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Nuclear Young Star Cluster in our Galactic Center

by

Siyao Jia

A dissertation submitted in partial satisfaction of the

requirements for the degree of

Doctor of Philosophy

in

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in the

Graduate Division

of the

University of California, Berkeley

Committee in charge:

Professor Jessica Lu, Chair

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Spring 2020

Nuclear Young Star Cluster in our Galactic Center

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Siyao Jia

Abstract

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Doctor of Philosophy in Astrophysics

University of California, Berkeley

Professor Jessica Lu, Chair

The formation of the young stars near the Supermassive Black Hole in our galactic center has been a mystery for a long time. In my thesis, I will present how we use the improved astrometric measurements, combined with well-measured radial velocities to study their dynamical structures and constrain their formation scenario.

First, we present an improved relative astrometry for stars within the central half parsec of our Galactic Center based on data obtained with the 10 m W. M. Keck Observatory from 1995 to 2017. The new methods used to improve the astrometric precision and accuracy include correcting for local astrometric distortions, applying a magnitude dependent additive error, and more carefully removing instances of stellar confusion. Additionally, we adopt jackknife methods to calculate velocity and acceleration uncertainties. The resulting median proper motion uncertainty is 0.05 mas yr^{-1} for our complete sample of 1184 stars in the central $10''$ (0.4 pc). We have detected 24 accelerating sources, 2.6 times more than the number of previously published accelerating sources, which extend out to $4''$ (0.16 pc) from the black hole. Based on S0-2's orbit, our new astrometric analysis has reduced the systematic error of the supermassive black hole (SMBH) by a factor of 2. The linear drift in our astrometric reference frame is also reduced in the North-South direction by a factor of 4. We also find the first potential astrometric binary candidate S0-27 in the Galactic center. These astrometric improvements provide a foundation for future studies of the origin and dynamics of the young stars around the SMBH, the structure and dynamics of the old nuclear star cluster, the SMBH's properties derived from orbits, and tests of General Relativity (GR) in a strong gravitational field.

Second, we measure the 3D kinematic structures of the young stars within the central 0.5 parsec of our Galactic center using the 10m Keck telescope with a time baseline from 1995 to 2015. Using these high-precision astrometric measurements of positions and proper motions, plus radial velocities from literature, we are able to constrain the orbital parameters for each young star. Our results clearly show the significant clockwise stellar disk as has been

proposed before but with a better disk membership measurement. Besides the previously identified clockwise stellar disk, we have found another almost edge-on stellar disk, but we cannot rule out the possibility that it is a stream associated with the IRS 13 group. Based on their disk membership, the young stars are divided into different dynamical subgroups, and the disk properties, e.g. eccentricity and semi-major axis, are compared between these three dynamical subgroups. In order to understand the asymmetric 2D density profile, we simulated a single-age, single-metallicity star cluster with observational disk properties to reproduce the spatial distribution for each dynamical subgroup, and then applied a differential extinction map. It is the first time that stellar populations are simulated and compared to observational density profile, which provides us with a deeper understanding of their dynamical structures and helps us distinguish different star formation scenarios around the supermassive black hole at our Galactic center.

To Luning Zhao, for all his love and support

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Chapter 1

Introduction

1.1 Supermassive Black Hole and Nuclear Cluster in our Galactic Center

Nuclear Star Clusters (NSCs) and Supermassive Black Holes (SMBHs) are found to coexist within the central parsec of many different types of galaxies [1]. Furthermore, there are clear indications that NSCs, SMBHs, and their host galaxies evolve together. For example, [2] found that $M - \sigma$ relationship, the empirical correlation between the mass of the SMBH M and the stellar velocity dispersion σ of a galaxy bulge, applies both for SMBHs and NSCs with similar slopes, although at a certain σ , NSCs are 10 times more massive than SMBHs. Additionally, Kormendy et al. (2013) [3] showed that the combined mass of the SMBH and NSC scales with a galaxies bulge mass with much less scatter than either the SMBH or the NSC separately, which suggests a strong dependency between their formation and growth. The formation mechanism for NSCs is still widely debated, as is the formation and growth of SMBHs, because the strong tidal force at the galactic center will disrupt normal star formation.

Located at a distance of 7.97 kpc [4], our Galactic Center (GC) has been observed extensively due to its close proximity. It hosts a compact radio source, Sgr A*, which is associated with a $3.98 \times 10^6 M_\odot$ SMBH [5, 6, 7, 8, 9, 10, 4]. A NSC surrounds Sgr A* and is the only NSC in which individual stars can be resolved with the largest telescopes such as the W.M. Keck Observatory and the Very Large Telescope (VLT) [8, 9, 10, 11, 12]. The total mass of the NSC is $\sim 10^7 M_\odot$ with a half-light radius $\sim 2\text{--}5$ pc [13, 14, 15]. The NSC at GC is composed mostly of old, late-type stars (age $\gtrsim 8$ Gyr). However, there is also a population of hot, young stars (4-8 Myr) that dominates the luminosity in the central parsec with a total mass of $\sim 10^4 M_\odot$ [16]. So far, spectroscopic observations have identified more than 150 early-type stars, including Wolf-Rayet stars, OB supergiants and OB main-sequence stars [17, 16, 18, 19, 20]. Because the age of this population (4 - 8 Myr) is much less than the relaxation timescale in the Galactic center (~ 1 Gyr, [21]), the origin of these stars can be constrained through studies of their dynamic structures. Therefore, our own GC provides

a unique test bed for understanding star formation around SMBHs. Only in the GC can we resolve individual stars and measure their motion, photometry, and spectroscopy with sufficient precision to constrain their dynamics and IMF and therefore constrain theories of star formation around SMBHs.

1.2 The Need for Accurate and Precise Astrometry

There are still many open issues that observations of the GC, particularly astrometric measurements, can address: (1) the spectroscopically identified (and thus relatively bright) old stellar population does not show evidence for a cusp, as predicted from theory [22, 23, 24], (2) the formation mechanism of the young populations is still not well understood [25, 26, 18, 27], (3) there are many unusual stars that may be the product of tidal interactions with the SMBH or with other stars in the region [28, 29, 30], and (4) short period stars, such as S0-2, provide an opportunity to test General Relativity (GR) in a strong field around a SMBH for the first time [31, 32]. We discuss each of these in more detail below.

Late-type populations are presumably dynamically relaxed and are predicted to have a steep core profile, or a cusp [33, 34]. Contrary to theoretical predictions, some observations have found that the surface density profile of late-type stars in our GC is flat, consistent with the NSC having no cusp. Several plausible dynamical scenarios have been suggested to explain the depletion of late-type giants, such as mass segregation, stellar collisions, or a recent merger event [22, 23, 24]. However other studies pointed out that this missing stellar cusp problem is limited to the brightest few percent stars due to observational difficulties and the observed density of the faintest stars is actually consistent with the existence of a stellar cusp [35, 36, 37]. Obtaining an unbiased measurement of the stellar distribution will be the key to solve the missing cusp problem. In particular, significant acceleration detections from astrometric measurements will greatly help with the line-of-sight distance. The line-of-sight distance can not be obtained in any other way for stars located 8 kpc away, but the absolute value of line-of-sight distance can be calculated if a significant acceleration in proper motion is detected.

The *in-situ* star formation mechanism is now widely accepted, where stars in the vicinity of SMBH are formed from a dense gas disk [25, 26, 16, 13, 38] (see more details about *in-situ* formation scenario in Section 1.3). Detailed measurements of dynamical properties of young stars will help in determining their origin. Previous publications have found that about 20% of the young stars rotate in a clockwise disk with an inner edge at 0.8 arcsec, which is an outcome of the *in-situ* formation mechanism [25, 16, 18, 24, 27, 39] (see more details about their stellar properties in Section 1.4). This 20% limit might be a lower limit if we consider binaries [40]. The dynamical structure of those young stars will greatly help us understand the star formation under extreme environments and even get a direct understanding of the SMBH's properties which cannot be obtained by other methods.

G2 is a dusty red object that was discovered in 2011 [28]. The fact that it is a very red point at L band, but very faint at K band, and shows recombination lines from Br- γ , makes

G2 look like a pure gas [41, 28, 42]. However G2 surprisingly survived its closest approach to Sgr A* in early 2014 [30, 43, 29, 44, 45], where a pure cloud would get tidally disrupted during periaipse passage. This implies that G2 has to be a stellar object, probably a binary merger product. Better characterization of the binary fraction is needed to answer whether this mechanism is able to explain the star formation in the GC. In Section 2.5, we find the first potential binary candidate using pure astrometry. As astrometric monitoring of the GC continues, more candidate astrometric binaries are likely to be detected, and we will be able to better constrain the binary fraction at the GC.

S0-2 is one of the most interesting young sources, which is very close to Sgr A* (< 0.5 arcsec) and has a short period of only 16 years. In fact, S0-2 has provided some of the most direct evidence for the presence of a SMBH in our GC and its orbit has been used to calculate the mass and distance to Sgr A* [8, 9, 10, 11, 12]. In 2018, S0-2 reached its second observable closest approach to the SMBH. While GR has been thoroughly tested in weak gravitational fields for considerable amount of times [46, 47], this event marked the first direct test of GR in a strong gravitational field around a SMBH by measuring the relativistic redshift in S0-2's radial velocity [48, 49]. Improved astrometry is necessary to characterize the full 3D orbit for S0-2 and will be even more important for future periaipse precession tests.

The key to solve those puzzles is to obtain accurate and precise astrometric measurements, specifically the positions, proper motions, accelerations and orbits. Previous observations have been conducted using VLT and Keck telescope: astrometric positions from imaging data reaches an accuracy of ~ 0.3 mas, a factor of 200 smaller than the image resolution in K band [50].

Achieving precise and accurate astrometry requires cross-matching and transforming stellar positions and photometry into a common coordinate system. However, this is particularly difficult in the GC region for several reasons [50]. (1) The GC field is very crowded. Even with AO, confusion occurs frequently as two or more stars move past each other. (2) Stars have high proper motions, especially for those close to the central SMBH, and therefore require careful consideration when aligning (cross-matching and transforming) epochs [39, 11]. (3) Long time baseline observations are critical for accurate astrometry, but the experimental setup can change with time and the associated distortion needs to be handled carefully [51, 52]. In my thesis, we implement effective methods to deal with the existing difficulties in accurate relative astrometry measurement, including confusion, imperfect point spread function (PSF), local distortion of non-standard AO epochs.

