to compare the calculated values with observations and, thereby, determine the conditions under which the process which produced Solar System s-process material must have occurred, and (d) to find a suitable astronomical site or sites for the process.

The general principles behind identifying s-process products are clear from Fig. 12. All nuclides along the "valley of beta stability" will have s-process contributions. Nuclei to the left of the valley (more proton-rich nuclei) must be produced by some other process (called p and discussed below). Because the abundances of these nuclei are always much less than that of those in the "valley," the p-process contributions to nuclei also made by the s process are generally negligible. Nuclei to the right of the valley must also be made by some other process (called r and discussed below). Their abundances are not small, however, and r-process contributions are not negligible. The "pure" s nuclides are, then, those which are "shielded" by another stable nuclide from contributions due to the decay of very neutron-rich material. ^{176}Lu and ^{186}Os are examples of nuclides made only by the s process. ^{176}Yb and ^{186}W, on the other hand, lie to the right of the valley, and are made only by the r process, while ^{180}Hf and ^{184}W will have both s and r contributions. The assignments of nuclides heavier than $A = 75$ among the processes has changed hardly at all from the Appendix of B²FH (1957) to Cameron (1973), although B²FH assigned to the s process some lighter nuclides which we now attribute to explosive carbon, oxygen, and silicon burning. Clearly, the "pure-s" nuclides will provide the best test of theories of the process.

The next problem is to calculate the abundances of s-process nuclides that ought to be produced under various conditions. For any one nucleon, with atomic weight A, the time rate of change of its abundance, N_A, when exposed to a flux of neutrons, F_n, will be given by

$$dN_A/dt = F_n(\sigma_{A-1}N_{A-1} - \sigma_A N_A).$$

Given sufficient time and neutrons, then, s-process nuclides should all be built up to the same value of σN (where the cross section must be the one at the energy where the captures actually occur). Given insufficient time and neutrons, the product σN should gradually decline with distance from the "seed" nucleus. This signature of the s process was already apparent in B²FH. A detailed test requires both good cross sections and accurate abundances. Since isotope ratios for Solar System material are nearly always better determined than elemental abundances, the ideal test would be two or more s-only isotopes of a single element. Very few elements meet the requirement of having more than one pure-s isotope. ^{148}Sm and ^{150}Sm were the first to be tested, and careful measurements of the neutron capture cross sections near 30 keV by Macklin *et al.* (1963) established that the σN's were, indeed, the same to within the errors of the measurements. The results for ^{122,123,124}Te are similarly as expected (Macklin and Gibbons, 1967). Over wider ranges of A, the σN product invariably declines with A. This is true not only for Solar System material (Gibbons and Macklin, 1968), but also for a variety of other stars with a wide range of total metal abundance (Danziger, 1966). The fundamental correctness of the s process can, therefore, be regarded as well established.

Two other things must be known before detailed theory can be attempted. One is cross sections, and the other is the identity of the "seed" nuclei which capture the neutrons. Laboratory cross sections are not quite as hard to come by as they were for the explosive processes, since all the nuclides involved will be, by definition, stable. Most of the measurements have been made by the Oak Ridge group, and values of many of the necessary cross sections as a function of energy are given by Macklin and Gibbons (1965) and by Allen *et al.* (1971). Beer *et al.* (1974) have also measured some cross sections for $A = 50$–64. Woosley *et al.* (1975) have fitted a semi-empirical formula to the Oak Ridge data and have considerably advanced the art of estimating unmeasured cross sections. B²FH originally chose the iron-peak nuclei, especially ^{56}Fe, as their seed. The arguments in favor of the correctness of this choice have been summarized by Clayton (1968, and in NATO, 1974). Some of them are the following: (1) The neutron capture cross sections of lighter elements are so much smaller that it takes a larger flux of neutrons to convert, say, ^{28}Si to Ba than to convert ^{56}Fe to Pb. Thus, unless iron is exceedingly deficient in the material that undergoes s processing, it will inevitably dominate lighter seeds. (2) The heavier elements are so much less abundant than the iron peak that they, too, must be relatively unimportant. For instance, Ba stars are seen in which $[\mathrm{Ba/Fe}] \geq 3$;

thus the seed must, under ordinary circumstances, be at least 1000 times more abundant than Ba, in these stars. (3) For the Solar System material, as well, heavier seeds (e.g., r-process products) are unimportant. At most 1% of the Solar System material can ever have been in an s-process region, otherwise the ratios $^{37}Ar/^{36}Ar$, $^{41}K/^{40}Ca$, and the like would all be much larger than they are, since the numerators will be produced from the denominators by slow neutron capture. Thus, the s-process seed for Solar System material must be at least 100 times as abundant as the sum of its products. This limits us to the iron peak, within which ^{56}Fe is the most abundant nuclide.

Complete analyses of the s process have been given by Clayton *et al.* (1961) and Seeger *et al.* (1965). Cameron (1959a), Ulrich (1973), and others have also made important contributions to our understanding of the process. There are three questions we can ask the theory: how many neutrons has the average seed been exposed to, on what time scale, and under what temperature and density conditions? It is important to ask them in this order since we will need the first answer to tackle the last question. It was already clear in the work of B^{2}FH that no single neutron exposure would suffice to explain s-process abundances over the entire range of A attributable to the process. Rather, the situation required large amounts of iron to be exposed to a rather small number of neutrons per seed nucleus to produce the nuclides out to $A \sim 100$, and a smaller amount of seed material to be exposed to a large number of neutrons to produce the rather flat abundance distribution at larger A. Clayton (1968) has considered such a bimodal distribution of exposures and the circumstances under which it might occur (e.g., low exposures during helium burning and high exposures during carbon burning in a massive star). The predicted abundance curve falls quite smoothly from iron and then flattens off, as required.

The standard treatment of the s process (Clayton *et al.*, 1961; Seeger *et al.*, 1965), on the other hand, considers a continuous distribution of neutron exposures. The exposure is generally expressed as the time-integrated neutron flux

$$\tau = 10^{-27}\int_0^t v_T n_n(t)dt,$$

where $v_T = (2kT/m_n)^{\frac{1}{2}}$ and n_n is the neutron density. A brief digression is in order to explain that most of the relevant neutron capture cross sections scale is about the same way with temperature. Thus, only the total amount of processing that occurs depends on temperature, and not the abundance ratios produced by the s process. τ is, therefore, a suitable parameter. Now, the rate of change of individual abundances, N_A, is given by

$$dN_A/dt = \sigma_{A-1}N_{A-1} - \sigma_A N_A$$

when the thermally averaged cross sections are measured in millibarns (10^{-27} cm^2). If we now define the σN product per initial ^{56}Fe nucleus resulting from its exposure to a flux τ as

$$\psi_A = \sigma_A N_A(\tau)/N(^{56}\mathrm{Fe})_{\tau=0}$$

then the exact solution to the abundance distribution is given by (Bateman, 1910)

$$\psi_A = \sum_{i=56}^{A} C_{Ai}\exp(-\sigma_i\tau),$$

where

$$C_{Ai} = \frac{\sigma_{56}\sigma_{57}\sigma_{58}\cdots\sigma_A}{(\sigma_{A-1}-\sigma_i)(\sigma_{A-2}-\sigma_i)\cdots(\sigma_{56}-\sigma_i)}.$$

This is not easily evaluated, even with modern, high-speed computers, since the successive cross sections are often almost equal. An alternative, approximate solution was developed by Clayton *et al.* (1961) in which

$$\psi_A(\tau) \simeq \frac{\lambda_A(\lambda_A t)^{m_A-1}}{\Gamma(m_A)}\exp(-\lambda_A\tau),$$

where

$$\lambda_A = \sum_{i=56}^{A}\frac{1}{\sigma_i}\Big/\sum_{i=56}^{A}\frac{1}{\sigma_i^2}$$

and

$$m_A = \left(\sum_{i=56}^{A}\frac{1}{\sigma_i^2}\right)^2\Big/\sum_{i=56}^{A}\frac{1}{\sigma_i^2}.$$

Other methods of approximation and ways of handling the exact solution have recently been examined by Clayton and Newman (1974), Clayton and Ward (1974), and Ward *et al.* (1975). Within the framework of either the exact solution or one of the approximations, if seeds are exposed to a distribution of fluxes, $\rho(\tau)$, so that $\rho(\tau)d\tau$ is the number of iron nuclei exposed to fluxes between τ and $\tau + d\tau$, then the resultant abundances will be given by

$$\sigma_A N_A = \int_0^\infty \rho(\tau)\psi_A(\tau)d\tau.$$

Neutron capture cross sections tend to reach a minimum at each magic number (closed shell), and then rise gradually to the next magic number. This results in ψ_A being fairly flat between shell closures and dropping steeply at magic values of the neutron number. This "ledge and precipice" structure will naturally be reflected in the abundances, unless $\rho(\tau)$ is chosen so as to just counteract it. In fact, the observed σN curve for Solar System material seems to show some of this steplike behavior, suggesting a smooth $\rho(\tau)$.

Seeger *et al.* (1965) analyzed both exponential and power law forms for $\rho(\tau)$. Reasonable fits to the solar abundance curve of s-process nuclides could be obtained for either. If $\rho(\tau)$ has the form

$$\rho(\tau) = G\exp(-\tau/\tau_0),$$

then the equation for $\sigma_A N_A$ can be integrated to give

$$\sigma_A N_A \simeq G[\lambda_A\tau_0/(\lambda_A\tau_0 + 1)]^{m_A}.$$

The best fit to the observed Solar System σN curve is obtained for $G = 10^4$, $\tau_0 = 0.17$. A similar analysis has been done by Ulrich (1973, and references therein) in terms of a parameter Λ, which is the number of neutrons per seed nucleus made available in each of a succession of bursts of neutron production. It is essentially $1/\tau_0$ and arises naturally out of the inclusion of convection in his treatment. The best fit to Solar System abundances is obtained for $\Lambda = 4$–5, in good accord with the Seeger *et al.* (1965) value for τ_0. Figure 13 shows a comparison of Ulrich's calculated σN curve for $\Lambda = 4.8$ with Solar System abundances. The *s*-process elements in stars of low metal abundance appear to have abundance ratios consistent with larger values of τ_0 (less exposure) (e.g., Danziger, 1966).

Now that the theory has told us the total neutron exposure required, we can ask it the time scale for the *s* process, as determined by the ratios of the products from the branches that occur at beta-unstable nuclides with long lifetimes. For "slightly unstable" nuclei like this, the beta decay half-life will be a strong function of ambient temperature and electron density (especially when the decay occurs through an excited state). Thus, two "branch points" which define the same time scale can be used to pick out the temperature and electron density (separately) at which that part of the *s*-process synthesis occurred. Such a pair of branches occurs in the $A = 79$–82 region as follows:

$^{79}\mathrm{Se}(n,\gamma)\ {}^{80}\mathrm{Se}(n,\gamma)\ {}^{81}\mathrm{Se}(e^-\bar{\nu}_e)\ {}^{81}\mathrm{Br}(n,\gamma)\ {}^{82}\mathrm{Br}(e^-\bar{\nu}_e)\ {}^{82}\mathrm{Kr}$

or

$^{79}\mathrm{Se}(e^-\bar{\nu}_e)\ {}^{79}\mathrm{Br}(n,\gamma)\ {}^{80}\mathrm{Br}(e^+\nu_e)\ {}^{80}\mathrm{Se}$

or

$^{80}\mathrm{Br}(e^-,\nu_e)\ {}^{80}\mathrm{Se}$

or

$^{80}\mathrm{Br}(e^-\bar{\nu}_e)\ {}^{80}\mathrm{Kr}(n,\gamma)\ {}^{81}\mathrm{Kr}(e^-,\nu_e)\ {}^{81}\mathrm{Br}(n,\gamma)\ {}^{82}\mathrm{Br}(e^-\bar{\nu}_e)\ {}^{82}\mathrm{Kr}$

or

$^{81}\mathrm{Kr}(\eta,\gamma)\ {}^{82}\mathrm{Kr}.$

Both ^{79}Se and ^{81}Kr can either beta decay or capture another neutron, but the beta decay lifetimes, which are 6.5×10^4 and 2.1×10^5 yr in the laboratory are shortened to 3.2 and 10 yr, respectively, at a temperature of 5×10^{8}°K and an electron density of 10^{27} cm^{-3}. The situation is slightly complicated by ^{80}Br, which can emit an electron, emit a positron, or capture an electron, but these are known to happen 92%, 3%, and 5% of the time, respectively, more or less independent of density and temperature. The electron capture and position emission branches feed back into the $^{79}\mathrm{Se}(n,\gamma)$ chain at ^{80}Se. Of the five branches that separate at ^{79}Se and reconverge at ^{82}Kr, only one passes through ^{80}Kr. Thus, the observed Solar System ratio of these two "pure-*s*" nuclides, $^{80}\mathrm{Kr}/{}^{82}\mathrm{Kr} = 0.20$, tells us that the ^{79}Se must have beta decayed about a quarter of the time. A similar branch occurs at ^{85}Kr.

Requiring such a branch to lead to the right product ratios will define a permissible plane in a three-dimensional space of temperature, density, and time between successive neutron captures. Two such planes will typically intersect in a line, and three in a point (or not at all). Thus, by requiring several nearby (in A) branch points to determine *the same* values of the three parameters, relatively narrow ranges of temperature, density, and neutron flux can be found, within which the synthesis of Solar System material must have occurred. Ward *et al.* (1975) have applied this method to the ^{79}Se branches and find $T = 4$–6×10^{8}°K, $n_e = 10^{26}$–10^{27} cm^{-3}, and a neutron flux of 10^{15}–10^{16} cm$^{-2}\cdot$sec^{-1}. For a nucleus with a neutron capture cross section of 1000 mb, the capture time scale is then 10 yr. The sum of all the capture times from Fe up to Bi is about 10^4 yr under these circumstances. Thus, any site we want to postulate for the *s* process must be capable of supplying this sort of neutron flux for at least that long, under suitable density and temperature conditions. Blake and Schramm (1975, and in NATO, 1974) have found similar requirements on the neutron flux and time scale.

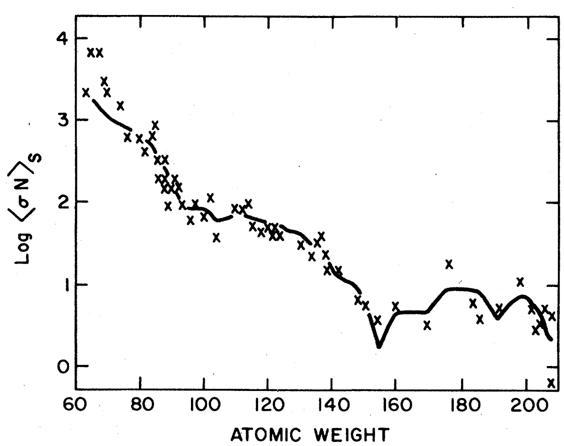

FIG. 13. A comparison of observed and calculated values of abundance times thermally averaged neutron capture cross section for nuclides produced primarily in the *s* process. The solid line is fitted to the calculations for individual nuclides by Ulrich (1973) and the crosses are the observed Solar System abundances (Cameron, 1973), corrected for *r*-process contributions as described by Seeger *et al.* (1965) times the measured cross sections of Allen *et al.* (1971), where these are available, or the interpolated ones used in Ulrich's (1973) calculations. The "ledge and precipice" structure of the calculated σN's is characteristic of a smooth distribution of neutron exposures during the processing. It does not conflict with the observations, but neither would a smoother falling-off of σN with atomic weight. The good agreement, as well as the rough constancy of σN over narrow ranges of A, is strong evidence that something like the *s* process has made a major contribution to Solar System abundances. The fluctuations in the range $A = 160$–200 come in the right places, but are not as large as the observed ones. Better calculations would probably improve the agreement (W. A. Fowler, private communication, 1975).

Ulrich (1973) has considered what happens if the nuclei undergoing the process are subjected to a burst of neutrons and then allowed to decay in peace for a while, before the next burst comes along. This allows the unstable nuclei to take longer to decay, without capturing another neutron in the meantime, so that the process can occur at rather lower temperatures and densities and on a longer time scale. His treatment corresponds to adding a single term, $-\Lambda_A N_A$ (attributed to the effects of convective mixing), to the right-hand side of the equation given above for dN_A/dt. Ulrich finds that 10-yr-long bursts, separated by intervals of several thousand years, with neutron flux of 4×10^{16} cm$^{-2}\cdot$sec^{-1} at a temperature of 2.5×10^{8}°K within the bursts, provides a good fit to the abundances of the nuclides produced through the ^{79}Se and ^{85}Kr branches.

These constraints on temperature and so forth apply *a priori* only to that fraction of the *s*-process material which had the right total neutron exposure, τ, to have produced most of the ^{80,82}Kr, and so forth. B^{2}FH originally found very different time scales from the branches at ^{79}Se and ^{151}Sm, but they did not allow for the effects of high temperature and density on the beta decay lifetimes. When these are included, the branching at ^{151}Sm and ^{154}Eu is consistent with the same conditions and time scales as the others (Ulrich, 1973). The flattening off of the abundance curve at $A \gtrsim 100$, which B^{2}FH originally attributed to a small amount of seed material having been exposed to higher neutron fluxes than average, we now tend to attribute to seed material having been exposed for a longer total time (either continuously or in bursts) to the same neutron flux.

Now that nuclear theory has told us the conditions under which *s*-process synthesis must occur, we can ask astronomy to provide a suitable site. The site or sites must be capable of providing something like an exponential distribution of total neutron exposure at the right temperature and on the right time scale. In addition, the products of the *s* process must be brought to the surface of the star in which they are produced, first so they can be seen (as in CH giants, Ba stars, and so forth, discussed in Sec. II.C), and second so that they can be returned to the interstellar medium without undergoing further nuclear reactions. To bring the processed material to the surface clearly requires convection or mixing. Mixing is also required to explain the neutron sources. The only reactions which can provide free neutrons at temperatures associated with hydrostatic phases of stellar evolution (required for the long time scales associated with the *s* process) are (α,n) reactions on ^{13}C, ^{17}O, ^{21}Ne, ^{22}Ne, ^{25}Mg, and ^{26}Mg. The abundances of these left behind after the completion of hydrogen burning by the CNO cycle are all small (most of the CNO is in the form of ^{14}N and so forth). Hence, by the time and in the place that helium burning starts so that high-energy alpha particles are available, there are no suitable targets around for them. Target nuclei for the (α,n) reactions must, then, somehow be brought into the helium burning region (which will also require mixing) or be produced *in situ*.

Cameron (1954) originally hoped that the equilibrium abundance of ^{13}C left behind by hydrogen burning might be an adequate neutron source, and that the *s* process could, therefore, go on during hydrostatic core helium burning. Then subsequent mixing could bring the products to the surface. The equilibrium amount is, in fact, very low. The primordial ^{13}C abundance in stars which have burned hydrogen by the p–p chain is also inconveniently low (^{13}C/^{56}Fe ~ 0.2, while we need a number of neutrons per seed nucleus). Hence ^{13}C cannot be an adequate neutron source unless it is continuously supplied by mixing down from a region in which hydrogen burning is still going on. This kind of mixing also tends to overcome the difficulty that ^{14}N will soak up all the neutrons via ^{14}N(n,p) ^{14}C before they can contribute to the *s* process if it is present in its equilibrium abundance (Cameron, 1955; B^{2}FH). ^{17}O, ^{21}Ne, and ^{25,26}Mg suffer from the same problem of low abundance under normal circumstances, coupled with the absence of a plausible "outside source" from which they might be brought into the helium burning region. Neutron production by ^{20}Ne(p,γ) ^{21}Na$(e^+\nu_e)$ ^{21}Ne(α,n) ^{24}Mg in regions with a high neon abundance was discussed by Marion and Fowler (1957) and more recently by Rolfs *et al.* (1975). Present stellar models do not provide a suitable site for the process. We will come back to models based on ^{13}C as the neutron source.

Cameron (1960) also pointed out that, once helium burning is well and truly under way, an additional neutron source becomes available via the series of reactions ^{14}N(α,γ) ^{18}F(e^+v) ^{18}O(α,γ) ^{22}Ne(α,n) ^{25}Mg. Temperatures high enough for this to be important can be expected to occur during helium core burning in very massive stars ($\gtrsim 9\,\mathfrak{M}_\odot$) and during helium shell burning in lower mass stars (~ 2–$8\,\mathfrak{M}_\odot$). Because CNO cycle hydrogen burning leaves behind it nearly all the CNO nuclei in the form of ^{14}N, this can provide a very copious neutron source. For Population I material, (C + N + O)/Fe is about 20, so that 20 neutrons per seed nucleus could be supplied in this way. We only need about five (Ulrich's Λ or the reciprocal of Clayton's τ_0), so there must be other neutron sinks if Solar System abundances are to result. With this neutron source, mixing is required only to bring the *s*-process products to the surface.

Stellar models for *s* processing based on ^{13}C neutrons build on an idea of Schwarschild and Härm (1967) and Sanders (1967). They were studying the phase of stellar evolution in which hydrogen and helium are burning in thin shells around an inert core. During shell burning, intermediate and low mass stars become somewhat unstable and undergo thermal pulses (discovered by Weigert, 1966), in which the rate of energy generation is greatly increased for a few years, and then dies down again into a quiescent phase lasting some thousands of years, followed by another thermal pulse, and so forth. The thermal pulses disturb the structure of the star a good deal. In particular, at the peak of any given pulse, a convective shell reaches up from the region of peak helium burning almost to the hydrogen–helium composition discontinuity which defines the base of the hydrogen-burning shell. If it were to reach all the way, then ^{12}C made by the helium burning further down could be exposed to protons to make ^{13}C and then taken back down to act as a neutron source.

One class of models of this type has been developed by Ulrich and Scalo (1972; Scalo and Ulrich, 1973; Ulrich, 1973). Their convective shell reaches just to, but not beyond, the hydrogen-rich region at the peak of each pulse. On each occasion, protons tunnel through the thin intervening radiative region between the base of the hydrogen-rich region and the outer edge of the convective zone. This makes the ^{13}C. Fresh seed nuclei are also brought down at the same time. Meanwhile, fresh ^{12}C and ^{4}He made in the shell sources and freshly produced *s*-process nuclei work their way back the other way into the hydrogen-rich zone. It can be shown (Ulrich, 1973) that this "add a little, take away a little" operation, when repeated many times, automatically yields the required exponential distribution of exposures. The processed material is finally brought to the surface because convective plumes driven by the $^{12}C(p,\gamma)\,^{13}N(e^{+}\nu)\,^{13}C$ reactions reach up to the base of the star's convective envelope. The chief difficulty with this sort of model is getting things to diffuse back and forth in exactly the right proportion. In addition, the existence of the plumes and their properties do not come out of any stellar model calculation, nor has the tendency of a molecular weight gradient to brake the convection been included. The problems are discussed by Iben (1974). The chief virtues of the model are that it yields the required exposure distribution in a natural way and that the properties of the stellar model cause the neutron exposure to occur on the right time scale automatically. The model is relevant for stars of 3–10 $\mathfrak{M}_\odot$.