1.3 *In-Situ* Star Formation Scenario

Observations of young stars at the GC currently favor *in situ* formation models, meaning that the young stars formed roughly at where we see them today, within 0.5 pc of the SMBH [16, 27, 53]. Alternative scenarios invoking the inspiral of a massive star cluster are unlikely since the cluster would dissolve and deposit stars with a much shallower surface density profile of $\Sigma \propto R^{-0.75}$ (where Σ is stars arcsec $^{-2}$ and R is the 2D radius on the plane of the

sky) than the observed r^{-2} relation [54, 16]. Moreover, this cluster must be extremely dense to overcome the tidal force. An initial mass of $> 10^5 M_\odot$ is required, which will contain a larger number of massive stars than currently observed at the GC [55, 56].

In situ formation is theoretically possible in an accretion disk around the SMBH, if it is massive enough to collapse vertically under its own self-gravity [57, 58, 59, 60, 25, 26]. When a disk reaches a surface density that is just large enough to initiate star formation, the first low mass protostars form. Feedbacks from those stars will then heat the disk up to a point where it becomes stable to self-gravitation, shutting off further disk fragmentation. In the mean time, those stars remain embedded in the disc and continue gaining mass at very high rates (super-Eddington). As a result, an average star created in such a disc may become very massive, depending on the feedback efficiency. This process happens from a few percent of a parsec to a few parsec with a peak effect at $R \sim 0.1 \text{ pc}$ [58, 61, 26, 62]. This scenario can explain the existence of disk stars and a top-heavy initial mass function (IMF) in our GC [63], although it might not be the only explanation since many formation scenarios still don't uniquely predict the IMF.

However, if the disk is formed through steady accretion of gas, stars with circular orbits are more likely to form, which may contradict the observed eccentricity distribution that peaks at $e=0.27$. Furthermore, only 20% of the young stars are on the disk and it is not clear whether 80% of the stars could be scattered from the disk in only 4 - 8 Myr. Modified *in situ* formation scenarios have been proposed, including: (a) the initial gas is not uniformly distributed, (b) star forms in repeated episodes, (c) after the gas disk collapse, stars dynamically evolved off the disk. These formation scenarios can be disentangled by comparing the dynamical structure and stellar mass function among different dynamical sub-groups. For example, scattering processes would likely lead to more low-mass stars out of the disk. If the disk and off-disk stars have similar MF, they are more likely to have formed from gas in complex structures, not a simple disk. If stars form in repeated episodes, they could have different post-main sequence turn-off points if the gap between the formation episodes is large enough.

1.4 Stellar Properties for Young Stars in the GC

Previous study shows those Galactic central young stars can be divided into three dynamical groups: (1) $\sim 12\%$ of the young stars (within 0.03 pc) are in the most inner part with high eccentricities $\bar{e} = 0.8$ and randomly orientated orbits. (2) $\sim 20\%$ of the young stars are on a well-defined clockwise (CW) rotating disk (0.03 pc-0.5 pc) with moderate eccentricities. (3) $\sim 68\%$ of them are off-disk stars over the same distance, which are more randomly distributed and also with moderate eccentricities [63, 39]. At smaller radii, dynamical effects will randomize the stellar orbits within 4-6 Myr, which is the case for the first group. The existence of this CW disk has already been verified to be significant ($\sim 20\sigma$) [39], but the kinematic structure analysis, particularly the precision of disk membership at larger radii, strongly depends on the astrometric precision.

The IMF of the entire nuclear cluster has been measured to be top-heavy with $dN/dm \propto m^{-\alpha}$ and $\alpha = 1.7$ (vs. $\alpha = 2.35$ in the Solar neighborhood) [63]. However, no distinction was made between young stars on and off the disk. We are going to derive the IMF for each of the dynamical sub-groups to look into their formation history. After getting the disk membership from dynamic structure analysis, we use Bayesian inference method developed in [63] to derive the stellar properties by comparing the observed cluster properties with the simulated synthetic cluster. In this process, we will consider binaries and multiple star system that are essential to derive the ages and IMF slope. A Bayesian inference method will give us the best fit properties of our cluster, with which we could use to determine the star formation history. For example, we can see whether the slope of the IMF α is the same for different sub-groups and whether it is flatter than the Salpeter slope of 2.35 in the Milky Way disk as showed by previous work [64].

1.5 Outline

Chapter 2

This chapter is about the improved Relative Astrometry for Velocities, Accelerations, and Orbits near the Supermassive Black Hole. The observational data set, including both Speckle and AO, and the data reduction process are presented in Section 2.1. In Section 2.2, the multi-epoch alignment procedure is explained in detail. With a clean astrometric catalog in hand, we derive proper motions, accelerations and orbital fits in Section 2.3. The improved astrometry and newly detected accelerating sources are presented in Section 2.4. We also briefly discuss the potential scientific cases that could benefit from our accurate astrometric measurements in Section 2.5. Finally, we summarize our work in Section 2.6.

Chapter 3

In this chapter, we simulate and compare stellar populations among different dynamic sub-groups within the central 0.5 pc of our Galaxy. Our observation set up and data reduction are presented in Section 3.1. Orbital parameters are derived in Section 3.2 and the disk membership analysis is shown in Section 3.3. Results of stellar disk properties and cluster simulation are presented in Section 3.4. We discuss our findings in Section 3.5 and summarize in Section 3.6.

Chapter 2

Improved Relative Astrometry for Velocities, Accelerations, and Orbits near the Supermassive Black Hole

2.1 Observations and Data Reduction

Observations

The central 10 arcsec region of the GC (approximately centered around Sgr A*) has been monitored from the W. M. Keck Observatory with diffraction-limited, near-infrared imaging cameras since 1995. Images used in our multi-epoch astrometric analysis have been obtained in two different modes: Speckle imaging from 1995 to 2005 (26 epochs) and laser guide star adaptive optics (LGS-AO) imaging from 2005 to 2017 (30 epochs).

All Speckle data sets were obtained in the K-band ($\lambda_0 = 2.2\mu\text{m}$) using the Near Infrared Camera (NIRC; [69, 70]) on the Keck I telescope with a field of view (FOV) $\sim 5'' \times 5''$ and a pixel scale of 20 mas. Speckle data is taken using very short exposure times (0.1s) to freeze the atmospheric distortion. Details of the Speckle observation in Table 2.1 can be found in [6, 7, 67], [65], [66] and [71].

Since 2005, we have used the Keck II LGS-AO system [72, 73] with the near infrared camera NIRC2 (PI: K. Matthews) in its narrow-field mode, which has a FOV of $\sim 10'' \times 10''$ and a plate scale of 9.952 mas/pixel [51]. After 2014, the adaptive optics system and NIRC2 camera were realigned, and the plate scale changed to 9.971 mas/pixel [52]. In this paper, we include new data from 2014 - 2017, increasing the time baseline of the LGS-AO data set by $\sim 30\%$ (2005 - 2017, compared to 2005-2013 in [12]¹). The new LGS-AO data sets (Table 2.2) are obtained in an identical manner and have comparable quality to our previous observations. Additional details about our observational setup are presented in [67,

¹The “2004 July” LGS-AO epoch in [39] is dropped in our analysis because of the poor image quality and sensitivity compared with the rest of epochs.

Table 2.1: Summary of Speckle Imaging Observations

Date		Frames		FWHM	Strehl ratio ^b	N _{stars}	K _{lim} ^c	σ _{pos} ^d	Data Source ^e
(U.T.)	(Decimal) ^a	Obtained	Used	(mas)	post- process		(mag)	(mas)	
1995 Jun 9-12	1995.439	15114	5286	46	0.62	380	16.4	1.47	Ref. 1
1996 Jun 26-27	1996.485	9261	2336	47	0.57	246	15.7	3.11	Ref. 1
1997 May 14	1997.367	3811	3486	46	0.65	358	16.4	1.29	Ref. 1
1998 May 14-15	1998.366	16531	7685	47	0.49	251	15.9	0.92	Ref. 2
1998 Jul 3-5	1998.505	9751	2053	42	0.85	226	16.2	0.86	Ref. 2
1998 Aug 4-6	1998.590	20375	11047	46	0.65	293	16.4	0.75	Ref. 2
1998 Oct 9	1998.771	4776	2015	47	0.52	216	16.0	1.32	Ref. 2
1999 May 2-4	1999.333	19512	9427	45	0.78	344	16.7	0.69	Ref. 2
1999 Jul 24-2	1999.559	19307	5776	44	0.77	303	16.8	0.37	Ref. 2
2000 Apr 21	2000.305	805	662	48	0.46	141	15.4	2.48	Ref. 3
2000 May 19-20	2000.381	21492	15591	45	0.62	402	16.9	0.56	Ref. 3
2000 Jul 19-20	2000.548	15124	10678	46	0.61	410	16.7	1.10	Ref. 3
2000 Oct 18	2000.797	2587	2247	47	0.46	209	15.8	1.70	Ref. 3
2001 May 7-9	2001.351	11343	6678	45	0.58	344	16.3	0.95	Ref. 3
2001 Jul 28-29	2001.572	15920	6654	46	0.73	351	16.9	0.57	Ref. 3
2002 Apr 23-24	2002.309	16130	13469	46	0.65	452	16.9	0.90	Ref. 3
2002 May 23-24	2002.391	18338	11860	44	0.74	436	17.1	0.58	Ref. 3
2002 Jul 19-20	2002.547	8878	4192	48	0.52	300	16.5	1.23	Ref. 3
2003 Apr 21-22	2003.303	14475	3715	48	0.53	185	15.5	1.25	Ref. 3
2003 Jul 22-23	2003.554	6948	2914	46	0.65	276	16.2	1.16	Ref. 3
2003 Sep 7-8	2003.682	9799	6324	46	0.67	356	16.6	1.80	Ref. 3
2004 Apr 29-30	2004.327	20140	6212	47	0.66	275	16.1	0.51	Ref. 4
2004 Jul 25-26	2004.564	14440	13085	47	0.61	379	16.9	0.90	Ref. 4
2004 Aug 29	2004.660	3040	2299	49	0.79	289	16.3	0.83	Ref. 4
2005 Apr 24-25	2005.312	15770	9644	47	0.54	282	16.3	0.70	Ref. 5
2005 Jul 26-27	2005.566	14820	5642	50	0.64	332	16.6	1.79	Ref. 5

^a Decimal year is defined as the Julian Epoch year: $2000.0 + (\text{MJD} - 51544.5)/365.25$;

^b Strehl ratio reported here is the post-process value from deconvolution method;

^c K_{lim} is the magnitude at which the cumulative distribution function of the observed K magnitudes reaches 90% of the total sample size;

^d Positional error taken as error on the mean from the three sub-images in each epoch and includes stars with K<15;

^e Data originally reported in (1) Ghez et al. (1998)[6], (2) Ghez et al. (2000) [7], (3) Ghez et al. (2005) [8], (4) Lu et al. (2005) [65], and (5) Rafelski et al. (2007) [66] ;

9, 27, 74, 39] and [12].

Image Processing

[12] combined the individual Speckle frames using a “holography” method [75] rather than the traditional shift-and-add method [17]. We improved this algorithm by using multiple reference stars, removing nearby confusion sources and subtracting sky background when extracting PSF. To estimate the positional errors, 100 realizations of each epoch were created using a bootstrap sampling with replacement to combine frames for that epoch. *Starfinder* is run on each of these realizations and the standard deviation of the positional measurements of stars are adopted as the uncertainty. We find that this method provides a more robust estimate of astrometric and photometric uncertainties [71]. The new holography method allows us to use more exposures and increases the sensitivity and field of view of the Speckle images, resulting in more stars detected at fainter magnitude. On average, 309 stars are detected in Speckle data sets down to a 90% limiting magnitude of $K_{\text{lim}} = 16.4$, with average position error of 1.1 mas in each direction.

The NIRC2 images were reduced using our standard NIRC2 reduction pipeline, which includes corrections for geometric distortion and differential atmospheric refraction. We used a new photometric calibration described in [76] to recalibrate all LGS-AO data sets. For observations between 2004-2013, we use the distortion solution from [51], while for observations obtained in 2014 and later, we use the new static distortion map from [52]. This change in optical distortion is a result of changes in the alignment of the AO system [52]. The output of the NIRC2 pipeline is a single combined image for each epoch of data along with three sub-set images used for error analysis [8]. Positional uncertainties, σ_{pos} , is the error on mean of the positions in three subset images for each star. Starlists containing astrometry and photometry were extracted using a PSF fitting algorithm *StarFinder* [77, 39]. On average, 1850 stars are detected in AO data sets down to a 90% limiting magnitude of $K_{\text{lim}} = 19.0$, with average position error of 0.09 mas in each direction. This exquisite precision is a result of the high signal-to-noise ratio (SNR). For a star at $K=15.5$ mag, the typical SNR is approximately 3000 and 8000 for Speckle and AO separately. Given that the $\text{FWHM}=45\text{-}65$ mas and the lowest possible centroiding error is $\sigma_{\text{pos}} \sim \text{FWHM}/\text{SNR}$, we could potentially reach positional uncertainties as low as 0.01 mas. However, the astrometric precision and accuracy is mostly limited by systematic errors due to uncertainties in the point-spread function (PSF) and noise from the halos of the surrounding stars [78].

2.2 Multi-Epoch Astrometric Alignment

The starlists from each epoch must be aligned (i.e. transformed and cross-matched) into an *absolute reference frame* in order to measure accurate proper motions, accelerations, and orbits. We define an *absolute reference frame* as described in [79] by measuring the proper motions of a set of IR stars that have been accurately transformed into a radio Sgr A*-

rest frame. The resulting catalog of *astrometric secondary standards* contains ~ 800 bright stars evenly distributed over the central $20'' \times 20''$. We adopt this catalog of positions and velocities as our *absolute reference frame*, and all of the deep, high-precision observations of the central $10'' \times 10''$ will be transformed into this common frame.

Our starlists are transformed in two steps:

1) We choose a reference epoch (2009 May 04, one of the well measured AO epochs), and transform starlists from other epochs into this reference epoch coordinate system by fitting a second order bivariate polynomial transformation. In this step, the set of *reference stars* used to calculate the transformation is selected from the catalog of *astrometric secondary standards* (see §2.2 and §2.2). The *reference stars* are first propagated from the reference epoch to other epochs using their known velocities from the *absolute reference frame* and the position angle and plate scale of the reference epoch. Then stars are matched between the propagated reference epoch and the starlist using a matching radius of 40 mas (i.e. ~ 4 pixel for AO and ~ 2 pixel for Speckle, see [27] and [51] for more details on matching). The best-fit transformation is then calculated by minimizing the residuals between the predicted positions and measured positions at each epoch for *reference stars*. At this point, all epochs are still in units of pixel in the reference-epoch coordinate system.

2) We then align the reference epoch coordinate (2009 May 04) to the *absolute reference frame* in order to transform all pixel positions into arcsec on the sky. This transformation is also accomplished by a second-order polynomial transformation in an iterative process using the 50 brightest stars first, and then increasing the number of stars in successive iterations as more stars are matched with a better transformation. This iterative process stops when the number of matched stars do not increase significantly.

We have implemented a number of improvements to our multi-epoch alignment process. First, in order to assess the impact of instrument and AO changes in the 24 years of astrometry, we establish a baseline alignment using a subset of data with identical instrumental and AO setups. This alignment of data from 2006 to 2014, which we call the *06-14* alignment, represents the highest precision subset of the astrometry. We use this data set to characterize systematic errors and stellar confusion events as described below.

In §2.2, we add a magnitude-dependent additive error to increase the accuracy of the astrometric errors; and we develop an algorithm to detect artifact sources on the edge of AO images due to elongated PSFs. In §2.2, we incorporate the remaining AO data by deriving local distortion maps to compensate for different experimental setups (e.g. NIRC2 data after 2014). Then, Speckle data is added to the alignment in §2.2. Finally, we utilize new thresholds to remove instances of stellar confusion in §2.2.

06-14 Alignment

Magnitude Dependent Additive Errors

The uncertainty of each star's position is comprised of 3 components: (1) σ_{pos} : the measurement precision on each star's position, (2) σ_{aln} : the uncertainty in the transformation

process for each star, (3) σ_{add} : an additive error term to capture additional sources of error. σ_{pos} is calculated as the error on the mean of a star's position measured from the three sub-set images as described in §2.1. The alignment error, σ_{aln} is estimated from a half-sample bootstrap analysis that repeats the alignment process 100 times with random sets of reference stars and is taken as the standard derivation of each star's position over the bootstrap samples. We found that the combination of σ_{pos} and σ_{aln} alone yields a χ^2 distribution on the acceleration fits with an unexpected tail of high values that are inconsistent with the standard χ^2 distribution. This is rectified with σ_{add} .

Here, χ^2 is defined as: $\chi_p^2 = \sum_{i \in \text{epochs}} (\vec{p}_i - \vec{p}_{\text{mod},i})^2 / \vec{\sigma}_{p,i}^2$, where $\vec{p}_i$ is the position at epoch i , $\vec{p}_{\text{mod},i}$ is the linear/acceleration fit at epoch i and $\vec{\sigma}_{p,i}$ is the positional uncertainty at epoch i (the quadratically-summed combination of σ_{pos} , σ_{aln} and σ_{add}). The linear/acceleration fit is discussed in detail in §2.3. We will use χ_{acc}^2 for χ^2 of acceleration fits and χ_{lin}^2 for χ^2 of linear fits. The fit in the x and y directions are independent of each other, so there are χ^2 values in each direction. Ultimately, the χ^2 distribution of a sample of stars is an important factor for determining the quality of our analysis.

The additive error term, σ_{add} , was previously determined to be 0.10 mas for AO data [39] and was assumed to be constant with time, position, and brightness. However, the most likely source of additional astrometric error is from inaccurate estimation of the PSF wings, which predominantly impacts the astrometry of neighboring sources. The impact of this effect is largest when the brightness of the neighboring source becomes comparable to the flux in the wings of the many surrounding bright stars. We improved the determination of σ_{add} by implementing a magnitude-dependent error term for AO data that is added in quadrature, as described below [see also 50, 80].

Defining the Sample of Good Stars In order to evaluate the χ_{acc}^2 distribution and determine the optimal additive error, σ_{add} , we define a sample of *good stars* chosen as those that (1) were detected in all 22 epochs of the *06-14* alignment, (2) showed no source confusion in any epoch, as defined in §2.2, (3) were located between 0.8'' and 10'' from the SMBH in order to eliminate stars close to the SMBH that show high-order motion beyond acceleration, (4) were not outliers with large acceleration uncertainties, and (5) had χ_{acc}^2 smaller than 95 (5 times the *dof*), where *dof* is defined as the number of detections minus the number of free parameters. For example, here we have 22 epochs in total, and we use an acceleration fit with 3 free parameters per axis thus the *dof* per axis equals to 22-3=19. The resulting sample of 370 *good stars* will be used to evaluate the χ^2 distribution for different additive errors.