A second class of ^{13}C models (Smith *et al.*, 1973; Sackmann *et al.*, 1974) requires the convective shell driven by the helium-burning pulse to merge with the convective envelope at pulse peak. The chief difficulty here is that no published model star has yet ever had such an unbroken convective zone. The virtues are, again, that the exponential exposure distribution and the time scales come naturally from the model. The models should apply to stars near 1 $\mathfrak{M}_\odot$.

A series of calculations by Sweigert (1975) emphasizes the fundamental difficulty in all models for low mass ($\lesssim$1–2 $\mathfrak{M}_\odot$) stars. The large pressure difference and composition gradient across the hydrogen burning shell act as a barrier against convection, convective overshoot, and diffusion coming from above or below the shell.

Models in which the neutrons are supplied by ^{22}Ne are again of two types. Peters (1968) and Couch *et al.* (1974) considered helium burning near the centers of rather massive stars (9 and 15 $\mathfrak{M}_\odot$, and $\gtrsim$15 $\mathfrak{M}_\odot$ respectively). Some *s*-process material is undoubtedly produced in this way, and the relative abundances in the region $A = 60$–90 look about right. There are several problems. The total neutron exposure is not large enough to produce the heavier *s*-process nuclides. In addition, very massive stars are rare enough that not enough material is likely to have been through their cores to yield the total amount of *s*-process products seen in the Solar System. Finally, it seems very difficult to get the products out of the star without subjecting them to further nuclear reactions, which would spoil the whole thing. In addition to the $A = 60$–90 nuclides, some ^{22}Ne and ^{25,26}Mg will clearly be made by this process. Neutron capture on lighter seeds may also give rise to ^{36}S, ^{40}Ar, ^{40}K, ^{43}Ca, ^{45}Sc, and ^{58}Fe in roughly their Solar System proportions. These are nuclides which are not easily made in the explosive processes discussed in Sec. III.D.

The other possibility, *s* processing during helium shell burning in intermediate mass stars, has been examined by Iben (1974). He followed the first 10 pulses of a 7 $\mathfrak{M}_\odot$ star undergoing shell hydrogen and helium burning. By the end of the tenth pulse, there was no question but that some *s*-process products were brought to the surface of the model. Reasonable extrapolations to later pulses indicate that amount should be adequate to account for observed *s*-process enhancements. The mechanism is slightly complicated. During an interpulse phase, the hydrogen-burning shell gradually works its way out, leaving behind a region containing 1%–2% ^{14}N. At the peak of the next pulse, a convective shell grows from the region of maximum helium burning out to just below the hydrogen–helium discontinuity. The convective shell, which contains about 0.002 $\mathfrak{M}_\odot$ of ^{12}C, ^{16}O, ^{4}He, and a bit of ^{22}Ne and heavier elements, picks up some fresh ^{14}N and some fresh seed nuclei and carries them down into the helium burning region. There the ^{14}N is turned into ^{22}Ne which acts as a neutron source for the newly arrived seed nuclei and for other partially processed *s* material that was already there. At the end of the pulse (which lasts about 30 yr), this region ceases to be convective. But, further out, the excess kinetic energy deposited by helium-burning causes matter within and beyond the helium burning region to expand. As the density drops, the ratio of gas pressure to radiation pressure decreases and the adiabatic temperature gradient drops, until, in regions successively further inward in mass, matter becomes unstable to convection. The convective envelope therefore grows gradually downward between pulses, and toward the end of the pulse sequence studied, it was reaching into matter which had previously been within the convective shell that included the helium burning region. Thus, after the pulse is over, the convective envelope can pick up freshly made ^{12}C, *s*-process material, and, probably, also small amounts of ^{16}O and ^{22}Ne from the helium burning and carry them to the surface. During this interpulse phase, the hydrogen-burning shell works its way still further out, leaving behind fresh ^{14}N again, and, when the next pulse comes along in about 2500 yr, the entire cycle repeats. The amount of enrichment per pulse will depend on the ratio of shell mass to envelope mass, but seems to be adequate. The model should apply to stars in the range $M = 2$–10 $\mathfrak{M}_\odot$. The model predicts definite relationships between *s*-process enhancement in giants and the $^{13}C/^{12}C$ and O/C ratios. The predictions do not obviously violate any of the observations. The model shares with other thermal pulse models the virtues of giving time scale and exponential distribution of exposures automatically. It has the special virtue that convection clearly carries the products to the surface. In addition, the average total exposure is uniquely predicted. Since ^{14}N (ultimately the neutron source) and ^{56}Fe (the seed) are brought down together and in a fixed ratio (for Pop I material), the average exposure depends only on the fraction of the neutrons that get captured by something else before the *s*-process chain finds them. ^{22}Ne and ^{25}Mg themselves both have rather large neutron capture cross sections. As a result, they and their immediate daughters soak up a good fraction of the 42 neutrons per ^{56}Fe that would be available from a complete conversion of CNO to ^{22}Ne, leaving only 4–5 neutrons per ^{56}Fe. This is just what is required to give $\Lambda = 4$–5 or $\tau_0 = 0.17$ since these two parameters are precisely the number of neutrons required per seed (Ulrich, 1973) and its reciprocal (Seeger *et al.*, 1965). Thus, the right average exposure is obtained and Solar System abundance ratios are guaranteed. An immediate consequence of this feature of the model is that the *s*-process progeny of ^{22}Ne as well as ^{56}Fe should be enhanced in stars where the process has occurred. The data of Table III can be compared with these predictions. The expected carbon excess and oxygen deficiency are clearly there in stars with enhancements of the heavy *s*-process elements. The ^{22}Ne products, Na, Mg, Si, Al, and P, do not display any clear trend. They are also generally not the elements most carefully studied in these stars, but an excess of Na has been reported in some S stars (Wallerstein, 1973). Further observations can obviously help to test this facet of the model.

The average neutron exposure obtained will depend somewhat on the temperature at which the reactions take place, since the ^{22}Ne and ^{25}Mg neutron capture cross sections have a different temperature dependence from those of ^{56}Fe and its daughters. The right average exposure is obtained for $T = (3\text{–}4) \times 10^{8}\,^{\circ}$K. This is in good accord both with the temperature in the helium burning region of stars with two shell sources and with the temperature required to make the time scales at the ^{79}Se and ^{85}Kr branch points come out right.

Any of these models in which a convective envelope reaches down into the hydrogen-burning zone also predicts a temporary phase of surface Li enrichment. The temperature at the base of the convective envelope gradually increases as the star evolves. When the temperature reaches about $4 \times 10^{7}\,^{\circ}$K, ^{7}Li is formed by $^{3}\mathrm{He}(\alpha,\gamma)\ ^{7}\mathrm{Be}(e^{+}\upsilon)\ ^{7}\mathrm{Li}$ in the envelope and can be brought to the surface before it is destroyed by protons. By the time the temperature at the base of the convective envelope reaches about $6 \times 10^{7}\,^{\circ}$K, the lithium will be destroyed before it can reach the surface. Thus, high surface lithium abundance on the surface of giants should occur during a transient phase in their evolution and be correlated with intermediate degrees of enhancement of carbon, while Li may reach the surface either ahead of or with *s*-process material.

Another coherent scheme which relates *s* processing, lithium enhancement, and variations in C:N:O ratios to various evolutionary stages in giants has been formulated by Wallerstein (1973).

In conclusion, then, the *s* process undoubtedly occurs. We see its signatures ($\sigma N \sim$ const and abundance peaks at the nuclei with "magic" $N = 50$, 82, and 126) in Solar System and other material. The conditions under which it must occur are quite well defined by the requirement that observed abundance ratios (especially the products formed in branches of the chain) be matched. The most difficult part is to find a suitable site for the process. At the present time, the phase of helium shell burning in intermediate mass stars in which thermal pulses occur seems to be the most promising, but many of the details of the models and their comparisons with observations remain to be worked out. The *s*-process material produced in these stars can be returned to the interstellar medium by stellar winds, planetary nebulae, and (presumably for the most massive of the stars) supernova explosions. Since the lithium is probably destroyed before hydrostatic evolution is complete, it can only be returned by stellar winds.

2. The *r* process

Suess and Urey (1956) originally called attention to the double peak structure in the curve of abundance vs atomic weight at $A = 130$–138 and 194–208 (see Fig. 2). They associated the heavier peak in each case with the magic neutron numbers $N = 82$ and 126 which occur at those A's and suggested that the lighter A peaks could reflect neutron capture nucleosynthesis under conditions of unusually high neutron density. The peaks would then have been formed at the magic N's and later beta decayed back to a smaller value of $N - Z$. We have seen in the previous section that the peaks at the magic N's are a result of neutron capture on a slow time scale (the *s* process), such that unstable nuclei can generally beta decay before another neutron is captured. B^{2}FH (1957) called the process that made the lower A peaks "rapid" or *r*, because neutrons were captured on such a rapid time scale that beta decays did not occur between captures, except when the next additional neutron would not be bound to the nucleus without an intervening decay. Such a process clearly forms nuclides well on the neutron-rich side of the valley of beta stability, which will subsequently decay to give the $A = 130$ and 194 peaks. A third pair of *s* and *r* peaks associated with the magic neutron number $N = 50$ occurs near $A = 80$–88. It is hard to see in Fig. 2(a) because of the general fall-off of abundances with distance from ^{56}Fe, but is clearer in Fig. 2(b), where the processes are separated, and in the *r*-process abundances shown by Seeger *et al.* (1965). In order to prevent most unstable nuclei from decaying and yet allow a few decays where no further capture would otherwise be possible, the time scale of rapid neutron capture must be a few seconds. The process will be terminated by fission. That is, the heaviest nucleus that can be built by the *r* process is the first one for which the lifetime against spontaneous or neutron induced fission is shorter than the beta decay lifetime. We will come back to the question of where this cutoff occurs.

Within the framework of this general picture of the *r* process, we have the same four tasks as in the case of the *s* process, that is, (a) to identify the nuclei produced by it, (b) to calculate the expected abundances of *r*-process nuclides as a function of the conditions under which the process occurred, (c) to compare the calculated values with observations and, thereby, determine the conditions under which the production of Solar System *r*-process material must have occurred, and (d) to find a suitable astronomical environment for the process. Finally, we will ask about the production of some particularly interesting nuclei, those used for nucleocosmochronology and the superheavies.

As can easily be seen from Fig. 12, there will be *r*-process contributions to the abundance of all nuclides which are not shielded from the decay of neutron-rich material by another (higher Z, same A) nuclide. Some of these nuclides, which lie on the line of beta stability (like ^{180}Hf and ^{184}W), will also have s-process contributions; others (like ^{176}Yb and ^{186}W) will be pure r. The r-only nuclei will provide the easiest test of theories of the process, but we can also use many that are made by both processes. Because the σN curve (product of abundance and neutron capture cross section) is smooth (at least over small ranges of A) for the s-process products, we can fit such a curve through the "pure-s" nuclides and use it to estimate the s-process contribution to nuclei made by both processes. The remainder of the abundance is then to be attributed to the r process. Such corrections can be made quite accurately, provided that the s-process contribution is not much more than half the total abundance (Seeger *et al.*, 1965; Allen *et al.*, 1971). The main features of the r-process abundance curve are the three peaks just mentioned, a broad bulge at $A = 150$–170 (believed to be due to a spheroidal deformation of the nuclei in this mass range which increases their binding energy, and, therefore, their relative production), slight distortions at the proton magic numbers $Z = 28$, 50, and 82 (due to extra long lifetime against beta decay of a nucleus which is about to decay to one which has a single extra proton outside a closed shell), and (perhaps) slight excesses near $A = 110$ and 145 due to the recycling, by fission, of nuclei built up to $A \sim 260$. (The sum of the A's at the two peaks is not equal to the A that fragments because some free neutrons are released.) Because the s process is terminated at $A = 209$ by alpha emission, all nuclei heavier than this (including uranium and thorium and the postulated "superheavies" with $Z \sim 106$–115) must be pure r products. Typical r-process calculations which reproduce the Solar System (meteoritic) abundances at the three peaks make at least twice as much Th (and U) as the meteoritic abundances, and are, therefore, more consistent with the photospheric abundance of Th (Seeger *et al.*, 1965).

As in the case of the explosive reactions and the s process, r-process calculations are generally done in parametrized form. The necessary parameters are temperature, total density, neutron density, and something about the time scale over which these are maintained. The process can be treated as either an equilibrium or an explosive one. The nuclear physics data necessary to calculate r-process abundances include nuclear masses (which determine whether the next neutron to be added will be bound or not), beta decay lifetimes (which determine how long a "neutron-saturated" nuclide will have to wait before it can capture again), and neutron capture cross sections (these are important only toward the end of the process as the system falls out of equilibrium). Because the nuclei under consideration are far from the region of beta stability, adequate laboratory data are generally not available. Thus the masses are normally estimated from the semi-empirical nuclear mass formula and the cross sections from nuclear models. The techniques for doing this have gradually increased in accuracy and sophistication (compare B^{2}FH; Becker and Fowler, 1959; Seeger *et al.*, 1965; Schramm, 1973a; and Woosley *et al.*, 1975).

In the most general case, the time rate of change of the abundance of an r-process nuclide will be given by

$$\begin{aligned} dn(N,Z)/dt = &-\lambda_{n\gamma}(N,Z)n(N,Z) \\ &+\lambda_{\gamma n}(N+1,Z)n(N+1,Z) \\ &+\lambda_{n\gamma}(N-1,Z)n(N-1,Z) \\ &-\lambda_{\gamma n}(N,Z)n(N,Z) \\ &+\lambda_{\beta}(N+1,Z-1)n(N+1,Z-1) \\ &-\lambda_{\beta}(N,Z)n(N,Z) \end{aligned}$$

where $\lambda_{n\gamma}$ is the rate of neutron capture, $\lambda_{\gamma n}$ is the rate of photoneutron emission, and λ_β is the beta decay rate. Most of the time, the (n,γ) and (γ,n) reactions will be in equilibrium over the range of nuclides with a particular value of Z. In addition, for this set of nuclei the neutron captures and emissions will be so rapid that beta decays are unimportant. Hence, for these nuclei, a statistical balance will be set up in which (neglecting statistical weight factors)

$$\log\frac{n(N+1,Z)}{n(n,Z)} = \log n_n - 34.07 - 1.5\log T_9 + \frac{5.04}{T_9}Q_n,$$

where n_n is the neutron density, T_9 is the temperature in units of 10^9°K, and Q_n is the energy of binding of the next neutron to the nucleus (N,Z) in MeV, i.e.,

$$Q_n = c^2[M(N,Z) + M_n - M(N+1,Z)].$$

This equilibrium will hold up to the first nucleus in which the next neutron is not bound at a particular temperature and density, that is, as long as

$$Q_n \geq (T_9/5.04)(34.07 + 1.5\log T_9 - \log n_n) = 2\text{ MeV}$$

at $T_9 = 1$ and $n_n = 10^{24}$.

The flow between nuclides with adjacent Z will be given by

$$dn(Z)/dt = -\lambda_\beta(Z)n(Z) + \lambda_\beta(Z-1)\,n(Z-1),$$

and, in equilibrium,

$$\lambda_\beta(Z)\,n(Z) = \lambda_\beta(Z-1)\,n(Z-1).$$

Under these circumstances, large abundances will be built up at "waiting points," the A's where a beta decay must occur before capture can continue. These large abundances will be reflected in the final abundance curve (after beta decay back to the region of stability) in the form of great excesses of the "favored" A's over their neighbors. No such great lumpiness is observed in Solar System r-process abundances. B^{2}FH assumed that it would be smoothed out by final neutron captures after the system had frozen out of the equilibrium. They introduced an arbitrary smoothing function over the intervals between "waiting points" in their calculations to allow for this. Under these circumstances, the path of r-process synthesis is uniquely determined (for given T and n_n) by the criterion on Q_n, and within that path, the relative abundance of each nuclide is given by

$$n(A) \propto \frac{1/\lambda_\beta(N_1Z)}{\Delta A},$$

where $\lambda_\beta(N,Z)$ is the beta decay rate at the "waiting point" just below the nuclide of interest, and ΔA is the interval in atomic weight to the next waiting point.

The alternative to this semi-equilibrium type of calculation is a dynamical or explosive r process in which temperature and density decrease smoothly from some maximum value and the relative abundances of the various nuclides are followed through the freezing out process. Calculations of the dynamical r process have been attempted by Cameron *et al.* (1970), Sato *et al.* (1973), and Schramm (1973a,b).

The first constraint which the observed abundance curve obviously imposes on the theory is that the "magic" N's must be reached at the values of A where the observed r-process peaks are. This partially constrains the path of the r process and so determines the minimum value of Q_n that must be bound in the vicinity of each of the three peaks. B²FH found that slightly different values of T_9 were necessary at fixed n_n (or different values of n_n at a fixed value of T_9) to match the positions of the three different peaks. For example, if $n_n = 10^{24}$, then $T_9 = 1.45$, 1.0, and 0.8 at the $N = 50$, 82, and 126 peaks, or, if $T_9 = 1$, then $n_n = 8 \times 10^{19}$, 3×10^{23}, and 1×10^{26} cm^{-3} at the three peaks. By adjusting a final free parameter (the excitation energy of the nucleus left behind after the beta decays, as a function of A) they were able to match the relative heights of the three peaks and, therefore, the over-all shape of the $n(A)$ curve for r-process nuclei, rather well. Such calculations correspond to the case where equilibrium among the various processes has just been reached, but significant fission has not yet occurred. Since the average atomic weight of the r-process nuclides in Solar System material is about 94, this requires about 38 neutrons per ^{56}Fe seed. The total time required is the sum of beta decay lifetimes at the various "waiting points," which B²FH found to be about 80 sec. If more neutrons and time were allowed, significant buildup into the unstable, high A region would have occurred, which, recycling back to $A \sim 110$ and 150 by fission, would continuously raise the relative abundance beyond this point until all the material was contained in high A nuclides. Under these circumstances large amounts of ^{254}Cf could be expected, whose subsequent spontaneous fission (lifetime = 56 days) B²FH suggested as an important contribution to the light output from supernovae of Type I (whose light curves decay roughly exponentially with a comparable lifetime). This then suggested that the light and heavy r-process elements were built in two different environments, presumably Type II and Type I supernovae, respectively. This would account both for the difference in number of neutrons per ^{56}Fe seed and for the slightly different temperatures required.

Subsequent equilibrium calculations of r-process synthesis (Seeger *et al.*, 1965, 1968) with improved nuclear parameters confirmed and strengthened the two most discouraging of the above conclusions: the need for artificial smoothing and the need for more than one environment. In the matter of smoothing, neither the superposition of several regions at slightly different temperatures nor the extra neutron captures that occur during the decay back to beta stability were anywhere near adequate (Seeger *et al.*, 1965, and Blake and Schramm, 1973, respectively). The environment problem can be most easily discussed in terms of two time scale parameters. The first is the total time over which the process is allowed to act (the exposure time). The second is the "cycling time" over which a nucleus is built up from $A \sim 110$ to $A \sim 260$ and comes back again by fission. This latter is small at high neutron densities (obviously) and low temperature (because gamma rays do less disrupting) and large at low densities and high temperatures. Clearly, if the exposure time is much less than the cycling time, little material will reach the high A region; while if the exposure time is much greater than the cycling time, most of it will end up there. The two time scales were about equal in the B²FH treatment, so that some matter reached the highest A's, but little recycling occurred. Their requirement for a lower temperature (or higher density) to make the $A = 198$ peak than was required for the other peaks was, then, equivalent to requiring a shorter cycle time for the material that made this peak. Improved nuclear physics (especially the mass formula) has emphasized the dichotomy: by the time any material has reached the $A = 198$ peak, the $A = 80$ peak has eroded away completely. Seeger *et al.* (1965, 1968) found that the first two peaks could be made together with an exposure time of about 4 and a cycle time of about 400 sec, while, with the same exposure time, the third peak required a cycle time of only 3 sec. While such a multiplicity of kinds of environments is by no means impossible as a property of the real world, the approximate constancy of the relative abundances of the heavy elements in a wide variety of unevolved stars (with very different total metal abundances) argues that the environments must either have always occurred together or otherwise have been well mixed.

In any case, all of the astronomical sites which have been suggested for the r process (supermassive objects, supernovae, tidal disruption of neutron stars, etc.) are more or less explosive in nature. This suggests that a proper r-process theory must be a dynamical one. It can be regarded as a bonus that the inclusion of the effects of changing temperature and density and freeze out in an explosion turns out to solve the problems of smoothing and of multiplicity of environments.

Schramm (1973a,b) has done the most extensive calculations of the dynamical r process to date. He considers an explosion in which density and temperature vary with time exponentially, on something like the dynamical time scale, that is,

$$\rho = \rho_0 \exp(-t/\tau_X), \quad T_9 = (T_9)_0 \exp(-t/3\tau_X),$$
$$\tau_X = \chi\, 446/\rho_0^{\frac{1}{2}} \text{ sec}.$$

Then choices of ρ_0, $(T_9)_0$, χ, the initial abundance of seed nuclei, $n_0(N,Z)$, and the initial fraction of the mass which is in the form of free neutrons, $(f_n)_0$, suffice to determine the subsequent behavior of the system uniquely. The differential equations for changes of abundances with time are converted to difference equations and solved numerically. Convergence can generally be achieved in a couple of iterations at each time step. Freeze out normally occurs because the temperature drops to the point that (γ,n) reactions produce no more free neutrons and the available ones have all been captured. A neutron-rich freezeout could also occur if the expansion time scale got to be faster than the neutron capture time scale. Of the various nuclear data required, nuclear masses and neutron capture cross sections appear to be adequately known, but the beta decay rates far from the region of nuclear stability are not. Several different ways of calculating them were therefore tried. The fundamental conclusions are not dependent on which of the ways was used.