Calculating the Additive Error To find the optimal σ_{add} term, we perform an *06-14* cross-epoch alignment with additive errors from 0.01 mas to 0.47 mas in steps of 0.01 mas. We then divide the *good stars* into 9 magnitude bins with ~ 41 stars per bin. The magnitude boundaries are: 10, 12.55, 13.61, 14.25, 15.11, 15.39, 15.65, 15.87, 16.52, and 18.72. For each magnitude bin and each trial additive error, the χ_{acc}^2 distribution of *good stars* is calculated

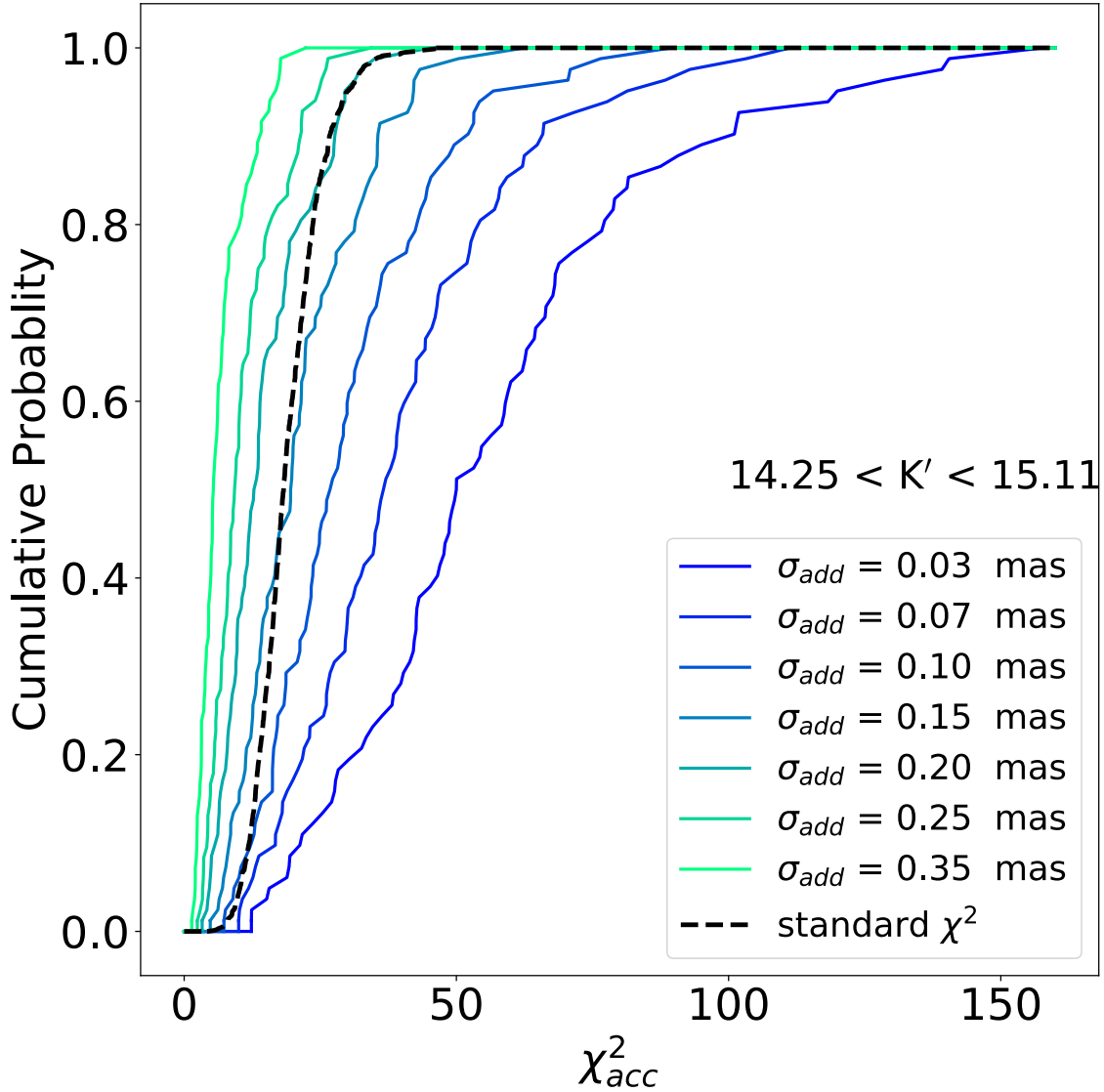


Figure 2.1: The cumulative χ_{acc}^2 distribution for different σ_{add} from the *good stars* in the magnitude bin, $14.25 < K < 15.11$. Different colored lines represent different σ_{add} added to the positional uncertainty in quadrature. With larger σ_{add} , the χ_{acc}^2 distribution shifts to smaller values. The standard cumulative χ_{acc}^2 distribution with $dof = 19$ is plotted in the black dashed line. The optimal σ_{add} of 0.15 mas produces a χ_{acc}^2 distribution that most closely matches the expectation.

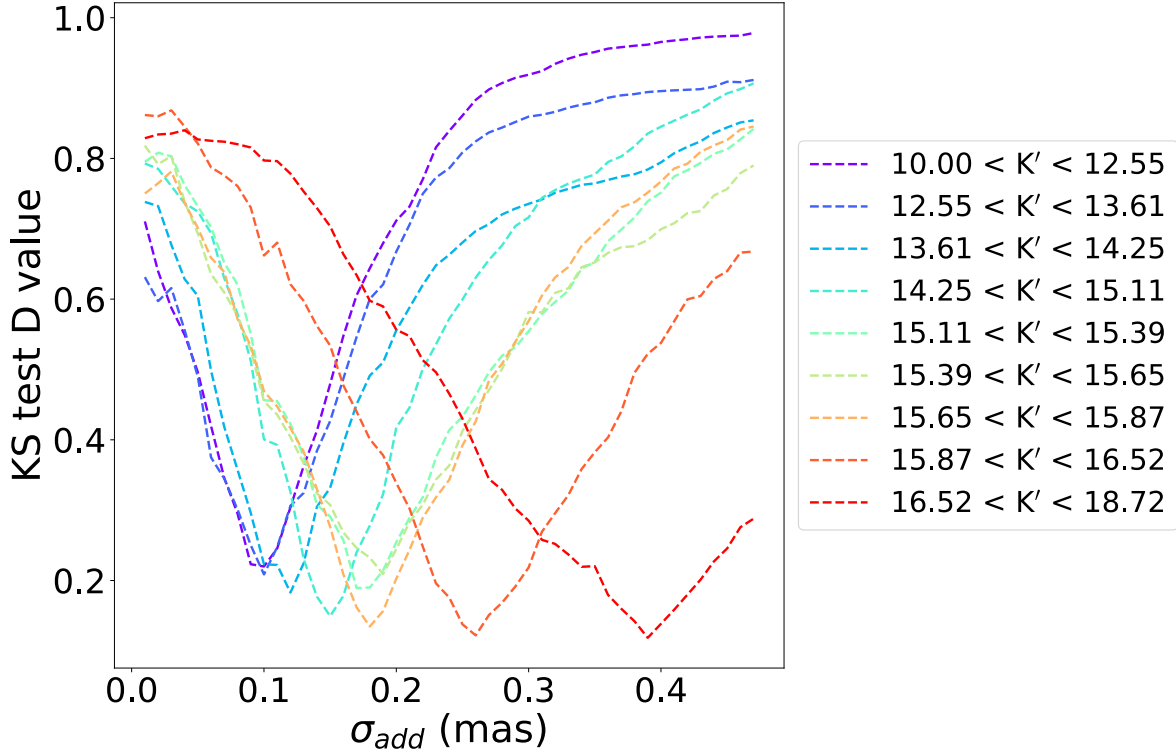


Figure 2.2: The KS-test between the observed χ_{acc}^2 distribution and a standard χ_{acc}^2 distribution. The KS test D value is plotted as a function of additive error. The smaller the D value is, the better agreement between the observed and standard χ_{acc}^2 distribution. We can see that for fainter stars, they need larger additive error to fit the standard χ_{acc}^2 distribution.

(Figure 2.1). Increasing the additive error causes the χ_{acc}^2 distribution to shift to smaller values.

The optimal additive error should exhibit a χ_{acc}^2 distribution that closely matches the standard χ_{acc}^2 distribution with dof=19. We perform a two-sample Kolmogorov-Smirnov test (KS-test) between the observed χ_{acc}^2 distribution and a standard χ_{acc}^2 distribution for each magnitude bin [81]. The KS test statistic (D value), which quantifies a distance between the observed χ_{acc}^2 distribution and the cumulative distribution function of the standard χ_{acc}^2 distribution, is plotted as a function of additive error in Figure 2.2. The smaller KS statistic D value suggests a better agreement with the theoretical predicted χ_{acc}^2 distribution. There is a clear trend showing that fainter stars require larger additive errors.

In Figure 2.3, we plot the best fit additive error in each magnitude bin. We derive an analytic function for the additive error as a function of magnitude given as

$$\sigma_{add}(mas) = 8 \times 10^{-5} \times e^{0.7 \times (K' - 4.9)} + 0.1 \quad (2.1)$$

With this magnitude-dependent σ_{add} , we now evaluate the contribution of each source

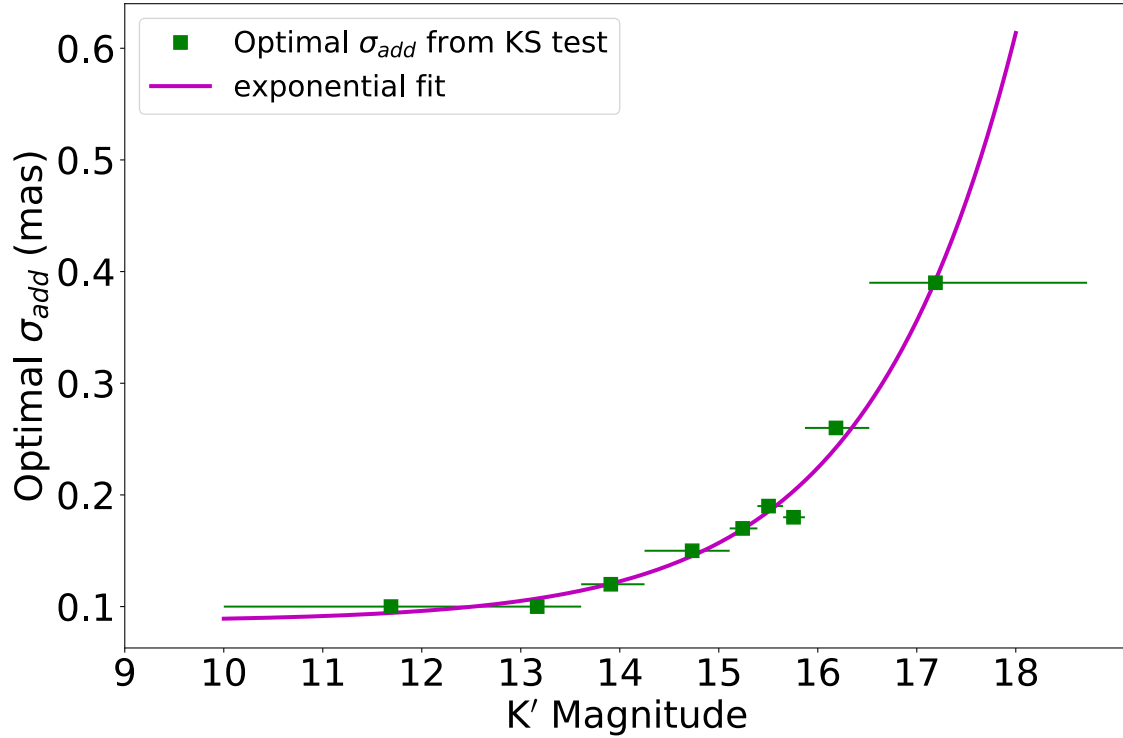


Figure 2.3: The optimal astrometric additive error, σ_{add} , as a function of brightness. The green squares are the optimal σ_{add} from minimizing KS test D value from Figure 2.2. The error bar shows the range of each magnitude bin, and the mean of each magnitude bin is showed on the green square. Based on the trend, we use an exponential function shown in Equation 2.1 to fit the observed σ_{add} as a function of K' . The best-fit exponential relation is shown in the magenta line.