Because of the large number of initial conditions to be chosen, it is not practical to explore all possible ranges of all of them to find a "best" solution. It is already clear that such a solution would not, in any case, be unique. All the calculations which approximately reproduced Solar System *r*-process abundances did, however, start from very high temperatures ($T_9 = 1$–4), rather high densities ($\sim 10^{5-8}$), and with a considerable fraction of the matter (10%–50%) in the form of free neutrons. The seed nuclei could be either the equilibrium nucleus under particular high density, high n/p conditions (e.g., ^{78}Ni), or alpha particles and their reaction products. The time scale parameter, χ, could always be taken as unity, and the explosion time scales were, therefore, very short (0.01–1.0 sec). If the initial n/p ratio is rather low, the neutrons may be exhausted before any fissioning takes place. Under these circumstances, the actinide region will not be populated uniformly, adding additional uncertainty to the ratio of U to Th produced. The calculated production ratio for these two (pure-*r*) elements is a necessary input for the considerations of nucleocosmochronology discussed in Sec. II.B.2 and below.

It is clear that the initial conditions for the *r* process are not well enough determined for us to pick out a unique site for the process from them alone. On the other hand, it is possible to choose the particular initial conditions that would be appropriate to a given site, and then see if the products bear any relation to the Solar System *r*-process abundance curve. It should, in principle, be possible at least to restrict the kinds of probable sites this way. Some of the sites which have been tested in this way are (a) novae, (b) neutron stars in the process of tidal disruption, (c) supermassive stars, and (d) supernovae with properties given by a variety of models. The details of the composition have been calculated only in the latter two cases.

Hoyle and Clayton (1974) have pointed out that the ^{13}C, whose production acts as the major source of luminosity in the nova models of Starrfield *et al.* (1974, and references therein), can act as a neutron source. In order for many neutrons to be released by $^{13}C(\alpha,n)\,^{16}O$ reactions, the white dwarf which gives rise to the nova must have a larger total mass (i.e., smaller radius) and a larger fraction of its mass in the nuclear burning surface region than is the case for a typical nova event. With a radius of 4×10^8 cm and $0.01\,\mathfrak{M}_\odot$ in the exploding region, an initial temperature of about 4×10^{8}°K is achieved and about 300 neutrons per seed nucleus are provided on a time scale of a few seconds. Two events per year of this type would have supplied $10^5\,\mathfrak{M}_\odot$ of *r*-process material over the age of the Galaxy, intermediate between the amounts of light and heavy *r*-process elements which we actually observe. The composition of the products has not been calculated, but the relatively low initial temperature and long time scale suggest that the products will be too heavily concentrated toward high A values for this mechanism to have supplied much of the Solar System *r*-process material.

Lattimer and Schramm (1974) have considered the process in which a neutron star in orbit around a black hole gradually spirals in closer due to the emission of gravitational radiation. When the neutron star reaches the Roche limit, it is tidally disrupted, and about $0.05\,\mathfrak{M}_\odot$ of material is expelled from the system completely by the "tube of toothpaste" effect, provided that the mass of the black hole is less than about $20\,\mathfrak{M}_\odot$. The neutron density in the extruded matter is obviously very high, and they consider it possible that rather efficient synthesis of *r*-process nuclides might occur. If all the matter is processed, then about one event every 2000 yr will have supplied the observed amount of *r*-process material over the history of the Galaxy. They do not regard this as improbable, given the observed statistics of binary star masses and separations.

Hoyle and Fowler (1963) first suggested that the site of the *r* process might be associated with the explosive outbursts of supermassive stars ($M > 10^4\,\mathfrak{M}_\odot$). Seeger *et al.* (1965, 1968) therefore sought the two kinds of environment their *r*-process calculations required among the models of supermassive star explosions by Fowler and Hoyle (1964). Since the explosion is essentially the reverse of a free fall, the necessary exposure time of 4 sec immediately yields the neutron density in the region where synthesis must occur. It is about 5×10^{26} cm^{-3}. Then, in order to get a cycle time of about 400 sec, the temperature in this region has to be about 2.4×10^{9}°K. According to the Fowler and Hoyle (1964) models this temperature implies objects of about $10^5\,\mathfrak{M}_\odot$ as the site for synthesis of the first two *r*-process peaks. As the explosion proceeds, the conditions of lower temperature ($T_9 = 1$) and neutron density ($n_n \sim 3 \times 10^{25}$ cm^{-3}) needed to produce the third, high A, peak will also be achieved. Models of supermassive stars which have been calculated more recently (e.g., Appenzeller and Tscharnuter, 1973, and references therein) do not undergo the relaxation oscillations of the Fowler and Hoyle (1964) models before exploding. Their temperatures, densities, and expansion time scales are therefore rather different, and it is not clear whether the necessary conditions can be achieved. For instance, the model which Audouze and Fricke (1973) found to produce interesting amounts of the light nuclides ^{13}C, ^{15}N, etc. (see Sec. III.C.1), achieved a maximum temperature of (3–4) $\times 10^{8}$°K at a density of about 6 g cm^{-3} over a time of 10^{3-4} sec. No dynamical *r*-process calculations suitable to supermassive star conditions appear to have been carried out. The advantages of supermassive objects as a site for *r*-process nucleosynthesis were discussed by Hoyle and Fowler (1963). There seem to be two chief disadvantages. First, it appears that stars with $M \sim 10^5\,\mathfrak{M}_\odot$ only explode if their initial metal abundance is greater than about 1%. Thus *r*-process material could only have begun to be returned to the interstellar medium by supermassive stars after the metal content of the regions of the Galaxy where they form had already reached nearly its present value. Stars of low total metal abundance should, then, have little or no r-process material. This does not seem to be the case. Second, there is no direct observational evidence for the existence of supermassive stars anywhere.

B^{2}FH originally believed that conditions suitable for r-process nucleosynthesis would be achieved in supernova explosions. This still seems to be the most promising site. The useful feature of supernovae is that the onset of dynamical instability at the center of a massive star when it has built up a core containing approximately the Chandrasekhar mass of inert material (mostly Fe, Ni, etc.) is exceedingly rapid. The actual onset of instability can occur in a variety of ways, including the endothermic photodisintegration of ^{56}Fe to give 13 alpha particles and 4 neutrons (as envisioned by B^{2}FH), the formation of electron–positron pairs, or electron capture by protons in nuclei (e.g., Cameron, 1959; Chiu 1964; Finzi and Wolf, 1967; Fraley, 1968). Which of these occurs can be very important to the star. The reader who wishes to explore the possibilities is referred to the following set of original model calculations: Arnett (1967), Colgate and White (1966), Leblanc and Wilson (1970), Wheeler *et al.* (1973), and Wilson (1974), and to the proceedings of any of the numerous recent conferences on supernovae, high-energy astrophysics, neutrino astronomy, gamma ray bursts, and so forth. Unfortunately, the r process is presently unable to distinguish the various kinds of supernovae. Or, alternatively, suitable initial conditions for the dynamic r process occur in all the models that have been studied (Schramm, 1973a,b). In particular, it does not matter whether the seed nuclei are initially present as, e.g., ^{78}Ni or are built up out of free alpha particles early in the explosion. In either case, suitable combinations of temperature, density, and neutron mass fraction can be found to give approximate agreement with solar r-process abundances. A relatively low initial density (10^5 g cm^{-3}) facilitates forming all three r-process peaks in a single explosion.

Two of the questions which the theory of the r process has often been asked recently are: Do you make superheavies? And what is the production ratio of the radioactive nuclei important for cosmochronology? These questions can be tackled independent of the site of the process by requiring the synthesis to occur under circumstances the produce Solar System abundances of the three r-process peaks.

B^{2}FH originally believed that fission would terminate the r process near $A = 260$. As a result, cycling through that region could build up large amounts of ^{254}Cf, which, in turn, could be responsible for the light output from supernovae. The standard r-process calculations of Seeger *et al.* (1965, 1968) also had fission terminating the r-process chain before the closed neutron shell at $N = 184$ was reached. Two things inspired a reconsideration of the situation. The first was the prediction (Myers and Swiatecki, 1966; Sobichewski *et al.*, 1966) that the next magic proton number (after $Z = 82$, lead) would occur, not at $Z = 126$ by analogy with the neutron shells, but at $Z = 114$. This made the superheavies seem more reachable and inspired a calculation by Viola (1969) of the termination of the r process by fission. He concluded that fission would cut off the synthesis at about $A = 275$. Almost simultaneously Fowler *et al.* (1969) announced that a balloon-borne cosmic ray detector had found a single nucleus with charge apparently equal to about 106. This inspired Berlovich and Novikov (1969) to recalculate r-process synthesis and conclude that superheavies could be formed. Schramm and Fowler (1971) were similarly optimistic. These early calculations did not attack the fissionability of the nuclides along the r-process path with the heaviest guns available in the arsenal of nuclear physics.

There are really four separate questions: (a) Is there an "island of stability" near $Z = 114$, $N = 184$? (b) Do the nuclei there have long enough lifetimes to be interesting for astronomy? (c) Can the r process reach the island if it exists? (d) Have any superheavies been observed? The theory of nuclear stability has been reviewed by Fix (1972). The existence of a region of relative stability of superheavy elements seems clear. The lifetimes are, however, very uncertain. Fiset and Nix (1972) find that reasonable calculations can give lifetimes that range from less than a year to millions of years. The r-process theorists are also in disagreement. Schramm and Fiset (1973) conclude that whether or not the island of stability is reached depends on details of the properties of heavy nuclei (especially pairing and spherical deformation) which are not adequately known at present. Johns and Reeves (1973), on the other hand, conclude that superheavies may well be made in those supernovae where the neutron density is rather low, because of the reduced probability of neutron-induced fission. Howard (in NATO, 1974) believes that the "island" is completely surrounded by "straits" in which the barrier against fission is no larger than the neutron separation energy, and, thus, that no r-process synthesis is likely to produce superheavies. These calculations have all used the "best available nuclear theory."

The problem of the existence and formation of the superheavies must, therefore, be handed over to the observers and experimenters. Some of the evidence has been discussed in Secs. II.E and II.B.2. No further superheavy cosmic rays have been found, although the detectors on Skylab greatly increased the amount of collecting time (Price and Shirk, 1974). It is possible that the $Z = 106$ event reflected a miscalibration of the tracks in nuclear emulsions (Price *et al.*, 1971). At present, it is possible to set only an upper limit, $N(\text{superheavies})/N(Z = 70\text{–}83) \leq 0.006$ (Price, in NATO, 1974). The other possible evidence for superheavy elements with astronomical long lifetimes is the anomalous Xe component in meteorites, whose isotope ratios are neither normal, due to the decay of ^{129}I, nor attributable to the fission of ^{244}Pu. This component has been attributed to the fission of a superheavy volatile element near $Z = 114$ by Anders and Heymann (1969). The astronomical case for stable superheavies, must, therefore, be judged "not proven." Further cosmic ray data may eventually clarify the situation. The question of lifetimes can be most effectively settled in the laboratory through production of superheavies by heavy-ion reactions (^{76}Ge impinging on a target with $Z = 80$–90 has been suggested) when/if the appropriate accelerators become available.

Rutherford (1929) first used properties of radioactive isotopes to establish a time scale for nucleosynthesis. By attributing all the lead in radioactive minerals to the decay of normal uranium (^{238}U) and a then-hypothetical isotope

TABLE IX. Nucleocosmochronology. Ratios of unstable isotopes that give information about the time scale of nucleosynthesis. The first column gives the ratio considered and the second the decay rates of the two nuclides concerned. The third column is the observed abundance ratio, either now or at the time the meteorites began to retain xenon ("at solidification"). The fourth column is the ratio in which the two nuclides are produced by the r process, and the fifth column the time scale implied by the comparison of observed ratio with production ratio. The first two time scales are a measure of the time between the end of the nucleosynthesis that contributed to Solar System material and solidification. The third and fifth are the time from the median time of r-process synthesis to the present, and the fourth and sixth the time from the median time of synthesis to solidification.[a]

Ratio	Decay rates (yr^{-1})	Observed abundance ratio	Production ratio	Time scale
$^{129}I/^{127}I$	4.077×10^{-8} Stable	$1.07 \pm 0.04 \times 10^{-4}$ (at solidification)	$1.5\ ^{+1.4}_{-0.5}$	$2.3 \pm 0.15 \times 10^8$
$^{244}Pu/^{232}Th$	8.474×10^{-9} 4.99×10^{-11}	0.0062 ± 0.002 (at solidification)	0.47 ± 0.1	$5.1 \pm 0.7 \times 10^8$
$^{235}U/^{238}U$	9.72×10^{-10} 1.537×10^{-10}	$1/137.8$ (now)	1.5 ± 0.5	$6.5 \pm 0.3 \times 10^9$
		0.313 ± 0.026 (at solidification)	1.5 ± 0.5	$1.9 \pm 0.3 \times 30^9$
$^{232}Th/^{238}U$	4.99×10^{-11} 1.537×10^{-10}	$3.9_{-0.2}^{+0.5}$ (now)	$1.9_{-0.3}^{+0.2}$	$6.9_{-1.6}^{+3} \times 10^9$
		$2.4_{-0.15}^{+0.35}$ (at solidification)	$1.9_{-0.3}^{+0.2}$	$2.3_{-1.6}^{+3} \times 10^9$

[a] Data from Schramm (1973, 1974).

which was the progenitor of actinium (^{235}U), he was able to calculate the half-life of ^{235}U as about 4×10^8 yr and thus to conclude that, even if the two isotopes had originally been produced in equal amounts, the uranium in the Earth could not be much older than 3.4×10^9 yr. From this, and an estimated age for the Sun of 7×10^{12} yr, he concluded that "the processes of production of elements like uranium were certainly taking place in the Sun 4×10^9 years ago and probably still continue today." If the word "Sun" is replaced by "Galaxy" his statement stands today.

The general subject of disentangling the history of nucleosynthesis over the lifetime of the Galaxy is called nucleocosmochronology and has been reviewed by Schramm (1973a, 1974). The general principle, of course, is that if we know the abundance ratio of two nuclides as they are produced, their lifetimes, and the ratio now (or at the time the Solar System was formed) we can calculate the average elapsed time from synthesis to the present or Solar System formation. A wide variety of nuclides have been suggested as capable of yielding information. Of these, ^{205}Pb (half-life 3×10^7 yr) and ^{176}Lu (half-life $= 2.2 \times 10^{10}$ yr) are produced entirely by the s process, ^{146}Sm (half-life $= 1.2 \times 10^8$ yr) by the p process, and ^{187}Re (half-life $= 4.3 \times 10^{10}$ yr) by both the s and r processes. For these and many others, either the observational techniques or the nuclear theory are inadequate for a meaningful comparison of production to observed abundance to be made. Thus, virtually all of the results of nucleocosmochronology to date have come from comparing the abundance ratio of two isotopes as they should be produced by r-process synthesis with the abundance ratio now or the abundance ratio at the time the meteorites solidified and began retaining volatile daughter nuclides. Besides Rutherford's uranium isotopes the nuclides ^{187}Re, ^{127}I, ^{244}Pu, and ^{232}Th have yielded useful numbers. The data on observed ratios of these are discussed in Sec. II.B.2. ^{129}I and ^{244}Pu, which have short half-lives, are sensitive to the length of time from the end of r-process production to the beginning of volatile (xenon) retention. U and Th, with their longer half-lives, provide a measure of the average ("center of mass") time at which Solar System r-process material was formed. The relative production of the actinides is given quite closely by the number of alpha-emitting progenitors each has (Seeger and Schramm, 1970), with 20% or so corrections for neutron capture during the freeze out (Blake and Schramm, 1973a). The production ratio of $^{129}I/^{127}I$ can also be estimated almost independent of the circumstances under which synthesis occurs. Since iodine falls on the left side of the $A = 130$ peak, the ratio is surely greater than 1. Fowler (1972) has used the observed isotope ratios of nearby, stable, tellurium to estimate the production ratio of $^{129}I/^{127}I$ as about 1.5. Table IX lists the available observational and nuclear data on the "useful" chronometers and the information that has been derived from them. The sources and reliability of both observations and calculations are further discussed by Schramm (1973a, 1974). In general, we can conclude that about 10^8 yr elapsed between the end of r-process nucleosynthesis and the beginning of xenon retention by the meteorites (since ^{129}I is now found only as its daughter ^{129}Xe, and ^{244}Pu only as its Xe fission products). In addition, synthesis had evidently been going on for some billions of years before the formation of the Solar System. The implications of these results for our knowledge of the age of the universe have been explored by Gott *et al.* (1974). They do not contradict anything else we think we know. The ^{187}Re chronology is now believed to be consistent with the results from U, Th, Pu, and I (Fowler, 1972; Woosley *et al.*, 1975).

Before summarizing what we think we know about the r process, it is only fair to say that very strong doubts have been expressed about the entire scheme. Suess and Zeh (1973) have presented an empirically based discussion of the abundances of the heavy nuclides and a critique of the standard approach to nucleosynthesis, while Amiet and Zeh (1967, 1968) suggest an alternative way of forming the neutron-rich heavy nuclei. They point out that synthesis at very high density ($\sim 10^9$ g cm^{-3}) occurs along a valley of beta stability which is shifted in the high N direction by the high Fermi level of the electrons. Thus, if hydrostatic synthesis were to occur under these circumstances and the products be ejected very quickly (compared to the time to establish equilibrium through beta decays), subsequent decay of the nuclides will produce something like standard

r-process abundances. Unless the ejection occurs very rapidly, equilibrium will be maintained as the density gradually drops, and the abundance peaks will shift to the magic neutron numbers, mimicking the s process. A suitable site for such a process may also be difficult to find. Additional counterarguments are given by Peters *et al.* (1972).

Ignoring these doubts, we can say that most theorists agree that the neutron-rich isotopes of the elements beyond iron were predominantly formed by the rapid capture of neutrons in an explosive (probably supernova) environment. A theory based on this assumption can reproduce the general trends of Solar System abundances of the neutron-rich nuclides, starting with a variety of reasonable initial conditions of temperature, density, and neutron mass fraction. The theory enables us to predict how much of various radioactive isotopes should have been produced at the same time as the stable ones and therefore to estimate the time scale over which the nucleosynthesis that contributed to Solar System material occurred. That time scale is consistent with other estimates of the age of the Galaxy.

3. The p process

Something like 32–36 nuclides, all on the proton-rich side of the region of beta stability, remain unaccounted for. Their abundances are nearly all lower than those of the neighboring nuclides. They cannot have been made by reactions among heavy charged particles since at temperatures high enough to penetrate the Coulomb barrier of, e.g., Fe, equilibrium is established and all the matter ends up in the iron peak. Nor can they have been made by neutron capture reactions, being shielded by stable, more neutron-rich isobars. It is clear that either (a) protons have been added by (p,γ) reactions, (b) neutrons have been removed by (γ,n) or (p,n) reactions, or (c) neutrons have been converted to protons by positron decay or photobeta reactions acting on other heavy nuclei. Most of these possibilities were explored by B^{2}FH and Cameron (1957). The nuclides made by this p (for proton) process are necessarily either secondary (if the progenitors were made in the r process) or tertiary (if the progenitors are s-process nuclei) products of nucleosynthesis. Their abundances can, therefore, be expected to vary widely among stars with different total metal abundances. Because no single element is dominated by a p-process isotope, this will only be testable by accurate measurements of isotope ratios outside the Solar System. No very relevant data are presently available, so that our only handle on the p process is the Solar System distribution of the nuclides attributed to it.

Figure 2(b) shows the Solar System abundances of nuclides heavier than the iron peak which are thought to be made by the p process. The distribution shows several interesting features. First, there is a strong resemblance in general shape between the abundance curve for p nuclei and those for the r- and s-process ones, but the p, or "bypassed," nuclides are significantly lower in abundance. (The total masses of s-, r-, and p-process products with $A \geq 70$ are in the approximate ratio 1.0:0.5:0.02.) Second, the abundance ratio of bypassed nuclei to neutron-capture products decreases gradually with increasing A. Third, the abundance curve shows features associated with the closed neutron shells at $N = 50$ and 82, but the peaks, at ^{92}Mo and ^{144}Sm, are shifted to slightly higher A. This suggests proton capture rather than neutron removal as the dominant process in their formation. There is probably also a peak at the closed proton shell $Z = 50$ which also favors proton capture as the dominant mechanism. Finally, in those elements which have more than one p-only isotope (Ru, Cd, Xe, Ba, Ce, Dy, Sn), the relative abundances of the several isotopes never differ by more than a factor of 2 or 3. This, too, seems more consistent with the capture of a couple of protons by parent nuclides of comparable abundance than with the photoexpulsion of neutrons, which would be expected to result in a much lower abundance for the lowest A isotope of each element.

These considerations led B^{2}FH to study the equilibrium that would be set up if ordinary, Population I material ($X = 0.7$, $Y = 0.28$, $Z = 0.02$) were heated above 10^9°K at a density ≥ 100 g cm^{-3} and then cooled on a time scale less than typical beta decay times of proton-rich nuclei (about 1000 sec). They concluded, under these circumstances, (a) that light nuclides (^{12}C, ^{16}O, etc.) could not possibly capture all the available protons, so plenty would be left for heavy nuclei to capture, (b) that an equilibrium would be established among adjacent nuclides, such that

$$\log\frac{n(A+1, Z+1)}{n(A,Z)} = \log n_p - 34.07 - 1.5\log T_9 + \frac{5.04}{T_9}Q_p,$$

where n_p is the proton number density, T_9 the temperature in units of 10^9°K, and Q_p the binding energy of the $(Z+1)$st proton, up to the point where the next proton would no longer be bound at the particular temperature T_9, and (c) that this would allow the addition of one to three protons to a typical seed nucleus, if the initial temperature were $(2\text{–}3) \times 10^9$°K. They suggested the envelopes of Type II supernovae as a suitable site for this explosive addition of protons to r and s seed nuclei. In order to provide sufficient p-process material in their picture, almost half the mass of a typical Type II supernova would have to have its r- and s-process nuclei converted to p nuclei. Several factors can modify this estimate. If all supernovae contribute (e.g., all stars with masses of 4–8 $\mathfrak{M}_\odot$ on the main sequence) then 100% efficient conversion in 10%–15% of the star is enough (Truran and Cameron, 1971). On the other hand, if (as seems more probable) the conversion is only 25%–50% efficient, we are back to needing about half the mass of the star, i.e., most of its remaining hydrogen-rich envelope. Finally, if the stellar envelope has been enriched in seed nuclei prior to the supernova event (by mixing to the surface of the products of the s process), the requirements on mass and efficiency drop down again by a factor of 10 or more.

We will return to this general scheme of synthesis under explosive, hydrogen-rich conditions after exploring several alternative forms of the p process which seem to be less successful. Franck-Kamenetskii (1961) explored the production of "bypassed" nuclides by cosmic ray spallation. More neutrons than protons are often knocked off, producing the right sorts of nuclides, and such spallation must occur. But the observed abundances of Li, Be, and B (both in the Solar System and in the cosmic rays themselves) tell us how much spallation has occurred, and the total amount is inadequate by about a factor of 10 to have produced the totality of p-process nuclides (Audouze, 1970; M. Meneguzzi and J. Audouze, private communication). The two very low abundance odd–odd nuclei, ^{138}La and ^{180}Ta, probably are produced in this way, and the heaviest p-process nuclides (isotopes of Hg, Ta, and W) may also be made by spallation of lead isotopes, which provide a relatively abundant seed.

At temperatures above 10^9°K, production of electron positron pairs begins to occur. The reactions $(A,Z) + e^+ \rightarrow (A, Z+1) + \bar{\nu}_e$ and $(A,Z) + \gamma \rightarrow (A, Z+1) + e^- + \bar{\nu}_e$ can then occur. These also produce suitable nuclides. Such weak interaction processes have been explored by Reeves and Stewart (1965), Arnould and Brihaye (1969), and Agnese *et al.* (1969). The chief difficulty is that the rate of these reactions varies too sharply from nucleus to nucleus to give the observed smoothness of the p-process abundance curve. This sort of process may have contributed to the abundance of ^{164}Er, which lies well above the main p abundance curve, and is 10 times more abundant than its p-process isotope ^{162}Er (Audouze, in NATO, 1974). But since ^{164}Er is also made in a branch of the s process, it is not clear that any such explanation is necessary.

Arnould (in NATO, 1974) has explored the production of bypassed nuclides under the conditions of hydrostatic oxygen burning. Comparable numbers of free neutrons and protons are available due to reactions like $^{16}O + {}^{16}O \rightarrow {}^{31}P + p$ and $^{16}O + {}^{16}O \rightarrow {}^{31}S + n$ as well as from (γ,n) and (γ,p) reactions on various nuclides. Starting with a normal Pop I complement of r- and s-process seeds, hydrostatic burning at $T = 2.2. \times 10^9$°K and a neutron density $n_n = 10^{14}$ cm^{-3} overproduced the bypassed nuclides with $A = 100$–144 by a factor of about 10^4 in roughly their solar proportions. The lighter nuclides are similarly overproduced at lower neutron densities, and the heavier ones at lower temperatures. Thus, if 0.01% of the Solar System material had been through hydrostatic oxygen burning, the p nuclei would be accounted for. This seems reasonable enough. The difficult part is to get the stuff out without having the relative abundances of these rare and rather delicate nuclei perturbed beyond recognition. Since the major products of hydrotstatic oxygen burning do not get out without further processing, this is a serious objection. In addition, the same conditions acting on much more abundant, lower A seeds (C, O, Ne, etc.) will make large amounts of other odd nuclei which we do not see, unless the seeds for the "p processing" were already considerably enriched in the hydrostatic oxygen burning zone due to a prior admixture of s-process material.

In view of the problems of getting hydrostatically produced material out undisturbed, Howard (in NATO, 1974) has considered the production of rare, proton-rich nuclides during explosive carbon and oxygen burning. We have already seen (Sec. III.D.1) that explosive carbon-burning conditions probably contribute many of the rare nuclides below $A = 70$. Under the conditions of explosive oxygen burning, all the s- and r-process seeds disintegrate back to iron, which is not terribly helpful. But at intermediate temperatures $(T_0 = (2.2$–$3.0) \times 10^9$°K) partial photodisintegration of lead and other massive seeds occurs. This can produce adequate amounts of the bypassed nuclei from ^{144}Sm to ^{196}Hg, in the proper relation to ^{29,30,31}Si (which are made under the same conditions), provided that the s-process seeds were initially overabundant by a factor of 5 to 10. The synthesis can occur over a wide range of densities (10^{1-4} g cm^{-3}). The production of the lighter bypassed nuclides, $A = 72$–144, is insufficient by a factor of 10–100 in these explosive calculations.

Novae have not been suggested as a site of the p process by Hoyle and Clayton (1974).

Calculations of the production of bypassed nuclei under conditions which might represent the hydrogen-rich outer envelope of a supernova during or just after the passage of a shock wave have been carried out by many workers since the pioneering work of B^{2}FH and Cameron (1957). Published work includes that of Ito (1961), Macklin (1970), and Truran and Cameron (1972). Further calculations by Andouze and Truran (1975), W. M. Howard and S. E. Woosley, and J. W. Truran and A. G. W. Cameron are in progress. Several conclusions are common to all the investigations. First, the production of low A bypassed nuclides is dominated by (p,γ) reactions and that of higher A's by (γ,n) reactions. Second, an initial temperature of at least 2×10^9°K at an initial density of 100 g cm^{-3} is required if significant production of p-process nuclides is to be achieved on the hydrodynamical time scale. In addition, the structure of the shock wave is such that sufficiently high temperatures are only reached in matter with a density of at least 10^4 g cm^{-3}. Given these conditions, then most of the p-process nuclides (i.e., 30 out of 36) are made in roughly Solar System proportions and in adequate amounts if 1% of the Solar System material has been through the process (that is, 1% of all r and s seeds converted to p nuclei). It is interesting that about 1% of Solar System material must also have been exposed to the r and s processes; the great difference in product abundances is due to the great difference in seed abundances (iron peak for r and s; r and s products in turn for p), not to difference in exposure. The nuclei which are not adequately produced in these calculations are, for the most part, those like ^{164}Er, ^{138}La, and ^{180}Ta which can be attributed to branching in the s process or cosmic ray spallation. The difficulty is that only the base of a typical presupernova envelope will be at high enough density for the requisite conditions to be achieved. This means that much less than the necessary one solar mass or so of material per supernova will have been exposed to p processing. The implication is that the envelope of the star must already have been greatly enriched in s-process elements to act as seeds before the start of the explosion. Since S stars, Ba II stars, and the like (see Sec. II.C for a discussion of these evolved stars with surface s-process enhancements) do not constitute the majority of red giants it is not clear that the necessary enrichment is generally present. The stellar models mentioned in Sec. E.1 suggest that s-process enhancement is most readily achieved in stars of lower masses than those which normally give rise to supernovae. In addition, the supernovae involved cannot be the same ones that Howard *et al.* (1971) suggested as the site of explosive helium burning to make ^{15}N, ^{18}O, etc., since they required a lower temperature in the He-rich zone than we require here in the H-rich zone.

In an effort to overcome this difficulty, Truran and Cameron (work in progress) have considered the synthesis that might occur in the supernova shock itself (according to the models of Colgate, 1974) before thermal equilibrium is established. Production of bypassed nuclides then proceeds throughout the hydrogen-rich envelope as the protons are accelerated to 1–10 MeV in the shock wave. The conversion of r and s seeds to p nuclei in this case seems to be much less efficient than in the standard explosive calculations, so that initial seed enhancement will probably still be required.

Production of the bypassed nuclei by (p,γ) and (γ,n) reactions in hydrogen-rich regions under explosive conditions appears to be fairly well established. The chief difficulty is in identifying an astronomical site in which a quantity of s- and r-process seeds sufficient to make the Solar System abundance of p-process products can be exposed to the necessary high temperatures and proton densities. A need for an initial enhancement of the r- and/or s-process seeds for the p process seems to be a feature of all the calculations which have been done so far.

F. Summing over the processes to get the yield per star and the yield per generation of heavy elements

The time has come to leave the realm of hard, theoretical fact, and re-enter that of observational speculation. What we would like to be able to do is first to compile a table (for instance) listing the amount of each stable isotope of the 94 (or more) naturally produced elements returned to the interstellar medium by a star as a function of initial mass and composition of the star. (For a typical star, matter will be returned at more than one stage in its life.) Then, by integrating over the distribution of stellar masses and compositions formed at a particular time and place (or over the entire galaxy), we would determine the predicted production of all elements per generation of stars (or over the history of the galaxy). Theory fails completely to tell us two of the vital parameters needed for this program: the mass of the remnant left behind after various kinds of stellar deaths, and the numbers of stars of various masses that will be formed from the interstellar medium (as a function of its composition or temperature or density, or whatever the important variables turn out to be). In addition, about half the stars in the solar neighborhood are members of binary (or higher multiplicity) systems (Batten, 1973), and theory tells us virtually nothing about the amount of mass lost from close binary systems (or what its composition should be) or how remnant size will be affected by a close companion. The missing information must be supplied by some judicious combination of observational evidence and divine inspiration.

Most of the models of galactic evolution attempted so far (Tinsley, 1974; Larson, 1974; Talbot and Arnett, 1974; and references in each to earlier work) simplify these problems almost beyond recognition. For instance, the remnant mass is usually assumed to be the same for all stars in a wide mass range (e.g., a $0.7\,\mathfrak{M}_\odot$ white dwarf for all stars with main sequence masses below $5\,\mathfrak{M}_\odot$ and a $1.4\,\mathfrak{M}_\odot$ neutron star for everything above $5\,\mathfrak{M}_\odot$). The distribution of stellar masses at birth is normally assumed to be that observed in the solar neighborhood at the present time ("constant initial mass function") or some other power law that varies in a simple way with time ("variable IMF"). The ejected composition is normally represented by only a few main constituents. For instance, Larson (1974) assumes an ejected mass of heavy elements $M(Z) = m[q_c + (1 - q_c)Z_f] - 1.4$, where m is the initial main sequence mass of the star, Z_f is its initial metal abundance, and q_c is the fraction of the star's mass interior to the hydrogen-burning shell. An analytic fit to evolutionary models gives $q_c = 0.156 + 0.536\log(m/9)$. Talbot and Arnett (1974) have treated the ejected composition in a little more detail, considering the amounts of C, O, Ne, Si, and Fe returned as a function of main sequences mass, as well as the secondary production of nitrogen and s-process nuclides. The effects of binary systems and of the initial composition of the star on its final structure are normally neglected completely.

There are several good reasons for this degree of simplification. One is the finite size of computers and computing budgets. In addition, the observations of chemical composition and other properties of galaxies (Sec. IV.A) to be matched by the models are themselves not very detailed and do not justify a more complicated treatment. Finally, the calculations and observations needed to improve our estimates of the amount and composition of ejected matter are still far from complete. Table X is a zeroth order attempt at expressing ejected composition as a function of main sequence mass. The various kinds of assumptions, calculations, and observations that went into it are discussed in the paragraphs that follow. Ejected masses are in $\mathfrak{M}_\odot$ from one star with initial composition X, Y, Z, the metals being initially present in solar proportions.

1. Remnant masses: Both theory (e.g., Paczyński, 1970) and observation suggest that the mass of a white dwarf increases with the mass of its main sequence progenitor. The observations include the low mass ($0.42\,\mathfrak{M}_\odot$) of 40 Eri B and the high mass of Sirius B ($1.0\,\mathfrak{M}_\odot$) coupled with the probable evolutionary state of their systems, and the high average masses of the white dwarfs in nova systems (Warner, 1973) and in the Hyades, where the present main sequence turnoff mass is about $2.3\,\mathfrak{M}_\odot$. The white dwarf remnant masses in the table are those calculated by Paczyński (1970a). White dwarf formation is taken to extend up to $6\,\mathfrak{M}_\odot$ from the probable presence of a white dwarf in the Pleiades. The possibility of a narrow range of stars which explode by degenerate carbon detonation (Sec. III.D.1) is indicated by the $6.5\,\mathfrak{M}_\odot$ star leaving no remnant. Stars massive enough to ignite carbon without complete disruption are all assumed to leave $1.4\,\mathfrak{M}_\odot$ cores behind in the form of neutron stars. There are other possibilities, including smaller mass neutron stars (Ergma and Paczyński, 1974) and black holes, but a comparison of the rate of pulsar formation with that of deaths of massive stars (e.g., Ostriker *et al.*, 1974) suggests that at least a major fraction of such deaths must give rise to neutron stars capable of pulsar activity, and the models of presupernova stars all have core masses which converge to $1.4\,\mathfrak{M}_\odot$, independent of total star mass (Arnett and Schramm, 1973). In addition, the three neutron stars for which we have some model-independent information appear to fall between 1.0 and $1.5\,\mathfrak{M}_\odot$. They are NP 0532

in the Crab Nebula (Carter and Quintana, 1973), Hercules X-1 (Middleditch *et al.*, 1975), and the "binary pulsar" Psr 1913 + 16 (Taylor and Hulse, 1975).

2. Mixing and the *s* process: Stars of relatively low mass will have all the material which has undergone nuclear reactions left behind in the remnant unless convection brings the products into the hydrogen-rich envelope before it is lost. The only evolutionary calculation in which the requisite mixing happens unequivocally is that of Iben (1974), discussed in Sec. III.E.1; the model is for a $7\,\mathfrak{M}_\odot$ star and might be expected to apply over the range 3–$10\,\mathfrak{M}_\odot$. All stars in this mass range are assumed to have the amount of mixing implied by the model (which, observationally, will result in some enhancement of helium and the *s*-process elements and shifts in the C:N:O element and isotope ratios). Although there are no completely satisfactory models with mixing for lower mass stars, we know that it must occur at least some of the time from the existence of the metal-poor (hence old and low-mass) CH subgiants whose surfaces are enriched in *s*-process material and carbon (Bond, 1974). Mixing sufficient to produce the CH subgiants phenomenon is assumed to happen in 10% of stars of 1–$3\,\mathfrak{M}_\odot$, while all low mass stars are given enough mixing of helium into their envelopes to yield the slight He enhancement seen in the planetary nebulae. The slight depletion of iron-peak elements caused by this amount of *s*-processing is neglected, but the changes in envelope abundance of H, He, and C are not. It is difficult to predict whether the envelope abundances of C and O will increase or decrease as a star evolves. Mixing up from the helium burning region (required to bring up the *s* nuclei) will increase them, but conversion to ^{14}N in the hydrogen-burning shell (or ^{13}C in lower mass stars) and the subsequent use of the ^{14}N as a neutron source tend to deplete them. The observations (see Table III) include both high and low *C* abundances associated with *s*-process enhancement, but oxygen is normally depleted and nitrogen enhanced. The evidence of the planetary nebulae (Table IV) is that all three are very slightly enhanced, nitrogen perhaps more than C and O. The table assumes enhancement of C, N, and O in the envelope by factors of 1.3, 1.6 ($Z/0.0175$), and 1.3, respectively, for stars below $7\,\mathfrak{M}_\odot$.

3. "Minor" processes: One of the general conclusions that can be drawn from the entire discussion of explosive nucleosynthesis (Sec. III.D) is that if the major constituents in the range from carbon to iron can be made to come out in the right proportions, then the less abundant ones will, too. We make use of this by considering only C, N, O, Ne, and Mg separately (including F and Na with Ne, and Al with Mg), and grouping the elements from Si to Ca as "Si" and the isotopes from ^{45}Sc to $A = 70$ as "Fe." The nuclides beyond $A = 70$ are all attributed to the *s*, *r*, and *p* processes (q.v.). Li, Be, and B, along with ^{2}H and ^{3}He (which seem to be more often destroyed in stars than produced there) are neglected completely. The *p* process is assumed to convert half the *r* and *s* seeds in the bottom 20% of the hydrogen-rich envelope to *p* nuclides in all stars which supernova. The overlap of this range with the range where Iben's models apply provides some of the necessary initial enhancement of seeds for the process.

TABLE X. Products of stars at the end point of their evolution as a function of main sequence mass. The amounts of all products are given in solar masses. "Old" H, *e*, and metals refer to unprocessed material re-ejected into the interstellar medium. The observations and theory leading up to this table are discussed in Sec. III.F. The last column is the composition of one solar mass of material having $X = 0.7018$, $Y = 0.2807$, and $Z = 0.0175$ divided among the heavy elements in accordance with the abundances given by Cameron (1973). The stars have an initial composition X, Y, Z, which affects the amounts of some of the products in a straightforward way. It is not, in general, known how the relative amounts of C, O, Ne, Mg, Si, Fe, etc. are affected by variation in initial composition. $Z' = Z/(0.0175)$.

Product \ Main sequence mass	0.8	1.5	3.0	6.0	6.5	7.0	15	22	50	"Sun"
Remnant	0.6	0.8	1.2	1.2	0.0	1.4	1.4	1.4	1.4	...
"Old" H	$0.2X$	$0.7X-0.082$	$1.8X-0.165$	$4.8X-0.40$	$4.5X-0.20$	$5.0X-0.25$	$11X$	$14X$	$28X$	0.7018
"Old" He	$0.2Y$	$0.7Y-0.021$	$1.8Y-0.041$	$4.8Y-0.010$	$4.5Y-0.05$	$5.0Y-0.021$	$11Y$	$14Y$	$28Y$	0.2807
"New" He	...	0.8	0.16	0.38	0.60	0.65	1.64	3.1	5.1	...
"Old" metals	$0.2Z$	$0.7Z$	$1.8Z$	$4.8Z$	$4.5Z$	$5.0Z$	$11Z$	$14Z$	$28Z$	0.0175
"New" C	...	6.6×10^{-4}	1.7×10^{-3}	4.5×10^{-3}	0.1	0.1	0.18	1.00	5.0	3.134×10^{-3}
N	...	$4.7\times10^{-4}Z'$	$1.2\times10^{-3}Z'$	$3.2\times10^{-3}Z'$	$3.3\times10^{-3}Z'$	$3.4\times10^{-3}Z'$	$0.026Z'$	$0.049Z'$	$0.081Z'$	1.126×10^{-3}
O	...	1.6×10^{-3}	4.1×10^{-3}	1.1×10^{-2}	0.1	0.1	0.18	1.27	6.5	7.577×10^{-3}
Ne	...	...	...	...	...	...	0.09	0.60	2.15	1.568×10^{-3}
Mg	...	...	...	...	...	...	0.18	0.36	0.67	6.155×10^{-4}
Si(Si to Ca)	...	...	...	...	...	...	0.18	0.24	0.67	1.147×10^{-3}
Fe(Sc to A = 70)	...	...	...	...	1.4	...	...	...	0.54	1.122×10^{-3}
s process	...	$4.4\times10^{-7}Z'$	$1.1\times10^{-6}Z'$	$1.0\times10^{-5}Z'$	$1.0\times10^{-5}Z'$	$1.1\times10^{-5}Z'$	...	...	...	2.121×10^{-7}
r process	...	...	...	...	...	...	...	...	5×10^{-5}	1.060×10^{-7}
p process	...	...	...	$2.8\times10^{-6}Z'$	$3.0\times10^{-6}Z'$	$7.0\times10^{-7}Z'$	$7.0\times10^{-7}Z'$	$8.9\times10^{-7}Z'$	$1.7\times10^{-6}Z'$	4.242×10^{-9}

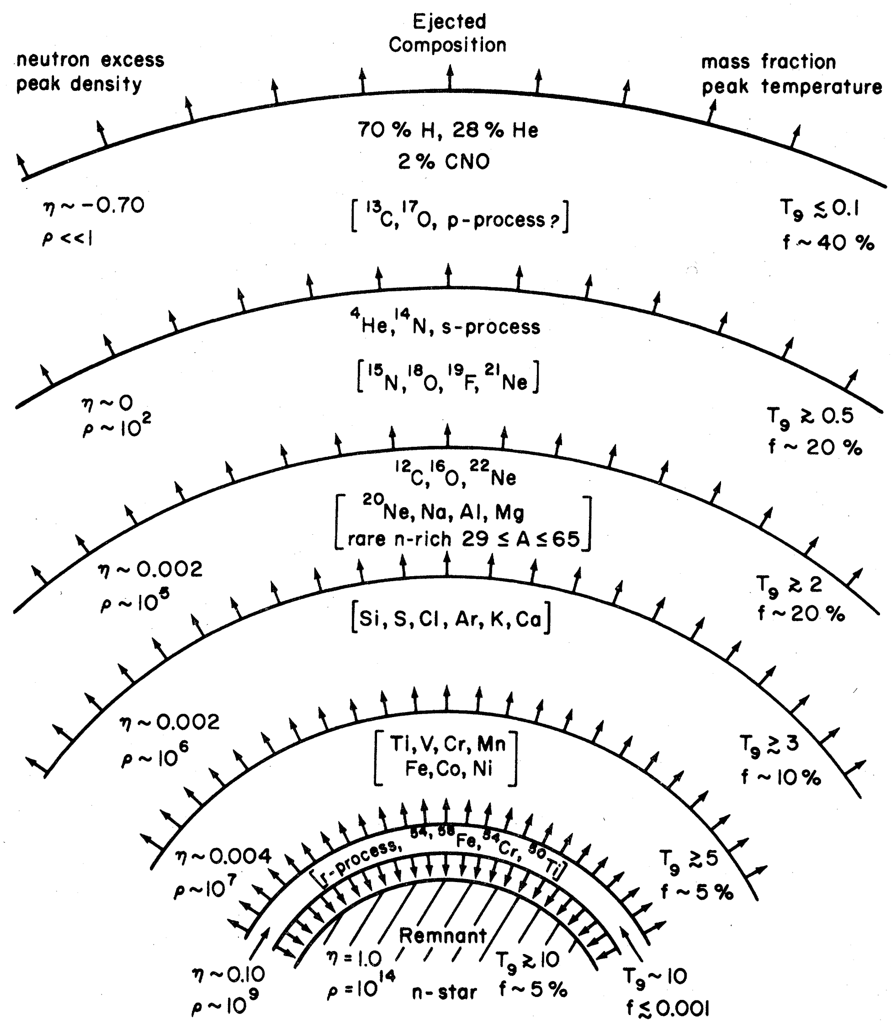

FIG. 14. A theorist's (Woosley, in NATO, 1974) composite supernova. It represents a star in which electron capture has transformed the center into a neutron star, resulting in the liberation of large amounts of gravitational potential energy. This energy (through a still-poorly-understood transport mechanism) ejects the outer layers of the star and heats the layers sufficiently that various nuclear reactions occur under explosive conditions. In each layer, the properties indicated are, from left to right: the neutron excess, $\eta = (n_n - n_p)/(n_n + n_p)$; the density, ρ, in units of g cm^{-3}; the major constitutents (nuclides whose abundances are largely determined by the explosion itself are in brackets; other nuclides are the products of earlier hydrostatic evolution); the temperature, T_9, in units of 10^9K, achieved just behind the supernova shock wave; and the fraction of the total stellar mass, f, in each zone. The precise position of the "mass cut" between the compact remnant and the ejected envelope will determine the abundance of iron-peak and r-process elements in the galaxy as a whole and in the cosmic rays, if they are directly accelerated by supernovae.