of positional uncertainty to the total. The 2006 May 03 epoch (one of the *06-14* epochs) is used as an example to show how σ_{pos} , σ_{aln} , and σ_{add} change with magnitude in Figure 2.4. We can see that both σ_{pos} and σ_{add} increase with magnitude, almost exponentially, and they are comparable with each other. In contrast, σ_{aln} captures the transformation uncertainty and is around 0.1 mas with a weak dependence on magnitude. This is likely due to the fact that σ_{aln} is slightly larger in the outskirts of the field, where the faint stars are more easily detected away from the central concentration of bright stars.

This magnitude-dependent additive error is incorporated in for all other AO epochs. The χ^2_{acc} distribution for *06-14* alignment with this magnitude dependent additive error is plotted in Figure 2.5 and shows good agreement with the standard χ^2_{acc} distribution.

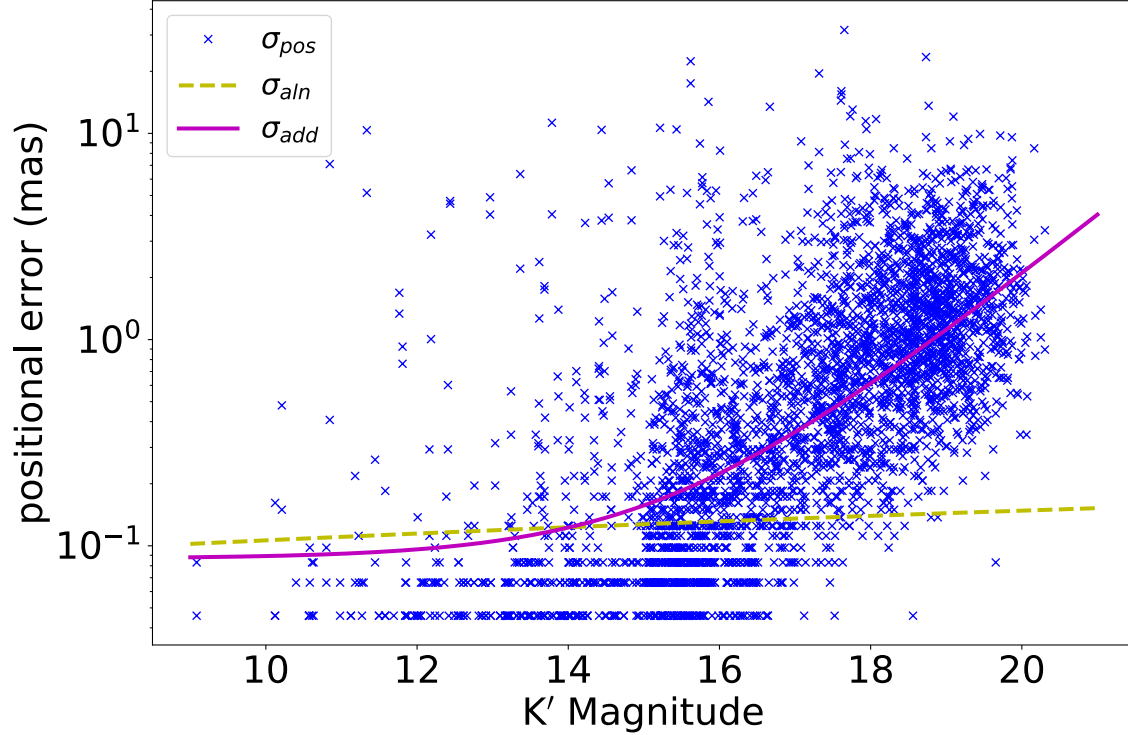


Figure 2.4: Different sources of positional uncertainty as a function of K' magnitude for 2006 May 03. The blue crosses represents σ_{pos} in both the x and y direction. The magenta solid line represents σ_{add} , which is the same as showed in Figure 2.3. The yellow dashed line represents σ_{aln} , which comes from the alignment transformation uncertainty. The σ_{aln} curve stays around 0.1 mas for all stars, almost independent of K' . σ_{pos} and σ_{add} is 0.1 mas for the bright stars and then increases rapidly for fainter stars. The turnover happens around $K' = 15$. Note that our positional uncertainty measurement precision is 0.01 mas, so the smallest σ_{pos} looks discrete.

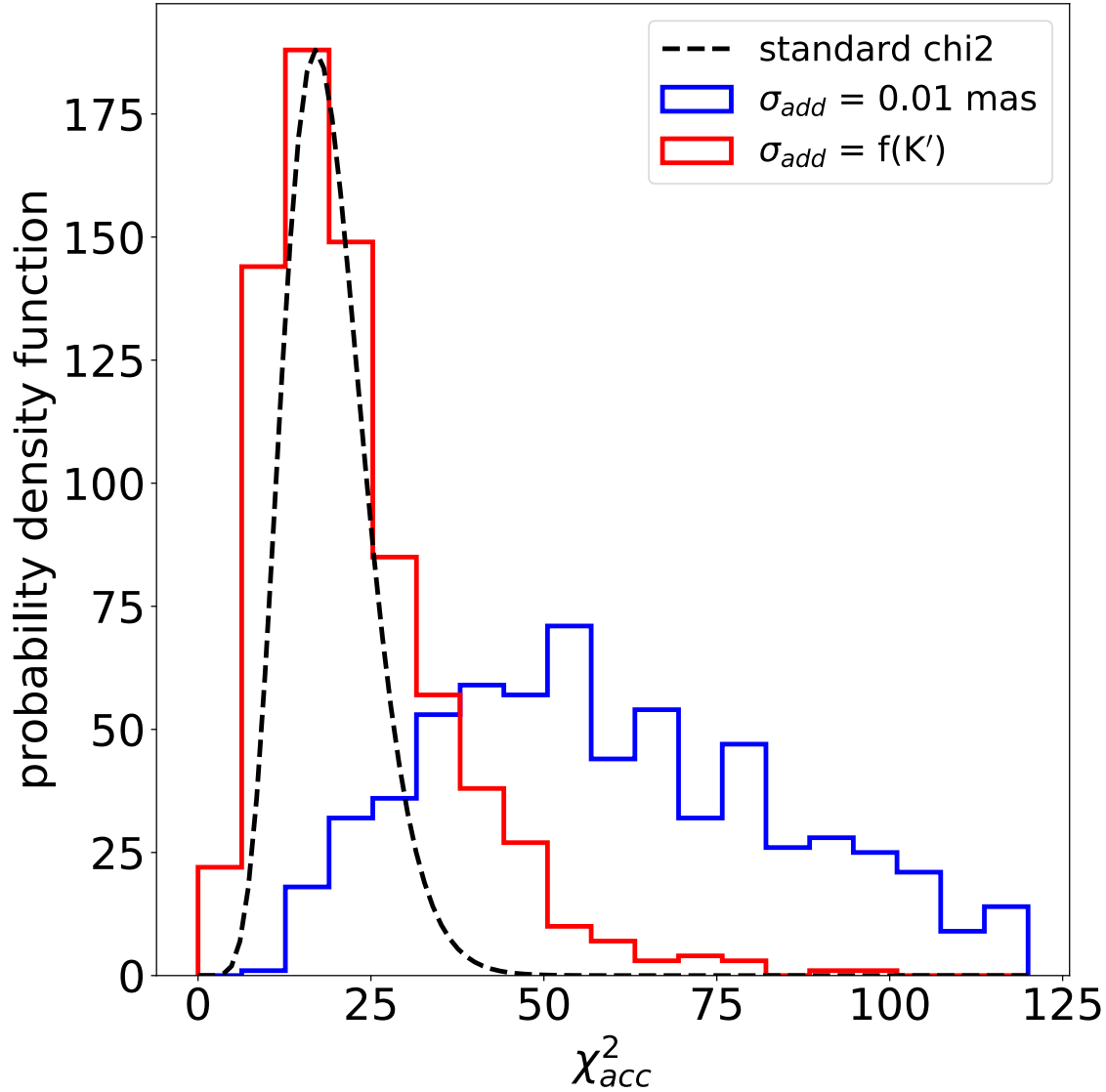


Figure 2.5: The χ_{acc}^2 probability density function (PDF) for the *06-14* alignment. The blue histogram is the original χ_{acc}^2 distribution with 0.01 mas additive error, which is approximately the χ_{acc}^2 distribution without any additive error. The red histogram is the final χ_{acc}^2 distribution with optimal magnitude dependent σ_{add} from Equation 2.1. The black dash line is the standard χ_{acc}^2 distribution with $dof = 19$, normalized to the maximum value of observed χ_{acc}^2 PDF.