4. "Major" constituents: The relative proportions of H, He, C, O, Ne, Mg, "Si," and "Fe" in the ejected composition were taken from the models of Arnett (1974a,b; in NATO, 1974; Talbot and Arnett, 1974). The models have been calculated only for an initial Pop I composition, and the effects of different initial composition are very poorly known (but, it is hoped, small!), except that dropping the initial metal abundance by a factor $\gtrsim 10$ increases the size of the zone between the hydrogen and helium burning shells and, thus, the amount of helium ejected (W. D. Arnett, private communication). The relative proportions of carbon and oxygen have been adjusted to reflect recent measurements and extrapolations of the nuclear reactions involved in helium burning (see Sec. III.C.2).

The last column in the table shows the fractional composition by mass of material which has $X = 0.7018$, $Y = 0.2807$, and $Z = 0.0175$ [divided among the heavy elements according to the abundances given by Cameron (1973)]. It is left as an exercise for the reader to attempt to reproduce the last column by summing various proportions of the other nine.

Figure 14 illustrates a slightly different form of "summing of processes." It represents a theorist's (S. Woosley) average Type II supernova (main sequence mass = 10–60 $\mathfrak{M}_\odot$). The constituents shown in brackets are produced on a hydrodynamic time scale during the supernova event. The others were produced during the hydrostatic evolution of the star. In addition to the ejected composition of the star, the figure shows temperature, density, neutron excess, $\eta = (n_n - n_p)/(n_n + n_p)$, and the fraction of the total mass in each of the zones.

If Table X were more accurate and more complete (including entries for stars of varying initial composition and for binaries) we could predict the nucleosynthesis by any population of stars simply by summing suitable proportions of the products of each type of star. Stars with evolutionary time scales longer than the present age of the Galaxy cannot be neglected because they remove a certain fraction of the gas from the interstellar medium without giving anything back. Such stars used up from 75% to 88% of the gas that has condensed into stars in the solar neighborhood. Of the gas that goes into stars above $1\,\mathfrak{M}_\odot$, about one-third remains behind in remnants, and two-thirds is eventually returned to the interstellar medium, if the distribution of stellar masses is that presently observed in the solar neighborhood. Of the 8%–17% of the gas that went into stars in a given generation and eventually comes out, 4%–8% of the mass returned is in the form of new heavy elements, a comparable amount is new helium, and the rest is unprocessed material (Arnett, in NATO, 1974), again if the distribution of masses is the local one. Finally, if new generations of stars use up old and new gas in the proportions in which they exist in the interstellar medium (i.e., the galaxy remains a well mixed closed system), it can be shown that (we return to this in Sec. IV)

$$X_i = \frac{q_i}{1-f}\ln\left(\frac{M(\text{gas}, t=0)}{M(\text{gas}, t=t)}\right),$$

where X_i is the present interstellar abundance by mass of the ith element, f is the fraction of all mass (per generation) which goes into stars which evolve in less than 10^{10} yr, q_i is the fraction of all mass that went into stars in a typical generation which comes out as the ith element (the "yield" of that element), and the M's are the total masses of the interstellar gas at some initial time and at the present. The effects of varying some of the many assumptions mentioned in this paragraph are discussed in Sec. IV.

It is easy to identify a number of observations which could greatly improve the contents of Table X and similar compilations. Among these are (a) measurement of the abundance of s-process elements in planetary nebulae and more data on C, N, and O in them, (b) detection of and abundance determinations for iron-peak and r-process elements in the spectra of supernovae or their young remnants, (c) any measurement of p-process isotope abundances anywhere outside the Solar System. Other badly needed information that might come either from theory or observations includes (a) masses of individual white dwarfs and neutron stars whose main sequence mass can somehow be estimated, (b) any information about how much mass is lost from close binary systems, when, and what its composition is, (c) an indication of how mixing and other properties of a star relevant to its nucleosynthesis are affected by variations in rotation rate and magnetic field, (d) information on variations of ejected composition as a function of initial composition, and (e) how the initial mass function (numbers of stars formed as a function of mass) depends on gas composition and whatever else it in fact depends on. The fact that these problems are easy to identify does not mean that the observations are easy, or even possible to do.

In conclusion, we have identified in the previous sections at least one nuclear process (and a plausible site for it) which is capable of producing each of the nuclides we see (with the possible exceptions of ^{44}Ca and ^{10,11}B if the boron abundance is as high as that given by Cameron, 1973). A sufficient amount of processing by each of the nuclear reaction chains considered seems possible in isolation. It is not yet possible to say for certain whether all of them can be made to occur in the proper proportions to produce the observed abundance distribution within a population of stars that bears some relation to the one we see in the solar neighborhood. It is perhaps encouraging that no absolute contradictions have been uncovered, in the form, for instance, of a process which is necessary to produce nuclide A, but which simultaneously vastly overproduces nuclide A′ or destroys all of nuclide A″. Neither do there seem to be any zones in a model of an evolved massive star which are not needed for nucleosynthesis in supernovae, or any essential zones which do not occur in the models. Once again, we must end by saying that not everyone agrees with the comprehensive scheme of nucleosynthesis in stars which has been advocated here. Suess (1968; Suess and Zeh, 1973) has suggested that the relative constancy of metal abundance and abundance ratios with place and time, as well as the details of the abundance vs A curve, is more consistent with nucleosynthesis in a single event, or small group of events, whose nature was more nearly fission of heavy nuclides than fusion of light ones.

IV. THE EVOLUTION OF GALAXIES

The final step in understanding nucleosynthesis must be to account for the distribution of the elements in our own and other galaxies. We would like to be able to explain both the variations in total heavy-element abundance from place to place and time to time and the variations in relative abundance. This latter can be a risky undertaking, given the present uncertainties in the observations. Talbot (1974a) had no sooner finished accounting for the data of Hearnshaw (1972), which implied that [C/Fe] and [Fe/H] are anticorrelated in disk stars of different total metal abundances, in terms of a model of galactic evolution, when Hearnshaw (1974) reported work on an additional group of 30 F and G dwarfs in which [C/H] $\sim$ 1.5 [Fe/H], a correlation in the opposite sense to that of the 1972 data.

A variety of kinds of observations of galaxies besides the chemical ones provide constraints on models of galactic evolution. These include total mass and ratio of star mass to gas mass, color, luminosity, morphology, and rates of star formation and supernovae. These are discussed in Sec. A, and the models in subsequent sections.

A. Observations of galaxies

The variations of total metal abundance and relative abundances with place and time in our own and other galaxies are discussed in Secs. II.C and II.F. The general trends seem to include a gradual increase of metal abundance with time, some of the nuclei which are produced by secondary processes (like nitrogen and the s-process nuclides) perhaps varying more steeply with time than those produced by primary processes (C, O, Fe, etc.), accompanied by a surprisingly small number of stars with

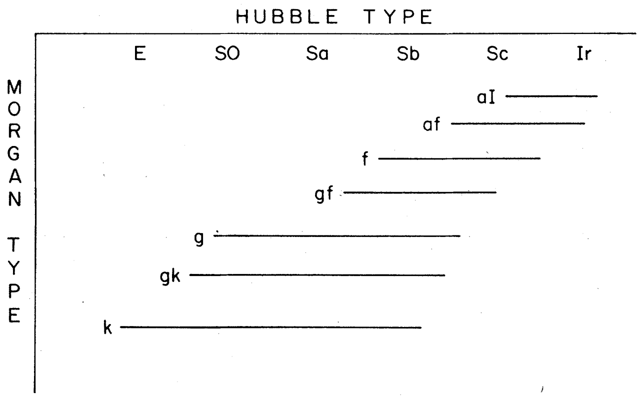

FIG. 15. The correlation of Hubble (1936; Sandage, 1961) and Morgan (1958) types of galaxies. The former is based on over-all morphology and the latter on degree of central concentration (which is closely correlated with spectral type of the central region).

very low metal abundance. The total metal abundance appears to be largest in the most massive galaxies (e.g., Faber, 1973), on the average, and, in addition, many galaxies show a gradient in heavy-element abundance from high at the center to low in the outer parts of disk or halo (see Fig. 9). Both these trends indicate that metal abundance is high in regions of high average density. In this section, we discuss other observed properties of galaxies that may constrain models of galactic evolution. The discussion is largely that of E. M. Burbidge (in NATO, 1974). Other interesting reviews of galactic properties and their correlations have been given by King (1971) and Brosche (1973). The properties of possible extensive galactic halos are discussed in the section on masses.

1. Classification and morphology

There are three systems of classifications of galaxies presently in use. The fundamental distinction, Elliptical, Spiral, and Irregular (with subdivisions within each and Barred as well as Normal Spirals) was first made by Hubble (1926, 1936), who regarded the sequence as an evolutionary one. This system, based on over-all morphology of the galaxies, was further developed by Sandage (1961) and de Vaucouleurs (1959, 1975). It includes elliptical galaxies (E0 to E7, in the direction of increasing ellipticity), spiral galaxies (S0, Sa, Sb, Sc; in the direction of increasing predominance and looseness of winding of the arms; intermediate types can be designated as Sab or Sb$^+$, etc.), and barred spiral galaxies (SB0, SBa, SBb, SBc; rules as for spirals, but with a conspicuous bar extending across the nucleus and the arms beginning from the ends of the bar), and irregular galaxies (IrI and IrII where I has lots of gas and II lots of dust; neither has any particular well-defined shape). The Ir II's are not a homogeneous group (Chromey, 1974). Transition types between spiral and irregular are designated Sd, Sm, and Im. It is no longer generally believed that evolution between major types occurs on time scales of less than 10^{10} yr, but precisely why a given galaxy has a particular morphology is not well understood theoretically. Ellipticals are systematically redder and have less gas and dust and fewer (if any) young massive stars than spirals, while irregulars have the largest fraction of gas and dust and young stars and are the bluest. These correlations immediately suggest rate of star formation as an important distinguishing feature. No very massive ($\gtrsim 10^{11}\mathfrak{M}_\odot$) irregulars are known, nor are there any very low mass ($\lesssim 10^9\mathfrak{M}_\odot$) Sa's and Sb's. The relative sphericity of the E's and great flattening of the S's also indicates that angular momentum per unit mass must be important.

Morgan (1958) recognized a strong correlation between the spectral types of the stars which dominate light output from the center of a galaxy and its degree of central concentration and developed a system of classification based on this correlation as its primary parameter. The classes run from a (light dominated by B and A stars; little or no central concentration) through af, f, fg, g, gk, to k (light dominated by K stars, probably giants, and greatest degree of central concentration). The Morgan types of galaxies are well correlated with Hubble types for objects which have been classified on both systems. The correlation is shown in Fig. 15. Morgan's classification also makes use of a "form" parameter, which includes types S (spiral), B (barred spiral), E (elliptical, including dwarf and giant ellipticals, dE and gE), Ir (irregular), Ep (peculiar ellipticals), L (ellipticals with very low surface brightness, like Sculptor and Fornax), D (ellipticals, but with a flatter light distribution; often the dominant galaxy in a cluster and a radio source), and N (ellipticals with bright, stellar nuclei). Finally, degree of ellipticity in the plane of the sky is indicated by inclination classes 1–7 (1 = circular). The classes as determined from degree of central condensation alone are surprisingly well correlated with colors and with E, S, I types as assigned by Hubble and Sandage or by van den Bergh (1960). This can be understood on evolutionary terms if a high degree of central concentration means that the gas density was initially very high at the center of the galaxy, so that star formation proceeded to completion very rapidly.

Van den Bergh (1960) has found that, for Sb to Ir galaxies, the degree of development of the spiral arms is closely correlated with the total luminosity of a galaxy. He calls the luminosity classes I, II, III, IV, and V, ranging from supergiant (brightest) to dwarf (faintest) by analogy with stellar spectral types. Not all luminosity classes are found for all morphological classes; Sd, Sm, and Im types occur only in classes III, IV, and V, while Sb's occur only in I, II, and III. Among ellipticals, the degree of central concentration is correlated with total luminosity or mass. Van den Bergh (1972a) finds that the crossover from dwarf ellipticals with little or no nucleus to giant ellipticals with a well-defined nucleus occurs at about $2 \times 10^9\mathfrak{M}_\odot$.

2. Luminosity distribution

The optical luminosity of external galaxies ranges from $M_V \sim -6$ to $M_V \sim -21$ (or -22 for QSOs). The distribution of that luminosity or surface brightness across the surfaces of galaxies follows two different patterns. Elliptical galaxies and nuclear bulge and halo component of spirals follow a power law distribution which Hubble (1936) described as $L = L_0/(r+a)^2$, where L_0 is the central luminosity per unit area, and a is some scale constant. De Vaucouleurs (1959) expressed the same phenomenon in terms of surface brightness, B, at radius r vs surface brightness B_e at the radius r_e which contains half

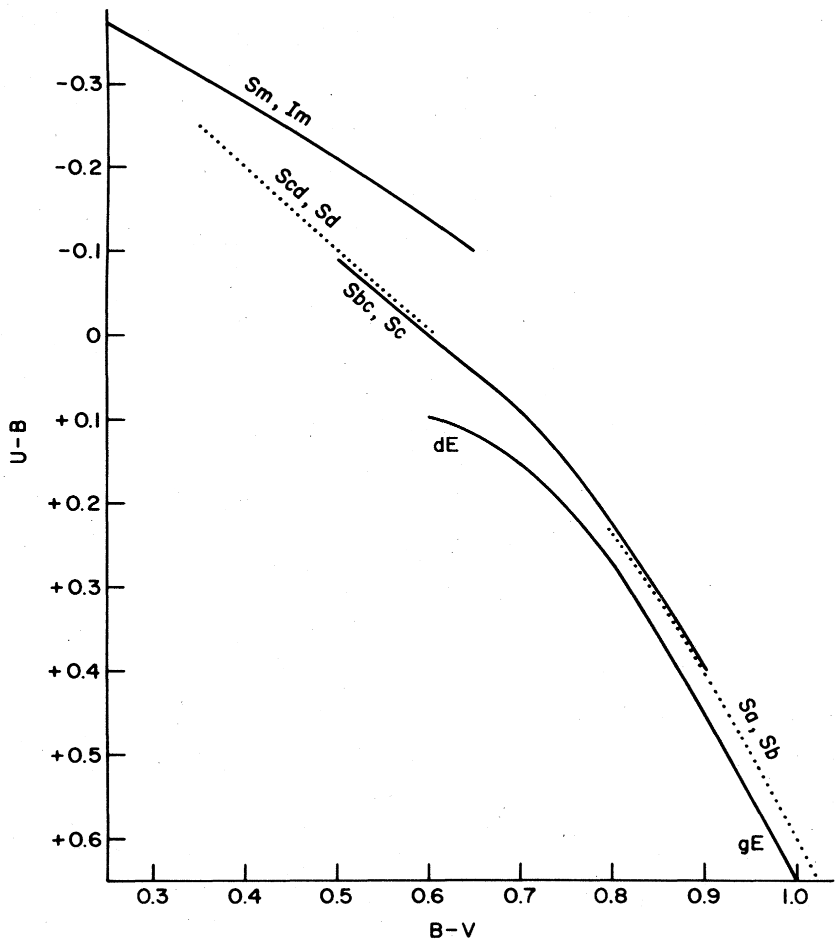

FIG. 16. Mean B-V and U-B colors for galaxies of various Hubble types. The data (de Vaucouleurs, 1960) can be matched by a variety of galactic evolution models.

the total luminosity of an elliptical galaxy as $\log(B/B_e) = 3.33\,[(r/r_e)^{\frac{1}{4}} - 1]$. The observed $L(r)$ distribution most frequently compared with theoretical models is that of Miller and Prendergast (1962) for NGC 3379. The disks of spiral galaxies, including SO's, on the other hand, show an exponential distribution of light with radius, after smoothing over the spiral arms (Freeman, 1970; Sandage *et al.*, 1970; Nordseick, 1973). This represents a much less steep falloff of light with radius over the region of the galaxy interior to the radius containing half the luminosity and can be qualitatively ascribed to rotation preventing material from falling rapidly to the center of the system.

3. Colors and populations

Integrated intrinsic colors of galaxies cover the range $B - V \sim +0.2$ to $+1.1$ and $U - B = -0.4$ to $+0.7$. Color is correlated with both Hubble morphological type (de Vaucouleurs, 1960) and with total luminosity within a given type (Baum, 1959; van den Bergh, 1971b) in the sense that irregulars are bluest and ellipticals reddest, and the lowest luminosity galaxies are the bluest of their types. The former effect represents the preponderance of old, low mass, red stars in E's and younger objects in S's and I's, and the latter effect the higher metal abundance in bright massive galaxies. There is considerable scatter about the mean for any given type. The mean relation of $B - V$, $U - B$ colors with Hubble type is shown in Fig. 16. The data is that of de Vaucouleurs (1960). The correlation of color with luminosity or mass can also be seen in spiral galaxies in the form of a tight relationship between $B - V$ and maximum rotational velocity (van den Bergh, 1971a). The range for ellipticals is from $B - V \sim 0.6$ (at $M_V = -8$ to -13) to $B - V \sim 0.9$ (at $M_V < -17$).

Many giant ellipticals display color gradients, being reddest at the center (Tifft, 1969). The color gradient in NGC 3379 (Miller and Prendergast, 1962) is frequently compared with the predictions of galactic evolution models. In the case of spirals, color varies with position as one would expect from the distribution of halo and disk populations in the galaxies. The extreme outer halo ($R \sim 25$ kiloparsec) of the one spiral (NGC 4565; Sb) which has been carefully studied has a redder ($\lambda\,9000 - \lambda\,5000$) color than either the nucleus of the galaxy or normal globular clusters (Spinrad, in NATO, 1974). A few gE's display anomalously blue colors (and sometimes early spectral type) in their nuclei (van den Bergh, 1972), apparently resulting from a recent burst of star formation. Galaxies with this property generally show other signs of activity in their nuclei. The presence of blue supergiants in the Ep/S0 galaxy NGC 205 (Baade, 1951) probably also indicates a small amount of recent star formation.

The color of a particular galaxy or galactic region depends both on the types of stars contributing the light (giants vs dwarfs; K vs M, etc.) and on their metal abundance. It is not, in general, possible to separate the two effects with *UBV* data alone. For instance, the red color of gE nuclei can be obtained from either a giant-dominated or a dwarf-dominated population derived from a reasonable distribution of stellar masses. Where spectroscopic data is available, it can help to resolve the ambiguities. We saw in Sec. II.F that the very strong CN bands and metal lines in some galactic nuclei can only be understood if an appreciable fraction of the stars there have metal abundances higher than solar by factors of 3 or more. Giant–dwarf discriminants also exist in the form of certain molecular bands. O'Connell (1974) concludes from the strength of the TiO bands in several elliptical and S0 galaxies that a significant fraction of their light comes from M giants. Frogel *et al.* (1975) and Whitford (1974) also conclude that giants are important in the nuclei of ellipticals. A similar conclusion can be reached for the nuclei of two Sb spirals from the CO and H_2O bands (Baldwin *et al.*, 1973). The luminosity of such giant-dominated regions is likely to vary considerably with time. Earlier work (quoted by O'Connell, 1974, and Baldwin *et al.*, 1973) had seemed to indicate that giants did not make an important contribution to the light from ellipical galaxies and spiral nuclei.

4. Masses

Masses for individual galaxies can be determined by four techniques: (a) measurement of rotation velocity as a function of radius (usually using the 21 cm line of HI and suitable only for gas-rich galaxies), (b) application of the virial theorem to the profiles of absorption lines in the integrated light of elliptical (gasless) galaxies, (c) measurements of the radial velocity and separation of pairs of galaxies which seem to be in binary orbits, and (d) application of the virial theorem to groups and clusters of galaxies. The first two techniques can be applied only to regions with significant luminosity (optical or radio), and there is no serious disagreement about the measurements or their interpretation over the bright parts of galaxies. The range of masses is quite broad, from dwarf galaxies of $10^5\,\mathfrak{M}_\odot$ (Hodge, 1971), which are not really distinguishable from globular clusters, to supergiant ellipticals, which may reach $10^{13}\,\mathfrak{M}_\odot$ (Arp and Bertola, 1971). On average, the sequence Ir, Sc, Sb, Sa, E is also a sequence of increasing mean mass, the means ranging from $\sim 10^{9.5}$ to $10^{11.5}\,\mathfrak{M}_\odot$, but there is a wide range within each type. Irregulars and spirals can easily vary by a factor of 10 either direction from their means and the elliptical class includes both the smallest dwarf and the largest supergiant galaxies. Within the part of the galaxy observed, the mass to light ratio (in solar units) is about 3 for irregulars, 5–15 for spirals, and 10–60 for ellipticals. This trend is what would be expected given the types of stellar populations that give most of the light from each type. In addition, M/L increases with distance from the center in those galaxies which have been carefully studied (M33 by Boulesteix and Monnet, 1970, and M31 by Roberts, 1975).

The amount of mass outside the optically bright regions of galaxies ($r \sim 20$ kiloparsecs for large systems) is much less well known, but there are several kinds of evidence that it may not be negligible. First, the curves of radial velocity vs radius are still quite flat at the largest radii ($\sim$30 kiloparsecs) observed (Roberts and Rots, 1973). This means that the outermost gas observed is still not following the Keplerian orbits ($V_{\rm rot} \propto R^{-\frac{1}{2}}$) that would indicate it is well outside most of the mass of the galaxy involved. Second, low luminosity halos have been directly observed well beyond the main part of the optical image of both elliptical (Arp and Bertola, 1971; Kormendy and Bahcall, 1974) and spiral (Spinrad, in NATO, 1974; Spinrad and Ostriker, 1974; Schild *et al.*, 1975) galaxies. If the light in these halos is largely due to stars, then its very red color suggests that the stars are low mass ones, so that M/L would continue to increase in these outer regions. Third, galactic masses measured by techniques (c) and (d) are nearly always much larger than those measured by (a) and (b), and the mass found increases about linearly with the effective radius out to which it is measured (Ostriker *et al.*, 1974; Einasto *et al.*, 1974). Thus, rotation curves give the smallest masses for spiral galaxies, the mass inferred from tidal effects on small companion galaxies is next, and masses derived from wide binary pairs and the virial theorem for groups and clusters are the largest. The implication is that, although out to 10–20 kiloparsecs the mass in a typical spiral halo is about equal to that in its disk, farther out there may be 10 times as much mass in the halo as in the inner, bright disk (Ostriker *et al.*, 1974). Fourth, in the absence of a massive halo, a flat, rotating gaseous disk is subject to several instabilities which, observationally, do not seem to occur (Ostriker and Peebles, 1973). Gaseous halos with this much mass can be excluded because they would contribute more x rays than we see from clusters of galaxies (Biermann and Silk, in NATO, 1974). The matter must be in the form of low luminosity stars. Within rich clusters, collisions of halos of neighboring galaxies can be expected to strip stars off (Gallagher and Ostriker, 1972), resulting in a population of intergalactic stars. Such encounters may also affect the structure of the cluster as a whole (Oemler, 1974). The evidence for these effects is not clearly present (Biermann and Silk, in NATO, 1974), but there is, at any rate, no evidence that they are not occurring. Typical galactic halos may then extend out to several hundred kiloparsecs and contain enough mass that clusters of galaxies are bound by them. The total M/L ratio is 200–300 for both spiral and elliptical galaxies under these circumstances. This increases the amount of mass in galaxies to a "cosmologically interesting" value (i.e., within striking distance of the amount needed to close the universe; Gott *et al.*, 1974). It is certain also to be important to how galaxies evolve. Burbidge (1975) strongly denies that the evidence requires or even allows the existence of such a halo for our own galaxy. We must evidently accept that all models of galactic evolution start with a mass of gas and stars which is uncertain by a factor of 10.

5. Gas mass and infall

Interstellar gas and dust is the raw material out of which new stars are formed, and stars are continuously returning matter to the interstellar medium. A knowledge of its composition, total mass, and distribution in various kinds of galaxies is, therefore, bound to be vital for an understanding of galactic evolution. We have discussed the composition of the interstellar medium in our own and other galaxies in Secs. II.D and II.F. The gas in our own galaxy presently available for forming new stars is characterized by a slightly higher metal abundance than the Sun ($Z \sim 0.02$–0.03), provided that the N, O, Ne, etc., in HII regions are assumed to be typical rather than the greatly depleted abundances of Ca, Fe, Ti, etc., in typical HI regions. The helium abundance is about 30% by mass. The average properties of other galaxies seem to be quite similar, except that the metal abundance can be much lower (down to $Z = 0.002$ or so) in galaxies of small total mass (see, e.g., Peimbert and Peimbert, 1974), and the helium abundance may be slightly lower ($Y \sim 0.23$) in low mass galaxies. The average ratio of dust to gas in our Galaxy is about $\rho_d/\rho_h = 0.5 \times 10^{-2}\rho_g$, where ρ_g is the density of a grain in g cm^{-3} (Aanestad and Purcell, 1973).

Temperature, densities, and distribution of both neutral and ionized hydrogen in galaxies have been studied by a large number of workers using many different techniques. Neutral regions are generally detected by 21 cm radio emission or by absorption lines in stellar spectra at optical or ultraviolet wavelengths, and ionized regions by optical or radio emission lines or thermal radio continuum. Two major reviews, Roberts (1975) on neutral gas and Burbidge and Burbidge (1975) on ionized gas, are part of the long-awaited Vol. 9 of the *Stars and Stellar Systems* compendium. Meanwhile summaries of much of the data are available in Roberts (1972), Davies (1972), and in a series of papers by the Burbidges and Prendergast in the *Astrophysical Journal* from 1959 to 1965. Kerr (1969) has discussed the details of the HI distribution in our own Galaxy.

Within our own Galaxy, the gas is strongly concentrated in spiral arms. When the distribution is smoothed over the arms, it becomes clear that the ionized hydrogen is rather strongly concentrated in a ring a couple of kiloparsecs wide, located about 5 kiloparsecs from the center of the Galaxy. The neutral hydrogen also has rather low abundance near the center ($n_H \lesssim 0.3$ cm^{-3}). The density gradually rises beyond about 4 kiloparsecs from the center to a broad, flat maximum ($n_H \sim 1$ cm^{-3}) from 7 to 11 kiloparsecs from the center. Beyond that, the density drops, but the thickness of the hydrogen disk increases, so that the amount of hydrogen per unit area of disk remains roughly constant. The neutral hydrogen appears to terminate about 15 kiloparsecs from the galactic center. Similar ringlike structure seems to be characteristic of most other spiral galaxies (see, e.g., Rots, 1974). Irregulars, on the other hand, have their neutral and ionized hydrogen concentrated toward their centers, although the precise 21 cm center is often displaced from the optical center in a way that cannot be entirely attributed to absorption by dust. The mass in neutral hydrogen is always considerably larger than that in ionized hydrogen for spiral and irregular galaxies. The neutral hydrogen diameter of a Galaxy is typically somewhat larger than its optical (Holmberg) diameter, and the ratio increases along the sequence from Sb to Ir types. Gas in molecular form (mostly H_2) may make a significant contribution to the total. For our Galaxy, Scoville and Solomon (1974; Solomon and Stecker, 1974; and Solomon, in NATO, 1974) have identified major concentrations of CO, which imply gas in molecular form near the galactic center and in a ring 4.6 kiloparsecs from the center whose total mass is probably comparable with that of atomic gas. Almost nothing is known about molecular gas in other galaxies, except that Andromeda could not have as strong CO emission from its center as our galaxy does without its having been detected (Solomon, in NATO, 1974). Strong CO emission has recently been detected in several other galaxies (Solomon, private communication, 1975).

The abundance of gas in detectable forms in elliptical galaxies is surprisingly low. About 15% of giant ellipticals show optical emission lines near their centers, which can be attributed to $10^{4-5}\,\mathfrak{M}_\odot$ of ionized gas at a typical HII region temperature (Osterbrock, 1962). A comparable amount of ionized gas appears to be required to produce the Faraday depolarization of the radio emission from some normal elliptical galaxies (Wardle and Sramek, 1974). Neutral hydrogen makes up less than about 0.5% of the mass of all ellipticals that have been studied (see, e.g., Knapp and Kerr, 1974, on NGC 4472). This is less than would be expected from most models of galactic evolution, inless the gas lost by the red giants in rather efficiently hidden by being turned back into stars, swallowed by black holes, swept out of the galaxy, or being heated or cooled to invisibility. Several of these are sufficiently possible that the problem is presently an interesting one rather than a serious one for galactic evolution (F. J. Kerr, 1974, and B. Paczyński, 1974, private communications).

Among the spiral and irregular galaxies, there are several correlations of gas (i.e., HI) mass (as tabulated, e.g., by Balkowski, 1973) with other properties. The ratios of HI mass to total mass and HI mass to total luminosity both increase monotonically along the sequence from Sa to Ir. Since Hubble type is well correlated with color, there is also a monotonic increase of gas fraction with increasing blueness. Finally, the total HI mass increases with the total luminosity of a galaxy, the average relationship being slightly different for different Hubble types. The four relationships are all roughly linear, and their ranges are: (a) $(M_{\rm HI}/\mathfrak{M}_\odot)/(L/L_\odot) = 0.2$ at Sa to 0.8 at Ir, (b) $(M_{\rm HI}/M_{\rm total}) = 0.03$ at Sa to 0.22 at Ir (up to 0.5 in the Small Magellanic Cloud), (c) $(M_{\rm HI}/M_{\rm total}) = 0.001$ at $B - V = +0.9$ to 0.5 at $B - V = -0.2$, and (d) $M_{\rm HI} = 8 \times 10^7\,\mathfrak{M}_\odot$ at $L = 10^8\,L_\odot$ to $10^{10}\,\mathfrak{M}_\odot$ at $L = 2 \times 10^{10}\,L_\odot$. In addition, within a given Hubble type, the ratio of gas mass to total mass tends to be highest in low luminosity galaxies (Balkowski *et al.*, 1974, and references therein). No galaxies are known whose total HI mass is greater than about $10^{10}\,\mathfrak{M}_\odot$. This may reflect the formation of molecules, regions of high optical depth, and stars when the total gas mass becomes greater than this. It is not presently known how the various correlations will be affected by including the mass of the gas that is in molecular form.

Finally, it is important for any theory of galactic evolution to know whether galaxies are closed systems, and matter is most likely to enter or leave galaxies in the form of gas. We have already alluded to the possibility of gas lost by stars being swept out of elliptical galaxies, and the low metal abundance of dwarf elliptical galaxies is apparently attributable to their having lost all their remaining gas after an initial burst of star formation (Larson, 1974b), possibly due to kinetic energy and heating supplied by the supernovae in that first stellar generation. Thus, mass loss from galaxies probably occurs. Significant mass gain has also been suggested. Oort (1970) pointed out that the properties of the high velocity HI clouds observed at high galactic latitudes were consistent with their being intergalactic matter in the process of falling down into the galactic plane. The gas would presumably consist of only H and He in their primordial proportions. Such infall would augment the local mass density by only 1%–2% per 10^9 yr, but since the material is all in the form of gas, it could influence the composition of the interstellar medium considerably. Infall was included in the galactic evolution models of Quirk and Tinsley (1973), Fowler (1972), Audouze and Tinsley (1974), and Biermann and Tinsley (1974) and was important in keeping the interstellar metal abundance roughly constant over the past few billion years and in maintaining the observed interstellar deuterium abundance. Although infall is supported by the work of Hulsbosch and Oort (1975) and Cohen (1975), it now seems at least as probable (Davies, 1972a; Verschuur, 1973) that the high velocity clouds represent extensions of the galactic spiral arms out of the galactic plane. There is, therefore, no strong evidence at present for addition of mass to the Galaxy, except its usefulness in making the models come out right.

6. Star formation rates

Since the fundamental process of galactic evolution is the interchange of matter between the insterstellar medium and stars, it is clearly vital to know the rate at which gas is being turned into stars as a function of time and place. Our present theory of star formation is not merely inadequate, but virtually nonexistent. We do not know how clouds unstable to gravitational collapse ($n_H = 10^{4-6}$ cm^{-3}, $T = 10$–30°K) arise or how a cloud decides what size fragments to break into or how these things depend on the average composition, density, or temperature of the interstellar medium. The efforts to date to arrive at a theoretical understanding of these problems have been reviewed by Larson (1973, 1974a). The largely observational discussion of star formation rates which follows is, therefore, at most occasionally illuminated or shadowed by theoretical insights.

The quantity desired is the number of stars of each mass and composition born out of a given amount of gas as a function of time. This quantity can conveniently be called $b(m,t,Z)$. We will return to the question of whether or not it is separable into $b(m,t,Z) = \psi(m)F(t)Q(Z)$. The total average rate of star formation over the history of our Galaxy is clearly $\sim 10^{11}$ stars in $\sim 10^{10}$ yr, or about 10 stars/yr. This average rate will be about the same for any galaxy of the same mass and age that has most of its mass in the form of stars. Thus, it cannot be very important in determining the course of galactic evolution. The rate at particular times is more interesting. In our own Galaxy, it appears that about $1\,\mathfrak{M}_\odot$/yr is going into new stars at the present time. This is considerably less than the average rate, but also not zero. The current star formation rate can be estimated from the total HII luminosity of the Galaxy (Mezger, in NATO, 1974). Studies of the Orion region tell us that the ratio of number of Lyman continuum photons emitted by new stars (and available to ionize HII) to the amount of mass that has gone into stars is about 2.2×10^{46} photons per solar mass. The totality of HII emission in the region 4–15 kiloparsecs from the galactic center tells us that 7.5×10^{52} photons have been required to ionize the gas. Thus there must be $3 \times 10^6\,\mathfrak{M}_\odot$ in stars formed recently enough to contribute to the ionization. Now the peak of Ly continuum emission occurs at about spectral type 06 (balancing the stars' temperatures with their relative abundance), and an 06 star lives about 3×10^6 yr. Thus, on average, $1\,\mathfrak{M}_\odot$/yr must have been turned into stars in our Galaxy in the recent past, assuming that stars are still being born with the distribution of masses that they have in the solar neighborhood. Since gas passes through the spiral arms, and thus has the opportunity to be compressed and make stars, at a rate of 50–100 $\mathfrak{M}_\odot$/yr (that is, all the gas every 2×10^8 yr), the efficiency of star formation is evidently only 1%–2%. Many of the observed differences among Hubble types are attributable to different rates of star formation at the present time. The irregulars are still forming stars about as rapidly as they ever have, while star formation in ellipticals virtually stopped after the first billion years or so. This point will be further discussed in connection with various galactic evolution models.

The composition dependence of the birthrate can be easily separated off from the mass and time dependences if the composition of stars being formed bears a simple relationship to the average composition of the gas in their vicinity at the time. The simplest possible relationship is identity of the two compositions, and many of the models have made this assumption. Both the other possibilities have also been discussed. Edmunds and Wickramasinghe (1974) have considered the expulsion by radiation pressure of metal-rich grains from collapsing protostars and concluded that massive stars could be significantly depleted in metals relative to the gas from which they formed. This possibility has not yet been incorporated into a model of galactic evolution. The alternative has been treated by Talbot and Arnett (1973; Talbot, 1974) under the title Metal Enhanced Star Formation (MESF). The motivation is that, by assuming stars are only formed where the metal abundance in the gas is higher than average, it is possible to account for the very small number of low Z stars in the solar neighborhood. The theoretical justification is that, in a two-component model of the interstellar medium, only the high-density, low temperature phase can exist above a certain critical pressure. Since the value of this critical pressure varies inversely with Z, the most metal-rich regions of the interstellar medium will collapse and form stars most readily. Variations of 20%–30% in metal abundance in the gas suffice to make the effect an important one. It has been included in the models of galactic evolution calculated by Talbot and Arnett (1973, 1974, 1975). Differences in metal abundance of this magnitude between young stars and HII regions or between HII regions are not presently detectable. Either very accurate observations or a much-improved theory of star formation will be necessary to decide whether the effect actually occurs. In any of the three cases ($Z_* = Z_{gas}$; $Z_* > Z_{gas}$; $Z_* < Z_{gas}$) birthrate as a function of composition can be separated from birthrate as a function of mass and time in a simple way, and the change in the average composition of the interstellar gas caused by forming stars can be calculated.

Separation of the mass and time dependences is more complicated. If the fraction of star-forming material that goes into stars of each mass is the same under all conditions, then $b(m,t)$ is separable as $b(m,t) = \psi(m)F(t)$, where the number of stars formed per unit time with masses between m and $m + dm$ is $b(m,t)dm$ and the normalization is given by

$$\int_{m_L}^{m_U} \psi(m)dm = 1,$$

m_L and m_U being the lower and upper limits to the masses of stars formed. Then all the model builder has to do is discover what $\psi_m(m)$ is here and now and try various forms of $F(t)$ to match observed properties of galaxies. This is clearly the simplest possible assumption and was used in the first models (van den Bergh, 1962; Schmidt, 1963). If applied to a completely homogeneous closed galaxy, this assumption of constant Initial Mass Function (IMF) leads to the production of many more low mass stars with low metal abundance than we see (Fig. 5; the calculations are discussed in Sec. IV.B). Schmidt's (1963) solution, also adopted by Truran and Cameron (1971) and by Quirk and Tinsley (1973), was a variable IMF, in which the first generation of stars contained only massive stars. Such models do not contradict other observed properties of galaxies, provided that most of the metals produced are retained in massive black holes (Truran and Cameron, 1971; Talbot, 1973). We have already seen that infall of $Z = 0$ gas and MESF can provide alternative solutions to the problem of the small number of stars with low metal abundance. Considerable improvement in the theory of both stellar and galactic evolution will be needed before models alone will be able to tell us whether the IMF is, in fact, a constant. At present, the work of Searle *et al.* (1973) favors, and that of Sargent and Tinsley (1974) opposes, constant IMF in elliptical galaxies.

Several observations tell us something about possible variations in the IMF. It has long been known (van den Bergh and Sher, 1960) that open clusters have fewer low mass stars than the general field. Since the cluster stars are, on average, younger than the field ones, this might be evidence of a time-variable IMF, though in the opposite direction from the variation suggested for galactic evolution models. Alternatively, the field-cluster difference may mean that the IMF depends on, e.g., the density or size of the collapsing interstellar cloud which forms the stars, but in a way that does not change systematically with time. Infrared astronomy also provided some relevant data. Within the Orion region, there are two clusters of infrared sources, each surrounded by a dense molecular cloud. The densities of the two clouds differ by a factor of about 10, and the individual infrared sources in the less dense cloud are the fainter by a factor of about 100. Since the sources are probably young stars or protostars, the implication is that low-density clouds form lower mass stars than high-density clouds (unpublished work by M. Werner, quoted by Larson in NATO, 1974). Again, it is not clear whether this implies any systematic variation with time. A study of the brightest stars in galaxies of different brightnesses provides evidence for a constant Initial Luminosity Function, and so, by implication, constant IMF (Sandage and Tammann, 1974). They point out that the positive correlation of brightest star brightness with total galaxy brightness has just the form that would be expected if each galaxy studied is sampling the same ILF and the most populous galaxies have the best chance of having at least one really bright star. Thus, the more stars a galaxy has, the brighter its brightest one is likely to be. Counts of stars of various ages and metal abundances also provide some support for a roughly constant IMF in the solar neighborhood (Bond, 1970; Clegg and Bell, 1973; Wielen, 1973). Reeves (in NATO, 1974) has used the results of nucleocosmochronology to deduce the approximate constancy of the IMF over the time during which Solar System radioactive nuclides were being produced. He points out that, in a simple model of galactic evolution (for a closed system which remains homogeneous and has the total gas mass dropping exponentially with time), the quantity X_{ji} defined below should be equal to unity, provided that the IMF has been constant over the time the elements were being made:

$$X_{ji} = \left(\frac{n_j}{n_i}\right)\left(\frac{f_i}{f_j}\right)\left(\frac{\tau_j}{\tau_i}\right),$$

where n_j and n_i are the amounts of the two radioactive nuclides j and i in the gas at a given time, τ_j and τ_i are their lifetimes, and f_j and f_i are the fractions of gas going into stars in one generation which are returned as those nuclides. Reeves finds that the ratios $^{244}Pu/^{235}U$, $^{235}U/^{238}U$, and $^{232}Th/^{244}Pu$ at the time the meteorites began to retain xenon all had values consistent with $X_{ji} = 1$, implying a roughly constant IMF over time scales of several billion years before the formation of the Solar System. The value of X_{ji} for the pair $^{129}I/^{127}I$ is much less than unity, implying that very little synthesis occurred in the 10^8 yr immediately preceding the formation of the Solar System. This is roughly the time between the passage of successive spiral arms through a given volume of gas in the density wave theory of spiral structure. Thus, there is observational evidence both for and against constancy of the Initial Mass Function.

Existing theory is similarly unable to decide definitely whether the IMF should vary or not. On the one hand, the Jeans mass (the smallest mass unstable to collapse by self-gravitation) decreases as gas density and metal abundance increase because both of these help cooling processes to keep up with the heating caused by collapse. On the other hand, calculations of how a cloud should fragment (Larson, 1973; Lynden-Bell, in NATO, 1974) suggest that the initial stellar core is much smaller than the star that will eventually form. The final star mass is then determined by the amount of accretion that occurs before radiation pressure blows away the remaining gas. The final phases of the accretion process for massive stars have been studied by Kahn (1974). His results generally confirm the *a priori* expectation that, if star formation is dominated by accretion, a high-density cloud will lead to massive stars, while due to the effects of dust (whose abundance depends on Z) on radiation pressure, a high metal abundance will result in lower mass stars, on the average. Larson and Starrfield (1971) have also considered the problem of how IMF should depend on gas density and Z, coming to slightly different conclusions, as do Appenzeller and Tscharnuter (1974). Reddish (1975) obtains yet another result.

Thus, theory cannot presently tell us whether the IMF should vary with time or place, or even what its form should be at the present time. The observational evidence on possible variations in mass function are similarly ambiguous. It is, however, possible to determine the Initial Mass Function in the recent past for the solar neighborhood in a relatively straightforward fashion. "All" we have to do is: (a) count the stars in a given volume of space as a function of absolute magnitude, (b) allow for incompleteness of the counts as a function of magnitude (especially important at the faint end of the distribution), (c) convert absolute visual magnitudes to masses, using stellar models, and (d) divide the number of stars in each mass interval by the lifetime of stars with that mass (derived from evolutionary calculations). This will give the average number of stars of a given mass which formed in the solar neighborhood per unit time, over a time interval which is, unfortunately, a steep function of mass. The "observed" IMF which is most often used is that of Salpeter (1955), who found that $\psi(m) \propto m^{-2.35}$ between some rather loosely determined upper and lower mass limits. The results of Limber (1960), Hartman (1970), and Larson (1973a) can be more conveniently expressed in terms of $\theta(M)$, the fraction of total mass that goes into stars per unit logarithmic interval in star mass. $\theta(M)$ has roughly the shape of a Gaussian, centered at $1.4\,\mathfrak{M}_\odot$ with a standard deviation of 0.66 in $\log_{10}(M)$. The two forms of the distribution are, of course, equivalent over the "well-observed" mass range, 0.6–10 $\mathfrak{M}_\odot$, but can differ quite considerably at the ends, depending on the choice of m_L and m_U. Both extrema of the IMF are of some importance, since the very massive stars contribute most of the enrichment in elements beyond carbon and oxygen, while the fraction of mass going into stars below $1\,\mathfrak{M}_\odot$ determines how efficient heavy-element production has to be in order to produce the present Z without using up all the gas. The upper and lower limits of observed stellar masses are about 60 and $0.06\,\mathfrak{M}_\odot$. The fraction of mass that goes into stars above $1\,\mathfrak{M}_\odot$ is uncertain by a factor of about 2, being about 25% for the van Rhijn (1936) luminosity function, but only 12% if the large number of M dwarfs implied by the work of Weistrop (1972) and Murray and Sanduleak (1972) is a universal phenomenon. These problems are frequently handled by the model builders by including m_L, m_U, and the slope of $\psi(m)$ as parameters to be varied over a family of models.