Reference Stars for AO

The selection of *reference stars* are critical as they are used to bring each epoch of observations to the Sgr A*-radio rest frame of reference. Ideally, we would like to maximize the number of reference stars to yield more accurate transformations between individual epoch and reference epoch (2009 May 04). However, in order to define the positions of these stars for each epoch, we want the reference stars to move linearly with accurate velocity measurements, since we assume a linear motion model to propagate between epochs.

From §2.2, the sample of *good stars* consists of 370 stars and their χ_{acc}^2 distribution is well-behaved. Among these 370 *good stars*, 18 have non-SMBH accelerations: significant tangential accelerations, significant positive radial accelerations, or negative radial accelerations that are higher than physically allowed from the SMBH (see details in §2.3). These non-SMBH accelerations likely come from confusion or binarity, so we exclude these non-SMBH-acceleration sources, leaving a clean sample of 352 stars. In order to obtain accurate velocity measurements in an absolute reference frame, we find the intersection of these 352 *good stars* and the *astrometric secondary standards*, which gives us 141 stars. These 141 stars are used as *reference stars* in order to derive the coordination transformations for all AO epochs.

Artifact Sources on the Edge of AO Epochs

The Keck AO observations deliver near-diffraction-limited spatial resolution; however, the PSF for the AO images is highly structured and varies across the field of view. We found that, near the edges of images from AO epochs, the PSF becomes elongated and structures in the first Airy ring are sometimes identified as a separate source. Figure 2.6 shows an example.

Active work is underway to account for PSF variations across the field [e.g. the AIROPA project 82]. However, in this paper, we use a simple yet straightforward way to mark and remove those artifact sources. Artifact sources have some common properties: (1) artifact sources are usually within 70 mas of their primary sources. (2) because the artifact sources are typically in the same relative position compared to their primary sources, the proper motions for the artifact sources and primary sources are similar. (3) the PSF elongation is usually in the same direction as the separation vector between the primary and artifact sources. Combined, these three properties enable us to find potential artifact sources coming from PSF field variability. Figure 2.7 uses epoch 2011 July 18 as an example to show how this correction works. First, we need to find pairs of sources based on the first 2 properties: we select every pair of stars that are within 70 mas of each other and have similar proper motions ($\delta\mu < 3 \text{ mas yr}^{-1}$). Each blue cross is a detected star in epoch 2011 July 18, and the red points are the pairs pass our selection criteria. The positional offset between each pair is shown as black and red arrows. The offsets are all in similar directions in the lower right corner, which is a strong evidence that they are artifact sources coming from PSF wings, and this agrees with what we see in Figure 2.6. For those stars that have the same offset

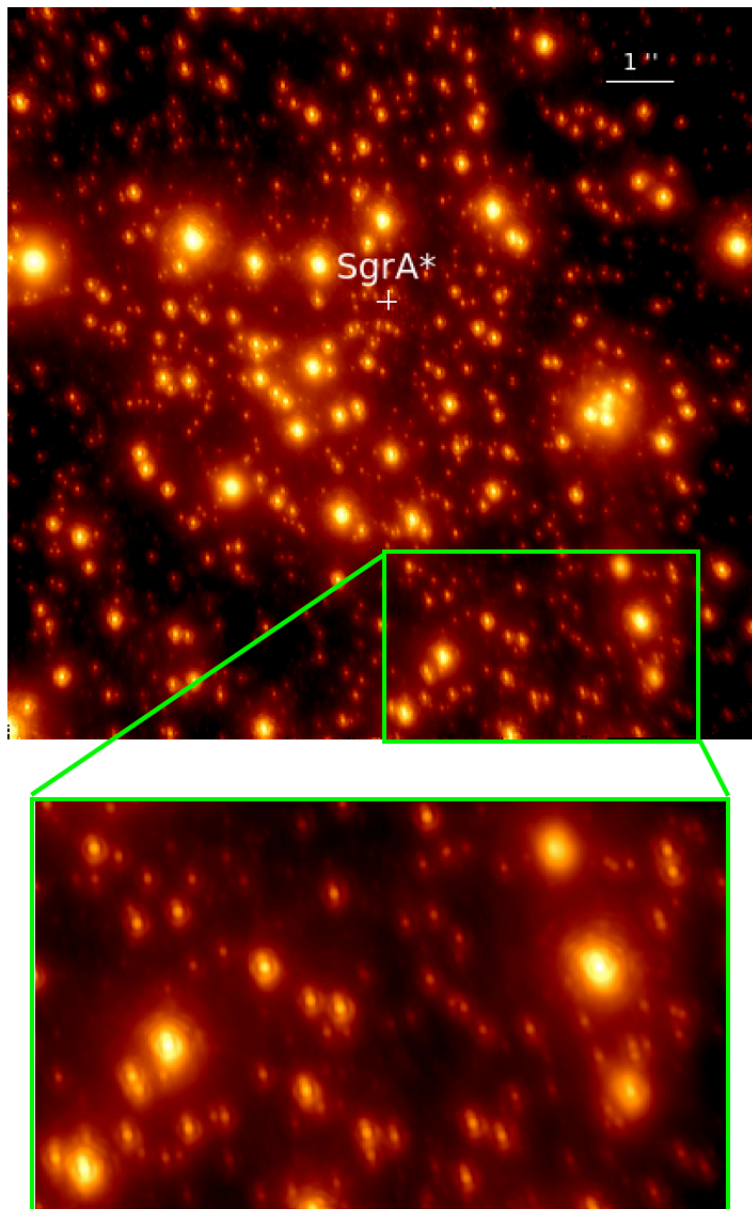


Figure 2.6: An image for epoch 2011 July. Stars in the central region are more circular as compared to stars in the lower right corner, which are all elongated in a similar direction. The PSF is most elongated in the direction away from the tip-tilt star.

direction, we will mark them as PSF artifact sources, which is shown in the red arrows in Figure 2.7. Among 2729 sources, 88 sources are found to be artifact sources in 26 AO epochs, accounting for about 3.2% of the total number of sources. The artifact sources are excluded from our final sample.

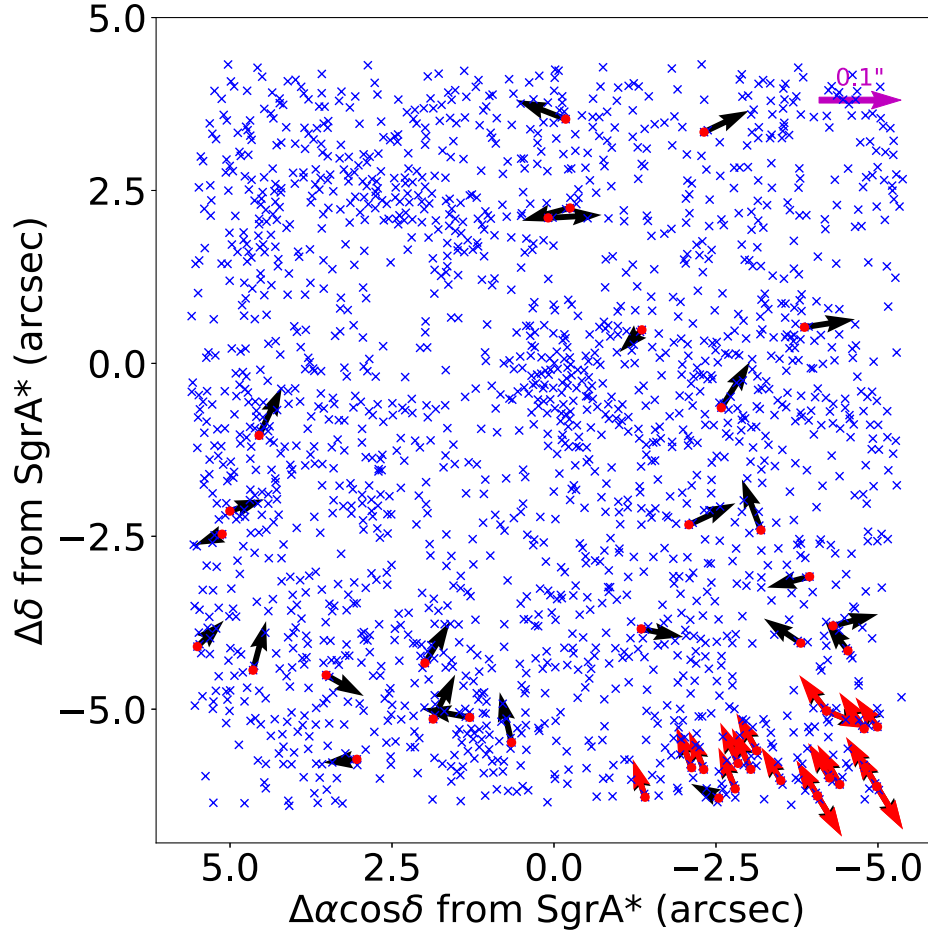


Figure 2.7: Artifact sources due to PSF field variability. Each blue cross is a star detected in epoch 2011 July 18. Then each star is paired up with every other star and the red points are marked when a pair of star has a positional offset smaller than 70 mas and a proper motion difference smaller than 3 mas yr^{-1} . The position offsets between those pairs are plotted in the black and red arrows. If a pair is just a coincidence, then the arrows should be randomly distributed; however, we see many arrows in the lower-right corner have the same direction, which suggests that PSF elongation has created these artificial sources. Those arrows with the same direction are colored in red as artifact sources which are excluded from the final sample.

Update Matching Velocity from *06-14* Alignment

The epochs from 2006 to 2014 are taken with exactly the same instrumental setup. As a result, these data give a much more precise velocity measurements compared to [39]. Therefore, we use the velocity from *06-14* alignment to update the velocities for those stars that are not *astrometric secondary standards* but still need to be matched based on their matching velocity. Matching velocities are updated for stars if: 1) Their projected distance to Sgr A* is larger than $0.4''$. This is because stars within $0.4''$ need a model beyond simple linear motion. 2) Their K band magnitude is brighter than 16 magnitude. 3) Their velocity uncertainty from *06-14* alignment is smaller than 2 mas yr^{-1} . 556 stars' velocities are updated from the *06-14* alignment. Comparing with [39], the median velocity uncertainty for these 556 stars is reduced from 0.14 mas yr^{-1} to 0.03 mas yr^{-1} , by almost a factor of 5.