The time dependence of the birthrate function, $F(t)$, is normally taken to be a declining exponential, $\exp(-t/\tau)$, where the time constant, τ, can vary from less than 10^9 yr for a massive galactic halo (Ostriker and Thuan, 1975) to at least the age of the universe for the bluest irregular galaxies (Larson and Tinsley, 1974). A time dependence of this general sort results more or less automatically if the star formation rate is presumed to vary directly with the amount of gas remaining to be turned into stars (that is, its average density). Some such dependence of star formation rate on gas density is implicit in most of the evolutionary calculations from that of Schmidt (1963) to the present, although the observational evidence for it is not very strong (Black and Kellman, 1974). Finally, it seems probable that star formation in fact occurs in spurts, moving from one place to another in a galaxy, and that it is only an average formation rate that varies smoothly with time. The evidence for this is clearest in small galaxies, where there is less smearing out of the events because there are fewer of them. Van den Bergh (1975, and quoted by E. M. Burbidge in NATO, 1974) has shown that the site of vigorous star formation has apparently moved around with time in the Small Magellanic Cloud. There is similar evidence for the Large Magellanic Cloud (Hodge, 1973; Baird *et al.*, 1974). In addition, Balkowski *et al.* (1974) suggest that the relatively common dwarf irregular galaxies are a quiescent phase, occasionally interrupted by rapid bursts of star formation, whose result is to produce a compact blue galaxy of the type studied by Searle and Sargent (1972). The confinement of young stars to the arms of spiral galaxies of course also implies some irregularity in star formation rates, as does the very blue $U-B$ color of some spirals with unusually high supernova rates.

7. Supernova rates

The observed supernova rate in our own and other galaxies constrains evolutionary models in two ways. First, it says how many events per year can now be contributing heavy elements to the interstellar medium. Second, if we only knew the ranges of main sequence masses which give rise to supernovae of various types, we would know the birthrate for these mass ranges in galaxies of various types, and thus have an important constraint on variations between galaxies in the Initial Mass Function. A comprehensive summary of observed supernova rates as a function of galaxy mass and Hubble type has been given by Tammann (1974). Table XI is derived from his data. Supernovae of unknown type were assigned by him to Types I and II in proportion as the identified types occur in the various galaxy classes. The two most striking features are the confinement of Type II supernovae to Sb and Sc galaxies and the nearly monotonic increase of both total rate and Type I rate from elliptical through spiral to irregular galaxies.

If our own Galaxy is a $1.7\times10^{11}\,\mathfrak{M}_\odot$ type Sb or Sbc galaxy, it should, on the basis of the data in Table XI

TABLE XI. Supernova rates per unit mass in galaxies as a function of Hubble Type (data from Tammann, 1974).

SNe per 100 yr per 10^{10} $M_\odot$	E/S0	Sa	Sb	Sc	IrI
Type I	0.007	0.08	0.08	0.29	0.38
Type II and other	0.0	0.0	0.11	0.28	0.0
Total	0.007	0.08	0.19 ± 0.10	0.57 ± 0.13	0.38 ± 0.19

have one supernova every 15 to 60 yr, about equally divided between Types I and II. The number of supernovae observed in the past 1000 years implies a very similar rate, given that they were all rather close to the Sun by galactic standards (e.g., 4 SN within 4.5 kiloparsecs in 1000 yr implies 23 yr between events throughout the Galaxy). Two other semi-observable quantities, the rates of formation of radio-emitting supernova remants (SNRs) and of pulsars, tell us something about the galactic supernova rate. These rates are even more uncertain, since they involve not only an estimation of the completeness of the counts, but a model-dependent estimate of ages of individual objects. Ilovaisky and Lequeux (1972) find that the mean interval between SNR births is probably in the range 25–75 yr, while the pulsar birthrate may be one per 30 yr (Ostriker and Gunn, 1970) or only one per 100–200 yr (work by A. Hewish quoted by Tammann, 1974). These rates seem to be somewhat lower than the expected total supernova rate, but are sufficiently uncertain to be consistent with SNRs and pulsars being formed by all supernovae, or only Type II supernovae, or only some Type II supernovae, or any other combination that appeals to the reader.

The total rate of supernova events in galaxies as a function of Hubble type is, therefore, known with at least moderate accuracy. Assignment of the various types of supernovae to stars of particular masses is much more difficult. The conventional association is Type I SNs with low mass stars and Type II's (and other less common types) with high mass stars on the basis of the occurrence of the former in elliptical galaxies (assumed to contain only old stars) and the latter in spiral arms. (SNe I's in spiral galaxies are not concentrated in the arms. S. van den Bergh, private communication, 1975.) What we mean by "massive" in this context was discussed in Sec. III.D.1 in connection with explosive carbon detonation. It appears to mean "all or most stars above a lower mass limit which falls in the range 4–8 $\mathfrak{M}_\odot$ and probably depends on composition, angular momentum, and so forth." Models of galactic evolution which are otherwise viable always have enough massive stars in the Sb and Sc galaxies to produce the required SNII rate. The difficulties of getting a major explosion out of a star of 1 $\mathfrak{M}_\odot$ or less have provided employment for a large number of theorists. The models generally involve either close binary systems (Hartwick, 1972; Whelan and Iben, 1973) or instabilities in white dwarfs (Finzi and Wolf, 1967; Hansen and Wheeler, 1969; Ostriker, 1971). Several of the theories are probably capable of yielding the required rates of SNIs, but none makes it clear why the old population in spiral and irregular galaxies should be so much more efficient a producer of Type I supernovae than that in elliptical galaxies. This and about 10 other difficulties with the conventional mass assignments, summarized by Tammann (1974), have led him and others to suggest that all supernovae may in fact descend from massive, young stars, the distinction (which may not be sharp) between Type I and Type II objects then being attributable to some less fundamental difference in a precursor than a large mass difference. If this is the case, then elliptical galaxies must be producing young stars at 1%–10% of the rate in spiral and irregular galaxies. This possibility cannot be excluded. First, we see evidence for blue (presumably young) stars in some elliptical and S0 galaxies (Baade, 1951; van den Bergh, 1972), including some which both produce supernovae and show other evidence of activity. Second, at least some models of the evolution of elliptical galaxies which reproduce their observed colors and other properties include enough star formation at the present time to account for the elliptical supernova rate (Larson and Tinsley, 1974). The intermediate possibility, two classes of Type I supernovae arising from old and young stars, has been discussed by Barbon *et al.* (1973) and Dallaporta (1973). This does not meet all the objections discussed by Tammann (1974) to low mass precursors for Type I SNe. The masses of supernova precursors will be further discussed by Arnett and Tinsley (1975).

Thus, we can regard as a possible but not certain constraint on models of galactic evolution the need to have a present rate of formation of "massive" stars sufficient to account for the total observed supernova rate. It must be kept in mind that this observed supernova rate, as derived by Tammann (1974), is a lower limit, since it includes no correction for incompleteness of searches for most galaxy types. In addition, it applies only at the present time and to whole galaxies. It cannot be expected that particular parts of galaxies will comply with the over-all rate or that it will not have changed greatly with time. It is, however, a plausible assumption that galaxies of a given type that agree now in the SN rate per unit luminosity (or per unit mass) should also have had the same rates in the past.

B. Homogeneous one-zone models with constant IMF

The object of models of galactic evolution is to start with a glob of gas of known mass, composition, and density, adopt some law of star formation (which may depend on the properties of the gas), feed in knowledge of the evolution of stars as a function of mass, and, thereby to predict as a function of time observable properties of galaxies including total luminosity and luminosity function (i.e., the distribution of stellar luminosities), color, mass remaining in gas, the distribution of stellar ages, the distribution of stellar chemical compositions, the ages of the chemical elements, gradients in gas density, color, and metal abundances, the supernova rate, and so forth. Many of these properties can be observed only in our own Galaxy (e.g., luminosity function, the distribution of stellar compositions, ages of the elements, etc.) and a few (e.g., integrated color) only in external galaxies. There are at present no good reviews of the field of galactic evolution, although one is contemplated (Andouze and Tinsley, 1976).

The first step in carrying out this program was taken by Salpeter (1955), who derived an expression for the birthrate of stars in the solar neighborhood as a function of mass. He found that the number of stars born (e.g., per year) per unit interval in mass, $\psi(m)$, scaled as $m^{-2.35}$ (or equivalently, the number born per unit logarithmic interval in mass scales as $m^{-1.35}$). His Initial Mass Function and ones like it are still commonly used in models of galactic evolution. The next stage of refinement, allowing the over-all birthrate to vary with time, but keeping the IMF, was introduced by van den Bergh (1957) and Schmidt (1959). Many of the properties of such a model can be derived in a simple, analytic way (see, e.g., Searle, 1973; Pagel, 1973b), while others require keeping track of enough properties of stars in enough different mass ranges that a computer must be added to the list of authors.

In particular, the chemical evolution of such a galaxy can be calculated analytically. Suppose, in some volume of gas (which may be an entire galaxy or a neighborhood which is known to remain isolated), that the rate of conversion of mass into stars, dS/dt, is proportional to some power of the gas density. Then, since the density decreases linearly with the fraction μ of the matter still in the form of gas

$$dS/dt = A\mu^n.$$

The mass fraction as a function of time will be affected both by star formation and by the return of gas from evolved stars. In the "instant recycling" approximation, the return is assumed to happen quickly, compared to other changes of interest. Then if b is the fraction of matter that goes into stars which eventually comes out again, the rate of change of μ is given by

$$d\mu/dt = -(1-b)(dS/dt) = -(1-b)A\mu^n.$$

Under these circumstances, b depends only upon the IMF and lies in the range 0.12 to 0.25, depending upon the number of late M dwarfs in the IMF, for a Salpeter-type birthrate function. If we let $A(1-b) = \tau$, some characteristic time scale, then

$$d\mu/dt = -\mu^n/\tau,$$

which has various analytic solutions, depending on the value of n:

if $n = 0$, $\mu = 1 - t/\tau$,
if $n = 1$, $\mu(t) = e^{-t/\tau}$,
if $n = 2$, $\mu(t) = 1/(1 + t/\tau)$,
etc.

From observational arguments, Schmidt (1963) concluded that n was between 1 and 2.

Now, suppose, in addition, that the interstellar medium always remains well mixed. Then $Z(t)$ will be well defined, and newly forming stars will always use up primitive and processed gas in the proportions in which they occur. In addition, for a constant IMF, we can define a unique quantity, called y, the yield, which is the fraction of mass going into stars per unit time which will eventually come out again in the form of heavy elements. Its value depends on the IMF and on the details of stellar evolution. Using reasonable birthrate functions and data like that in Table X, it is in the range 0.003–0.013. Using this concept, we can write down a second differential equation, for the change with time of the interstellar metal fraction, μZ,

$$\frac{d}{dt}(\mu Z) = Z\frac{d\mu}{dt} - y\frac{d\mu}{dt}(1-Z),$$

where the first term represents loss due to star formation and the second, gain due to mass ejection from massive stars. It reduces to

$$\frac{dZ}{dt} = -\frac{y}{\mu}\frac{d\mu}{dt} \quad \text{or} \quad \frac{dZ}{dt} = \frac{y}{\tau}\mu^{n-1}.$$

In the first form, there is an approximate analytic solution,

$$Z = y\ln(1/\mu).$$

Using this, we can immediately account for the systematic variations of metal abundance with Hubble type and the gradients in Z across galaxies. The highest metal abundances are to be found in the regions where the residual gas abundance is the lowest, in good accord with the observations. And, as along as μ remains in the range 0.005 to 0.5 (roughly the observed range), Z will only vary from $0.7y$ to $5.3y$, thus accounting for the relative uniformity in metal abundance among normal galaxies of a wide variety of types. Notice that these results do not depend on dS/dt having any particular relationship to time or μ, and the time dependence cancels out of the differential equation.

The time variation of Z follows also from the dependence of dZ/dt on μ and is linear with time for $n=1$. For larger n, Z increases rapidly with time, but flattens off as μ drops toward zero. This behavior is in reasonably good accord with the observed behavior of $Z(t)$ shown in Fig. 4. So far, the simplest conceivable model is doing very well. In addition, computerized versions of it (Talbot and Arnett, 1971; Fowler, 1972; Searle *et al.*, 1973; Tinsley, 1968, and in NATO, 1974) can account for an optional epoch in the past when Z(gas) was higher than it is now, before being diluted by unprocessed material being expelled from evolving low mass stars (allowing for the possible existence of super metal rich stars), the variations of $U-B$, $B-V$ color along the Hubble sequence, the mass to luminosity ratio and its relation to color along the Hubble sequence, the relationship of gas fraction and color; the detailed spectral energy distributions of galaxies as derived from scanner observations, and supernova rates, all by varying the time scale, τ, for star formation from small values for elliptical galaxies to large ones for irregulars. Why the time scale should have varied from place to place in the required way is evidently a problem in the dynamics of galaxy formation, to be accounted for by a more complex theory. Such a dynamical theory should be able to explain the reason for the existence of abundance and color gradients, the relationship of total mass and Z, the variation of star formation rate along the Hubble sequence of galaxy types, and the differences among halo, disk, and nuclear populations in spiral galaxies. We will return to the properties of dynamical models in the next section.

The one-zone, homogeneous model of galactic evolution makes one further prediction, the number of stars as a function of metal abundance, which fails completely to agree with observations. Consider a class of stars (e.g., G dwarfs) all of whose members, formed since the birth of the Galaxy, should still be around and all of which have about the same mass. Let $N(Z)$ be the number of such stars whose metal abundance is equal to or less than some value Z. Under these assumptions, $N(Z)$ will be proportional to the total mass of stars formed in the

Galaxy (or region) up to the time when the metal abundance reaches Z. Then, if Z_1 is the metal abundance now and μ_1 is the gas fraction now, it immediately follows from the dependence of Z and μ that

$$\frac{N(Z)}{N(Z_1)} = \frac{1-\mu}{1-\mu_1},$$

and, in addition, $Z/Z_1 = (\ln\mu)/(\ln\mu_1)$, or, combining the two equations,

$$\frac{N(Z)}{N(Z_1)} = \frac{1-\mu_1{}^{Z/Z_1}}{1-\mu_1}.$$

The present day value of μ_1 in the solar neighborhood is not precisely known, but is in the range 0.05–0.15. A typical predicted curve of $N(Z)/N(Z_1)$ for an intermediate value of μ_1 is shown in Fig. 5. It clearly predicts far too many stars of low metal abundance.

This is a fatal objection to the simple-minded model of evolution of the solar neighborhood. Although we have almost no information about $N(Z)$ elsewhere in the universe, it seems improbable that the solar neighborhood should be unique in having an $N(Z)$ which cannot be accounted for by one-zone, homogeneous models. Thus, builders of model galaxies have been forced to abandon one or more of the basic assumptions that lead to the simple model. These assumptions are (a) constancy of the Initial Mass Function, (b) homogeneity of the Galaxy or region considered, and (c) absence of flow of matter in or out of the Galaxy or region considered. Models have been built which violate each of these assumptions and most of the possible combinations.

C. Models with variable IMF

The work of van den Bergh (1962) and Schmidt (1963) made it clear that something must be wrong with the simple model since it predicted too many low-metal stars. In addition, the work of Eggen *et al.* (1962) and Sandage *et al.* (1970) in explaining the gradients in metal abundance in our own Galaxy and the differences between elliptical and spiral galaxies in terms of gas flow from halo to disk made it clear that the solar neighborhood could not really be considered as a closed system. Schmidt's solution to the problem with $N(Z)$ was the adoption of a variable Initial Mass Function, in which earlier generations of stars put more of their mass into massive stars, so that metal enrichment proceeded rapidly and very few low mass, low-metal abundance stars would be left over from that period. This is equivalent to replacing a constant yield by a variable one,

$$y = y_0\mu^q,$$

hence

$$\frac{dZ}{dt} = -y_0\mu^{(q-1)}\frac{d\mu}{dt}$$

and

$$Z = y_0(1-\mu^q)/q.$$

A value of q near 1.5 provides a reasonably good fit to the data of Fig. 5. There is, however, a major difficulty. We know that y_1 (the yield at the present time) is about 0.01 (from the present IMF and stellar evolution theory); thus, if $q = 1.5$ and the present gas fraction μ is at most 0.15, then y_0, the initial yield, must have been at least $(0.01)/(0.15)^{1.5} = 0.17$. Hence, the equation for Z in terms of y_0, μ, and q implies that the present metal abundance Z_1 should be about 0.17/1.5 or more than 10%. This is an embarrassing overproduction of heavy elements. Truran and Cameron (1971) in their variable IMF model for galactic evolution concealed the extra metals in black hole remnants of massive stars. We cannot exclude this possibility on the basis of present stellar evolution theory or galactic dynamics. The invisible black holes conveniently supply a portion of the mass in the solar neighborhood which is known to be there from studies of stellar orbits, but is not seen in the form of normal stars and gas (Oort, 1965).

Failing this wholesale concealment of heavy elements, it has been claimed (Thuan *et al.*, 1975) that no homogeneous, closed model, no matter what form is chosen for birthrate as a function of time, is capable of reproducing the observed properties of the solar neighborhood. Thuan *et al.* (1975) parametrize the birthrate as a function $B(M,t)$ such that $B(M,t)\ dMdt$ is the number of stars of mass M formed in the interval dM at time t in the interval dt per square parsec by unknown processes from the existing stock of gas. It has the form $B(M,t) = B_0(M) \times \exp(-t/\tau_M)$, where $\tau_M = (A + B\log M)^{-1}$ and there are no restrictions on the signs or sizes of A and B. Using this birthrate function and the physics of stellar evolution, an initial mass of gas (the solar neighborhood value of $M_{\rm total}$ is about 75 $\mathfrak{M}_\odot$ parsec^{-2}; Oort, 1965) can be evolved forward in time. The models are stopped at a time equal to the present age of the Galaxy. The three unknowns, $B_0(M)$, A, and B, are then uniquely determined by the requirement that the model match present observed values of (a) the local luminosity function, (b) metal abundance in the gas, and (c) remaining gas fraction. This yields a "standard model" with a considerable allowable range around it due to uncertainties in the input data (age of the galaxy, initial mass, stellar evolution physics, etc.). A variety of tests are then applied to the standard model and its allowable variants. It meets many of these well. The right mass/luminosity ratio is automatically produced, given that the observed luminosity function has been matched. Correct birthrates for white dwarfs and supernovae are also found when reasonable ranges of masses are assumed for the progenitors. The average age of the metals in the Solar System is also correct, to within the undertainties of the determination by nucleocosmochronology (see Sec. II.B.2). A sufficiently large fraction of the remaining gas has never been through stars in the standard model that there is no difficulty in interpreting the observed interstellar deuterium abundance as a relic of the Big Bang (provided we live in a low-density universe; see Sec. III.A). It is, however, found that throughout the range of otherwise allowable models one of two serious difficulties occurs. Either there are far too many low mass, low Z stars (the standard problem), or the rate of formation of low mass stars is rapidly increasing at the present time and all the gas is about to be converted to low mass stars. In the latter case, we live in an exceedingly unusual time. Since most other spiral galaxies also have about 10% residual gas and M/L ratios and Z values similar to those in our Galaxy, they would presumably all also be about to turn the last of their gas into stars. This seems implausible. It is, therefore, necessary to reject the entire class of homogeneous, closed system models with this parametrization of $B(M,t)$. Tinsley (private communication, 1975) and Johns and Reeves (private communication, 1975) find that closed homogeneous models are possible if greater freedom is allowed in the birthrate function (for instance, one in which there is a sharp lower mass cutoff which varies with time).

The most serious objection to these homogeneous models is, of course, that they are bound to be a tremendous oversimplification of the real state of affairs.

D. Models with infall or inhomogeneity

Two other modifications of the standard, local model are possible. We can abandon either the "closed system" assumption or the "well mixed" assumption. Quirk and Tinsley (1973) included the possibility of $Z = 0$ gas being added continuously to the evolving region. Such infall, combined with a variable IMF, can simultaneously give the present metal abundance and the observed $N(Z)$ distribution, without having the last bit of gas be about to disappear into low mass stars. Models of this type have also been considered by Larson (1972) and Fowler (1972). The infalling material keeps the metal abundance in the gas from building up past a steady-state level which is comparable with the yield, y, as well as providing a continuous source of gas for new star formation. Deuterium abundance is not a problem in such models, since the infalling material carries the primordial amount. The chief difficulty with such models is that the required infall may not occur (Verschuur, 1973). In addition, no homogeneous model can account for the observed variation in metal abundance of stars with the same age (Fig. 4) without some additional assumptions.

Inhomogeneous evolution was discussed qualitatively by Searle (1973) and quantitatively by Talbot and Arnett (1973; Talbot, 1974). Each suggested that stars would be formed preferentially in regions of higher-than-average metal abundance. Searle had in mind a large scale effect occurring during the collapse of the Galaxy from spheroidal to disk form, while Talbot and Arnett contemplated more local variations within the galactic disk. By making star formation sufficiently fussy about the Z where it occurs, one can reduce the number of low mass, low-metal stars to an arbitrarily low level in this picture, without producing an excessive metal abundance at the present time or using up all the gas in the near future. Satisfactory models can be obtained with either constant or variable Initial Mass Function. The locally inhomogeneous models automatically produce a variation in Z for stars of the same age. The existence of a spread in Z values in stars of given age (Hearnshaw, 1972; Dixon, 1966; Eggen, 1964) provides *a priori* evidence of the existence of the required inhomogeneities in the instellar medium, but Edmunds (1974) maintains that known mixing processes are sufficiently efficient that the inhomogeneities required to make Metal Enhanced Star Formation an important effect cannot exist. He attributes the observed spread in Z to a radial abundance gradient or to systematic expulsion of metallic grains from the atmospheres of stars during formation (Edmunds and Wickramashinghe, 1974).

All of these modifications (variable IMF, infall, inhomogeneities) are fundamentally ways of hiding, or avoiding the formation of, large numbers of low mass, low-metal abundance stars in the solar neighborhood. All are at least partially successful, but all are unrealistic in that they do not include effects of the initial collapse of a galaxy or the later systematic flows of gas within it. For the short time scales encountered in collapse, inhomogeneity is a necessary state of affairs. Neither mixing nor infall can occur simultaneously over the whole galaxy. The Ultimate Model of galactic evolution will have to be both a dynamical and an inhomogeneous one.

E. Dynamical models of galactic evolution

The Ultimate Model of galactic evolution has not yet been calculated, but most of the recent models do include some of the effects of the interaction between collpase of the Galaxy and chemical evolution that were implicit in the work of Eggen *et al.* (1962) and Searle (1973). Three different types of collapse have been considered: (a) that of a spheroidal cloud to form a spheroidal galaxy (Larson, 1974; Larson and Tinsley, 1974), (b) that of a flat disk, representing a spiral or irregular galaxy (Talbot and Arnett, 1974), and (c) from an initially spheroidal halo toward a flat disk (Ostriker and Thuan, 1975).

1. Spherical galaxies

The evolution of a collapsing spherical gas cloud has been followed by Larson (1974). The cloud initially has a constant density, a mass of $10^{11}\mathfrak{M}_\odot$, a radius of 30–50 kiloparsecs (usually, but values from 10 to 20 000 kiloparsecs were tried), which either remains fixed or expands gradually with the Hubble expansion, and a uniform average velocity dispersion of 6 to 55 km/sec. The evolution of the system is described in terms of fluid-dynamical variables for two fluids, one of gas and one of stars. The rates of dissipation of kinetic energy of internal motions in the gas and of star formation are taken to be proportional to the dynamical time scale, that is,

$$(1/\alpha_g)(d\alpha_g/dt) = -C_D t_f^{-1},$$

and

$$(1/\rho_g)(d\rho_g/dt) = -C_s t_f^{-1},$$

where

$$t_f = (3\pi/32G\rho)^{\frac{1}{2}}$$

and ρ is the mean density of matter inside the radius r at which t_f is evaluated. The acceptable models all have values of the proportionality constants C_D and C_S close to unity. The near identity of the three time scales is very reasonable if dissipation and star formation are both primarily due to shock waves associated with the collapse. A few models in which star formation and gaseous dissipation rates depended only on the local gas density were also calculated. Initial mass functions of both the power law and the Gaussian type (discussed in Sec. IV.A) have been tried. In order to calculate stellar mass loss and heavy-element production rate, the stars are divided into a number of mass intervals. The lower mass limit is varied to allow for the uncertainty in the solar neighborhood density of M dwarfs. The input of energy and momentum, as well as mass, into the gas due to mass ejection from stars is included in the calculations.

Output from the models (Larson and Tinsley, 1974) which can be compared to observations includes the distribution with radius of (a) gas and star densities and luminosities, (b) star formation rate, (c) star and gas metal abundances, (d) star and gas velocity dispersions, (e) *UBV* colors, and (f) mass to luminosity ratio as well as integrated luminosity, color, continuum spectral energy distribution and supernova rate, all as a function of time. The main effect of the gradual dissipation of gas kinetic energy is a continuing flow of the gas toward the center of the model. Thus a central density peak is gradually built up, and since the gas is being gradually enriched in metals as it flows in, an abundance gradient also develops. In the models with a fixed boundary, most of the star formation is over in a few billion years, as is most of the metal enrichment. The inner regions of the models have very few metal-poor stars, and remain somewhat bluer than the rest of the galaxy, due to ongoing star formation. Some, but not all, elliptical galaxies do, in fact, have blue nuclei. After the completion of most star formation, the models get gradually fainter with time, implying the need for significant evolutionary corrections to the absolute magnitudes of elliptical galaxies if they are to be used in a cosmological test. In the models with expanding boundaries, star formation and metal enrichment continue to ages of 10–12 billion years and the systems remain bluer.

Within the range of models calculated, acceptable matches can be found for most of the observed dynamical and photometric properties of most galaxy types. Properties which can be matched include the supernova rate as a function of Hubble type (if all supernovae are due to massive stars), the distribution of surface brightness (e.g., Miller and Prendergast, 1962, for NGC 3379), the gradient of stellar metal abundance (e.g., Spinrad *et al.*, 1972, for NGC 3379), $N(Z)$ in the solar neighborhood (if our galactic disk is assumed to have formed in the same way as the nuclei of elliptical galaxies; see also Tinsley, 1975a, on this point), the velocity dispersion in elliptical nuclei (e.g., Burbidge *et al.*, 1961, on NGC 3379), some of the energetic properties of quasars, if they are assumed to be the developing nuclei of giant elliptical galaxies, *UBV* colors as a function of Hubble type, mass to luminosity ratios, and continuum narrow band spectral energy distributions. Some preliminary work on rotating models suggests that the metal abundance gradient for them will be largely along the rotation axis, which is also in good accord with observation in both our Galaxy and NGC 3115 (Spinrad *et al.*, 1972).

The major set of observed effects not accounted for by the models are the variations of total metal abundance and degree of central concentration with total galactic mass. In addition, the models for ellipticals all had the blue nuclei which are observed in only some galaxies. Both of these problems suggest that loss of gas from the systems may be important. If gas is lost early from low mass systems, it is not available to flow toward the center. Thus, central density concentrations and abundance gradients cannot be built up. Star formation must also stop, and the total metal enrichment will always remain small. Eventual loss of the residual gas from larger elliptical galaxies will prevent their nuclei from being blue. Larson (1975) has considered the effects of supernovae on the early evolution of galaxies and finds that they probably produce enough heating to expel gas in a hot galactic wind, at least in low mass galaxies, as originally suggested by Mathews and Baker (1971). Gas loss due to this mechanism becomes important about two billion years after the beginning of the collapse of the galaxy. For an initial mass of $3 \times 10^9 \mathfrak{M}_\odot$, half the mass is lost in the form of gas due to supernovae without ever having been turned into stars. The fraction of mass lost drops to 10% at an initial mass of $10^{11} \mathfrak{M}_\odot$ and is greater than 50% for initial masses less than $10^9 \mathfrak{M}_\odot$. The predicted density and metal abundance gradients as a function of total galactic mass are in good accord with observations, significant gradients being formed only in galaxies above about $3 \times 10^9 \mathfrak{M}_\odot$, which is roughly the crossover found by van den Bergh (1972a) between dwarf ellipticals with no noticeable nucleus and normal or giant ellipticals with bright nuclei. The total metal abundance achieved after 10–12 billion years rises with galactic mass up to $10^{11} \mathfrak{M}_\odot$, at which point all gas is always retained. The increase is about two orders of magnitude in Z from the smallest models studied ($10^5 \mathfrak{M}_\odot$) to the largest ($10^{13} \mathfrak{M}_\odot$). This agrees with the observations for low mass galaxies, but the total observed metal abundance appears to continue to increase with mass beyond $10^{11} \mathfrak{M}_\odot$ (Faber, 1973; Sandage, 1972a). Loss of gas due to supernovae does not appear to account for the redness of the nuclei of most giant ellipticals or for their over-all low observed gas (HI) abundance.

2. Disk galaxies

Talbot and Arnett (1975) have considered the evolution of a flat, round disk of gas with an exponential distribution of surface density $[\rho(r) = \rho_0 \exp(-r/r_0)]$ assumed to have been left behind after the rapid collapse of an initially spherical halo. The fraction of the matter that ends up in such a disk will depend on the fraction of the initial cloud which had a high initial angular momentum (Sandage *et al.*, 1970). Since the models neglect the remaining halo component, they will apply most nearly to late type spiral (Sb, Sc) and irregular galaxies. The disks constitute a two parameter family in central surface density, σ_0, and scale size, r_0.

The models assume that the disk remains symmetric about the rotation axis and that there is no radial gas flow. The disk is taken to be in hydrostatic equilibrium in the vertical direction so that at each radius $p_i = \rho_i v_i^2$, where p_i is the pressure due to each component (gas, stars), ρ_i is its volume density, and v_i is its turbulent velocity plus some correction factor to allow for the effects of cosmic rays, magnetic fields, and so forth. The turbulent velocity is determined by the balance between dissipation and the turbulent input due to OB stars and supernovae. As a result, any star formation rate which depends on gas density leads to an equilibrium situation, since an increase in star formation increases the turbulent velocity, thereby reducing the gas density and the star formation rate. As a result either Metal Enhanced Star Formation or homo-

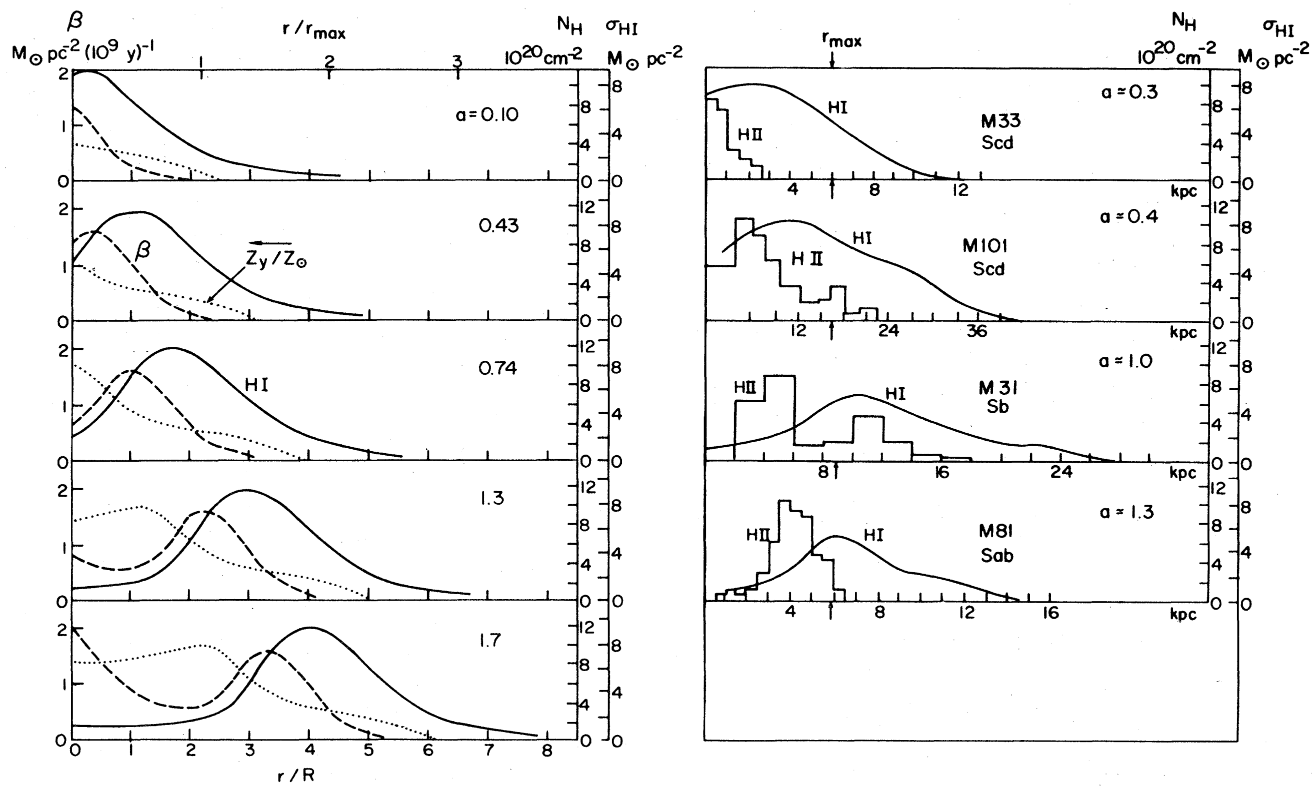

FIG. 17. A comparison of the observed distribution of HI and HII in galaxies as a function of Hubble type, with the calculated distribution of tota residual gas and Strömgren spheres as a function of central density of the galactic disk in the models of Talbot and Arnett (1975). The five models are arranged in order of increasing central density. The solid line is the radial distribution of surface density of HI in the models; the dashed line is the rate of formation of new, massive stars (and, thus, roughly the surface density of HII); and the dotted line is the heavy element abundance in newly formed stars, also as a function of position in the model. The four observed galaxies are arranged along the Hubble sequence. The smooth curve is the distribution of 21 cm emission, and the histogram curve the distribution of HII regions, in arbitrary units. The data are derived from sources listed by Talbot and Arnett (1975). The agreement between observed and modeled HI and HII distributions is remarkably good, considering that only one parameter in the model, the central density, was varied.

geneous evolution in regions at a given distance from the center leads to much the same model. Given a particular initial mass function (which is taken to be constant in these models) and a knowledge of stellar evolution (luminosity, color, and amount and composition of matter ejected as a function of mass, composition, and age), each radius zone in the model can be evolved forward in time using a prescribed rule for star formation rate and gaseous dissipation. Properties of the models which can be calculated as a function of time include total gas mass, remaining gas as a function of radius, average stellar metal abundance and its radial gradient, average M/L ratio and its radial gradient (high at the center where only low mass stars remain and in the outer regions where the matter is still gas; low in the region of active star formation), integrated colors and luminosities and their radial gradients, and the supernova rate and its correlation with integrated color.

Models have been calculated for central densities from 30 to 2000 $\mathfrak{M}_\odot$ parsecs^{-2} and a variety of size scales. The dependences of some of the observed properties of a 13 billion year old galaxy on σ_0 and r_0 are given by:

$$B - V = 0.22 \log\sigma_0 + \text{const},$$
$$\log[M(\text{HI})/M_{\text{tot}}) = -0.62\sigma_0 + \text{const},$$
$$\log(M_{\text{tot}}) = \log\sigma_0 + 2\log r_0 + \text{const},$$
$$\log L(\text{blue}) = +1.2\log\sigma_0 + 2\log r_0 + \text{const}.$$

Models can be found covering the range of observed integral properties of galaxies from Sa to Ir. Clearly a low value of σ_0 leads to a blue color, large residual gas mass, a small total mass, and small total luminosity, and conversely for a large value of σ_0. It is not possible to correlate the necessary σ_0's quantitatively with observed values because the central density of a real galaxy has a large contribution from its spheroidal (halo) component. Observed supernova rates and the thickening of the gas disk at large radii are reproduced. Regions of high initial density pass through a maximum in metal abundance at an age of a few billion years, which could account for super metal rich stars. Observed radial gradients of abundance and gas density in galaxies are also well matched. Figure 17 shows observed and modeled radial gradients in the amount of total gas (i.e., HI) and ionized gas (HII; assumed proportional to the OB star density) as a function of Hubble type for the observed galaxies and σ_0 for the modeled ones. The agreement is striking. The number of stars as a function of metal abundance in the solar neighborhood is also reproduced in the models with MESF.

3. Models with massive halos

Ostriker and Thuan (1975) have explicitly included the dynamical phase in which matter shed by dying halo stars accumulates in a smaller, more rapidly rotating disk in their calculations. The resulting disk can be entirely secondary or have contained 10%–100% of its final mass at the time nuclear processing started in the halo. The model with 100% of the disk there initially, of course, corresponds to one of the simple models that we saw in Sec. III.B do not work. In addition to the initial ratio of halo mass to disk mass, the model assumes a birthrate function in the halo, normally a power law in mass, like the Salpeter birthrate function, and a decaying exponential in time, with a time constant of 10^{8-9} yr, appropriate to the dynamical collapse of the halo. The halo birthrate function is normalized so that a total $75\,\mathfrak{M}_\odot$ persec^{-2} of matter is found in the disk after 12 billion years. The IMF in the disk is also independent of time and its form is adjusted to give the presently observed disk luminosity function. The disk IMF is, therefore, not necessarily a power law, but a Salpeter IMF produces an acceptable model, provided that some massive stars are presently seen only as infrared sources due to shielding by gas and dust. The disk birthrate function is normalized by the requirement that the gas density be $10\,\mathfrak{M}_\odot$ parsec^{-2} at an age of 12 billion years. Gas is lost by stars in the halo through both supernovae and planetary nebulae. Half the material ejected by supernovae is assumed to be blown out of the galaxy, and half to fall down into the disk. All the matter lost in planetary nebulae is retained in the disk. Gas is transferred instantly from halo to disk so that no reprocessing occurs in the halo.

The model predicts the usual chemical, light, and mass distribution properties as a function of time. Most of these (M/L in the disk, the shape of the $Z(t)$ curve, and so forth) can be brought into good accord with the observations for models with the disk ranging from being 100% primary to 100% secondary. Models in which at least 25% of the disk is made of gas shed from halo stars also reproduce well the observed $N(Z)$ curve for the solar neighborhood. This happens in a very natural way, since the low mass, low-metal abundance stars remain behind in the halo where we do not observe them. The model with a completely secondary disk presents several problems. First, the initial disk metal abundance is higher ($Z \sim 0.08$) than that of even super metal rich stars. Second, the average age of the heavy elements present in the disk gas 4.6×10^9 yr ago and available for incorporation into the Solar System is larger (though not by very much) than the results of nucleocosmochronology allow it to be. Finally, since all the disk gas has been through stars, no primeval deuterium can remain in it. This is a problem only if no place other than the Big Bang can be found in which to produce ^{2}H (see Secs. III.A and III.B). These three difficulties are eliminated if the disk consists of only $\frac{1}{4}$ to $\frac{3}{4}$ secondary gas and the rest was there before significant evolution of halo stars took place. In this case, up to 30% of the gas presently in the disk has never been in stars. The Big Bang production of deuterium need, therefore, only have been about three times the presently observed value ($D/H = 1.4 \times 10^{-5}$). The universe is thereby constrained to have only about 10% of the total density of mass–energy necessary to close it. This does not appear to contradict any other observations (Gott *et al.*, 1974).

Thus a range of models with partly-secondary disks exists which pass all of the tests to which they have so far been subjected. These models have several interesting properties. The final ratio of disk mass to halo mass is about 1/20. The galaxies are, therefore, protected from the kinds of instabilities in flat disks studied by Ostriker and Peebles (1973). The total mass of a typical spiral galaxy and its mass/luminosity ratio are about 20 times those of the disk component alone, in good accord with the masses and M/L's deduced by Einasto *et al.* (1974) and Ostriker *et al.* (1974). The halo luminosity of these models is evolving rather rapidly with time as did the luminosity of Tinsley's (1972) models for spheroidal galaxies. Evolutionary corrections to M_V of elliptical galaxies observed at red shifts greater than 0.2 are, therefore, predicted to be an important factor in determining the deceleration parameter, q_0. The effect is to convert an apparent observed q_0 near one to a "real" q_0 much less than one. This, like the need for deuterium production, implies an open universe (Gott *et al.*, 1974).

V. CONCLUSIONS AND EXHORTATIONS

Many, many pages ago we set out to try to discover the distribution of the chemical elements and their isotopes throughout the universe, and the nature of the nuclear processes which had produced them, and to incorporate the processes and the objects in which they occur into a model of the history of galaxies that would account for how the elements came to be where they are. How successful this effort has been can be debated. On the one hand, Tinsley (in NATO, 1974) has pointed out that we can outline a consistent scenario of the history of the universe which seems to account for nearly everything. The scenario has five stages. First, a Big Bang universe (with an expansion rate H_0 consistent with observations, the constraints of nucleocosmochronology, and the simplest possible cosmological model, and a deceleration parameter, q_0, given by the amount of matter we see in galaxies if they have massive halos) synthesizes ^{1}H, ^{2}H, ^{3}He, ^{4}He, and ^{7}Li in about their observed proportions during an early, hot, dense phase. Second, perturbations in the early universe give rise to protogalaxies. This is very hard to do in a low-density universe, but that is someone else's problem. Third, galaxies are formed as collapsing gas clouds in which the dominant effects are: (a) The relative amounts of high and low angular momentum gas determines where the galaxy will fall on the Hubble sequence. (b) A portion of the gas gets locked into low mass stars in an extended halo, the amount remaining free determining the luminosity class of the galaxy, and (c) The dynamics of the collapse gives the distribution in the visible part of the galaxy of metals, color, stellar populations and the like. This stage left our Galaxy with a first generation of low-metal abundance stars in a halo, and, after 2×10^9 yr, a residual disk whose metal abundance was already about half that of the Solar System. Fourth, the galaxies continue to evolve on a longer time scale with the following dominant processes: (a) continued star formation, evidently with a nearly constant Initial Mass Function, at an exponentially decaying rate (the time constant depending on the local gas density) in the disk, which may or may not develop a spiral structure, depending, perhaps, on its angular momentum and the rate of energy input into the interstellar gas by supernova, (b) explosive or quasistatic stellar mass loss in the disk, which gives a slow increase in Z and a change in the helium abundance, Y, comparable with the change in Z, (c) production of ^6Li, ^9Be, and $^{10,11}B$ by cosmic ray spallation in the interstellar medium, and (d) continuous dilution by low-metal abundance gas which is shed by the evolving halo stars near $1\,\mathfrak{M}_\odot$ and keeps the disk composition rather constant. Fifth, a portion of the heavy elements trapped in the solar nebula gives rise to the Earth, cosmobiological evolution, people, and astronomers.

Viewed in this way, the main problems all seem to have been solved, and we may wonder what we are all going to do for the rest of our careers. On the other hand, Sargent (in NATO, 1974) reminds us that we have almost no empirical information on the formation of stars and none on galaxies; we know nothing observationally about the evolution of galaxies, and we seemingly do not know the mass of our own Galaxy to within a factor of 10; nor do we have much direct observational evidence that the iron-peak elements are made in stars or anywhere else. Thus, several of the vital stages in the evolutionary scenario are accomplished only by saying "let there be galaxies," "let there be stars," and so forth. In addition, Reeves (in NATO, 1974) has cautioned that we must have great reluctance against accepting new data uncritically, forgetting embarassing old data, accepting new theories uncritically, forgetting alternative unfashionable theories, and being taken in by authoritarianism, fashion, and bandwagon effects. This is obviously good advice in any field of scientific endeavor, but particularly so here where the theoretical superstructure is massive compared to the foundation of observations and experiments on which it is based.

We referred repeatedly in Sec. III to the lack of necessary laboratory data on nuclear masses, lifetimes, and cross sections. Every additional one of these that can be measured is valuable, not only because it can improve a particular nucleosynthesis calculation, but also because it provides another point on which to fit and test semi-empirical theories for nuclear properties that cannot be measured. Fowler *et al.* (1975) draw attention to specific uncertain, but important, cross sections. New observations are equally vital. We know very little about isotope ratios outside the Solar System. Thus, among other problems, there is no direct evidence for the existence of p-process nuclides anywhere but in the Earth, Moon, and meteorites. Information on variations between stars and galaxies in relative abundances of the heavy elements is also woefully inadequate. The variations which have been found seem to be both smaller and less repeatable from place to place than would be expected from most of the models of nucleosynthesis and galactic evolution. On the theoretical side, the most obvious deficiency is the absence of any understanding of star formation, but we also need to know much more about the late phases of stellar evolution in order to be able to predict how much of each element should be expelled by a typical supernova, or even what sort of star gives rise to a typical supernova.

Sargent's final benediction to the NATO Institute was "Go home and observe," and, although the effect was somewhat spoiled by a subsequent complicated announcement about the supper arrangements, most of us left with a renewed zeal for observing, experimenting, calculating, or whatever we think we do best. There does not seem to be any immediate danger of running out of unsolved problems.

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