Adding Other AO Epochs: Local Distortion

Among the 30 AO epochs, 8 were not taken in the standard *06-14* setup including the 2005, 2015, 2016, and 2017 epochs. They have higher order residual distortion from changes in the AO system optics, so the standard second order polynomial transformation is insufficient to place these images into a common reference frame. Therefore, we need to make local distortion maps for these epochs. Our local distortion maps are calculated based on residuals, which is different from geometric distortions based on the on-sky measurements in [51] and [52]. The way we calculate the distortion map is described as follows.

First, the *06-14* alignment is used to fit acceleration models for stars within $0.8''$ and linear models for stars outside of $0.8''$ from Sgr A*. With the acceleration/linear fit, all stars are propagated to each of the non-standard epochs listed above. The differences between the propagated positions and the observed positions at that epoch are used to estimate the local distortion map. When calculating the distortion map, the following cuts are made to reduce noise in the final distortion map: 1) Only stars brighter than 17 mag are used. 2) Outliers are removed by dividing the detected stars into 6×6 spatial around the position of Sgr A*, calculating the mean and standard deviation of the distortion in each box, and trimming stars with offsets that are more than 2.5σ from the mean. This gives us the final sample that is used to calculate the local distortion map. On average, 530 stars are used in the final sample.

A high order Legendre transformation is fit to the residuals to construct the local distortion map. To calculate the uncertainty of the local distortion map, we use the standard deviation among 100 local distortion maps estimated from a full-sample bootstrap with replacement. The distortion and its uncertainty can become very large on the edge of the distortion map as this region is sometimes outside the field of view for a given epoch and the distortion must be extrapolated from very few stars. Therefore, for those “edge” values, whose distortion or distortion uncertainty is larger than 0.3 pixel, we use the nearest “non-edge” value to replace it.

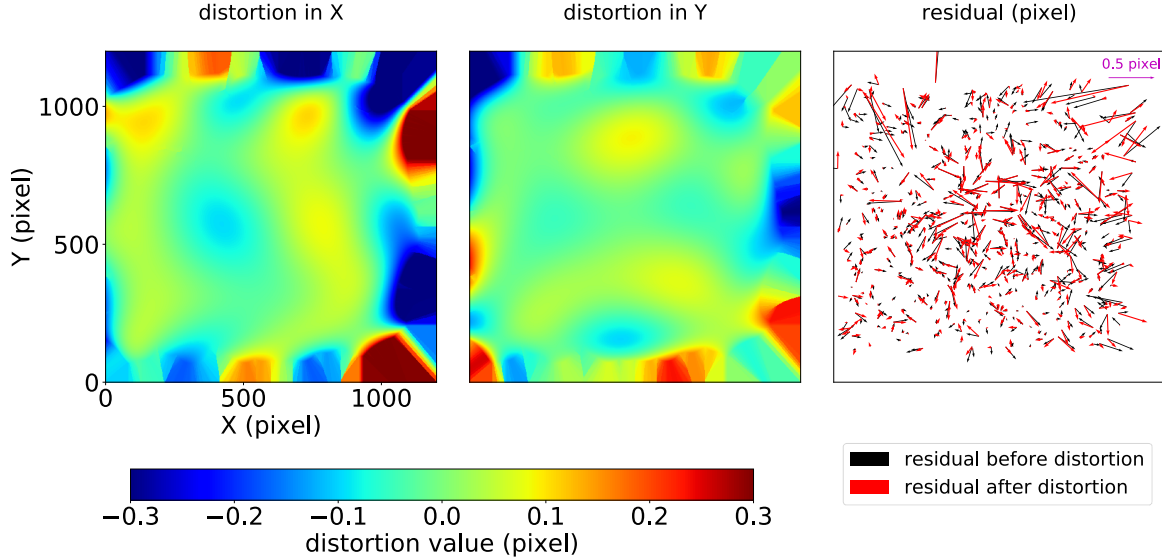


Figure 2.8: The local distortion map for 2016 May 03 in the x direction (*left*) and y direction (*middle*), color coded by the distortion value. On the *right*, the individual star’s residuals from the multi-epoch acceleration fit are shown both before (*black*) and after (*red*) the local distortion solution is applied. The arrows are the offset between the *06-14* alignment propagated position and the 2016 May 03 position. The distortion map is the Legendre transformation calculated from the black arrows. Note that there are very few stars on the edge, so we use the nearest distortion value to extrapolate into those regions. The red arrows are reduced by a factor of $1/3$ compared to the black arrows, especially at corners, indicating the effectiveness of the distortion map.

Finally, we need to find the best Legendre transformation order for our distortion map. The residuals will always improve as we increase the transformation order; but the number of free parameters also increases for a higher order transformation. So we use the F-ratio test to find when an increase in the Legendre transformation order no longer *significantly* improves the residuals. This is done by finding the point when $(1-p)$ value for F-ratio approaches 0. The F-ratio is calculated when increasing Legendre transformation order, and the p -value is the probability of obtaining this F-ratio (see Equation (B2) in [83]). Lower p -values indicate more significant benefits to increasing the order of the transformation polynomials. A similar test has been done in [83].

In this way, we derive the final local distortion maps for all non-*06-14* AO epochs in arcsec. The distortion maps are converted into pixel coordinates for each image using the previously derived transformations. The local distortion map for 2016 May 03 is shown in the first two panels of Figure 2.8 as an example. We apply the local distortion to the stars’

original position and add the distortion uncertainty in quadrature. The 3rd panel of Figure 2.8 shows how the residuals are reduced between the *06-14* alignment propagated starlist and 2016 May 03 starlist after applying the distortion map. The median residual has been reduced from 0.06 pixel (black arrows) to 0.04 pixel (red arrows).

Table 2.3 summarizes the local distortion median value and typical uncertainties for all 8 non-06 epochs. To compare with the previously mentioned positional uncertainties in §2.2, we plot the uncertainties as a function of time in Figure 2.9. From this plot, we can see AO epochs have reduced σ_{pos} and σ_{aln} by almost an order of magnitude relative to Speckle epochs. For Speckle epochs, σ_{pos} is slightly larger than σ_{aln} and both contributes the total positional error. For AO epochs, σ_{add} and σ_{pos} are constant among epochs, but σ_{aln} increases with time after 2011, the reason being our reference epoch is 2009, so alignment transformation gets worse when further away from reference epoch. For epochs with local distortion map, local distortion map error σ_{dist} is slightly larger than other uncertainties.

Adding Speckle Data

New Holography Data In the new Holography analysis (see §2.1), the stars' positional uncertainties are more accurately measured with bootstrapping, which captures unknown uncertainties from confusion or imperfect PSF. So there is no need to add extra additive error. The typical positional error for bright stars within $2''$ from Sgr A* is plotted in Figure 2.9.

Reference Stars for Speckle From §2.2, we have 141 *reference stars* for AO epochs, but the Speckle data has a smaller field of view and shallower detection limit. Therefore we make a radius cut of stars within $4''$ from Sgr A*. This gives us 43 stars out of 141 stars. More importantly, the Speckle images were not taken in stationary mode, so the field changed over night. Stars on the edge of Speckle images will have less frame coverage compared with stars in the inner region, therefore they have relatively poor astrometric measurements. To account for this effect, we require stars to be detected in more than 60% of frames in Speckle epochs relative to IRS16C, which is one of the brightest and cleanest star in our field of view. This 60% criteria is lower compared to [12] because new Holography has more frames. The spatial distribution of AO and Speckle *reference stars* is plotted in Figure 2.10.

In conclusion, we have 141 *reference stars* for AO epochs and 43 *reference stars* for Speckle epochs. This is less than what [12] used in her paper, mainly because we pursue the high quality of *reference stars* over the quantity. On average, we have 21 stars used as *reference stars* for Speckle epochs and 133 stars for AO epochs, while [12] has 61 stars for Speckle epochs and 230 stars for AO epochs.

Speckle Edge removal Since the number of frames that contribute to a given pixel near the edges of the Speckle images can be very low because of the field rotation, stars would have poor astrometric measurements on the edge. Therefore we decided to remove detections

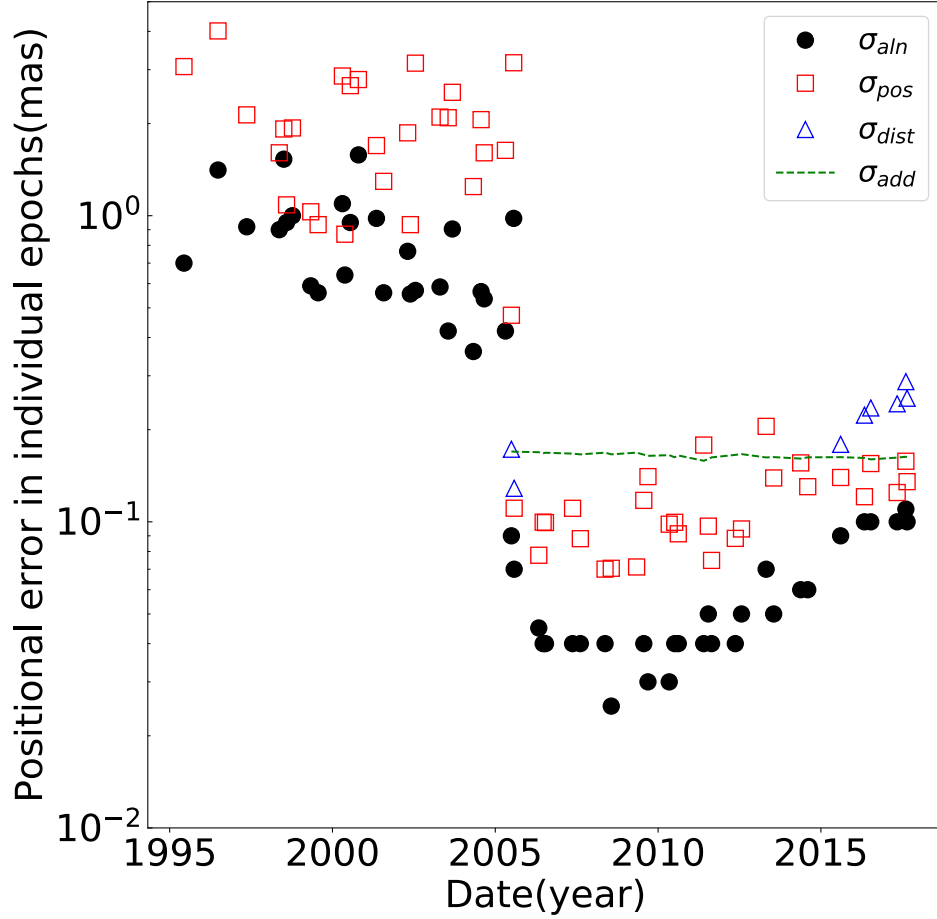


Figure 2.9: The different types of positional errors in individual epochs. Here the positional uncertainty is the median value for stars brighter than 16 magnitude and within $2''$ from Sgr A*. We only choose those stars because Speckle epochs have a smaller field of view and are much shallower (see details in §2.2). Red open squares, black filled circles and Green dashed lines are σ_{pos} , σ_{aln} , σ_{add} respectively (see details in §2.2). Blue open triangles show the distortion error added to non-06 epochs, σ_{dist} . Notice that σ_{pos} and σ_{aln} exist for both Speckle and AO epochs, but σ_{add} only exist for AO epochs (see §2.2) and σ_{dist} only applies to epochs with local corrections.

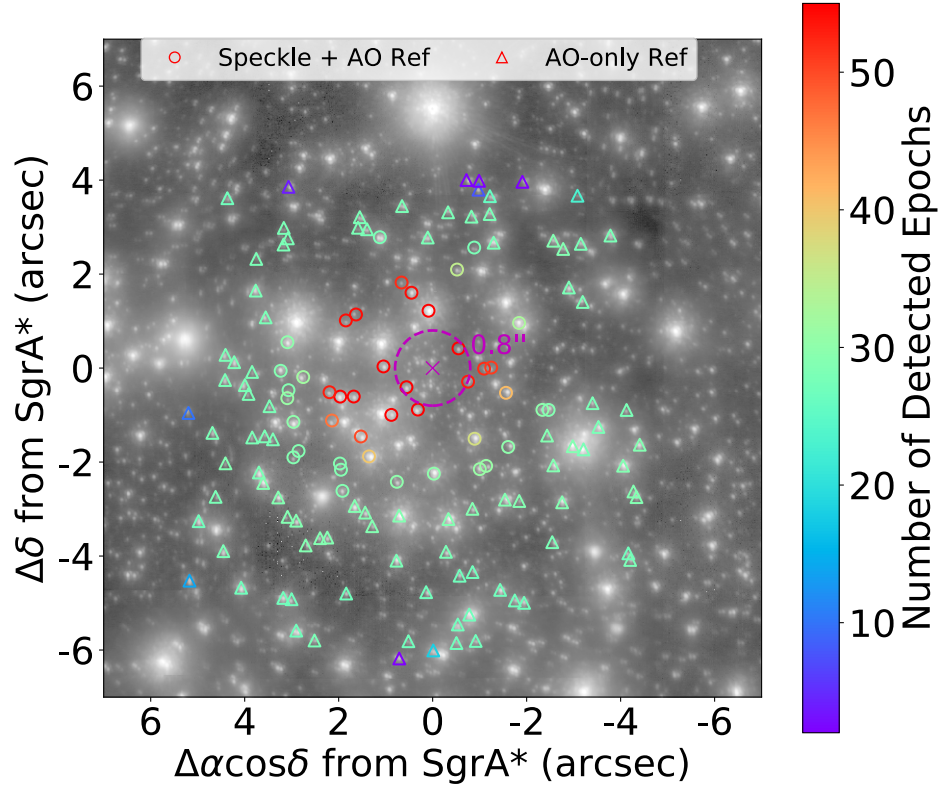


Figure 2.10: The spatial distribution of *reference stars* color coded with the number of epochs when they are used for calculating transformation. The triangles are AO only *reference stars* while the circles are Speckle+AO *reference stars*. Most of the *reference stars* out of $4''$ are only used in less than 30 epochs, because those stars are not in Speckle’s field of view, and are only detected AO epochs. *reference stars* are also required to be detected in more than 60% of frames compared with IRS16C, in order to make sure they have reliable position measurements. So there are a few stars in the outer region which are only used in less than 10 epochs.

without enough frame coverage (less than 60% frames relative to IRS16C) for Speckle epochs. As a result, 3601 detections from 1020 stars were removed, which is almost 44.8% of all stars over all Speckle epochs.

Improved Matching in Crowded Region: Confusion Removal

Up to this point, 2,700 stars are identified with a total of 56,061 measurements across 56 epochs. However, some measurements are biased or incorrect due to mis-matches from stellar crowding and confusion. When matching starlists from different epochs, we require that a new measurement must fall within a radius of 40 mas of the predicted position. However, when two stars get too close to each other (e.g. within the first airy ring), *StarFinder* cannot easily distinguish them. In this case, only one source will be detected and the position for this source is biased as it is the flux-weighted average of the two stars. Fortunately, the probability of mis-matches for any one star changes with time as stars move past each other in projection; so the individual stars are not entirely lost if we can remove the biased epochs.

We choose to remove instances of confusion and mis-matching using the following method: For each star A, we find nearby stars that are within 100 mas. For every nearby star B, if star B is 5 mags fainter than star A or brighter, then star B is defined as a *possible confusion source*. Notice that when star B is fainter than star A by more than 5 magnitude, the confusion from star B will not affect star A's position in a significant way. Both star A and star B are required to be detected in more than 10 epochs for a reliable proper motion measurement. For star A and its possible confusion sources, if only one source is detected in some epoch where there should be two based on their proper motion prediction, this detection will be removed. Figure 2.11 shows an example on confusion removal. In total, 5677 detections are removed because of confusion, affecting 751 stars, accounting for 10% of the total detections.

2.3 Proper Motions, Accelerations and Orbits

With the well measured astrometric positions from §2.2, we can measure the proper motion, acceleration and orbits for a final sample of 1148 stars. In order to obtain a precise estimate of the proper motion, we require stars to be detected in more than 20 epochs (2/3 of all AO epochs), which gives us a sample of 1184 stars. In this sample, the median positional uncertainty is 2.37 mas for speckle epochs and 0.25 mas for AO epochs. The positions over time are then used to fit a kinematic model for each star consisting of either a first-order (linear) polynomial, a second-order (acceleration) polynomial, or a full Keplerian orbit. We use a full orbit fit for the 33 stars within $0.5''$ of the SMBH. For the stars outside $0.5''$, we first fit a second-order polynomial fit to derive the acceleration for all stars. A jackknife is used to get a robust estimate of the acceleration uncertainty. We find 103 significant accelerating stars at the $> 5\sigma$ level, which are then further divided into SMBH-acceleration sources and non-SMBH-acceleration sources in §2.3. Then the remaining 1048 stars are fit

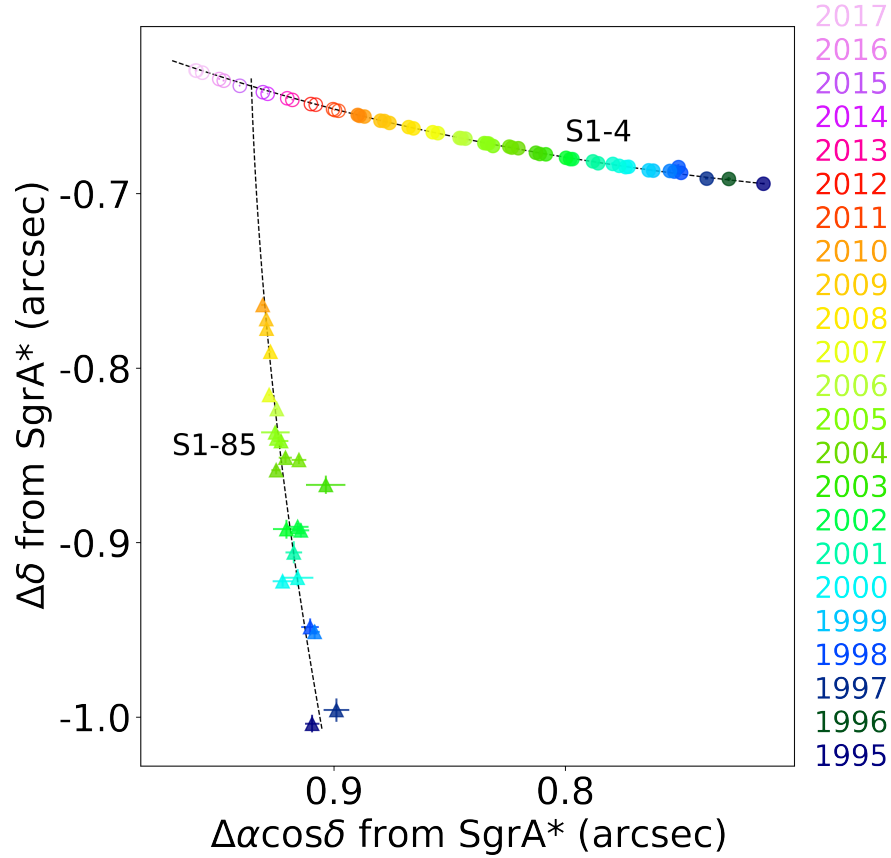


Figure 2.11: An example of how epochs with potential source confusion are removed: circles represent star S1-4 ($K'=12.3$) and triangles represent star S1-85 ($K'=15.3$). Different color shows stars' positions over time from 1995 to 2017. The two stars move closer to each other. From 1995 to 2010, S1-4 and S1-85 are separated enough to both be detected. But after 2011, they are too close to be distinguished and only S1-4 is detected. In this case, we remove those detections for S1-4 after 2011 as shown in the open circles, because the position for S1-4 is biased by S1-85 in those epochs.

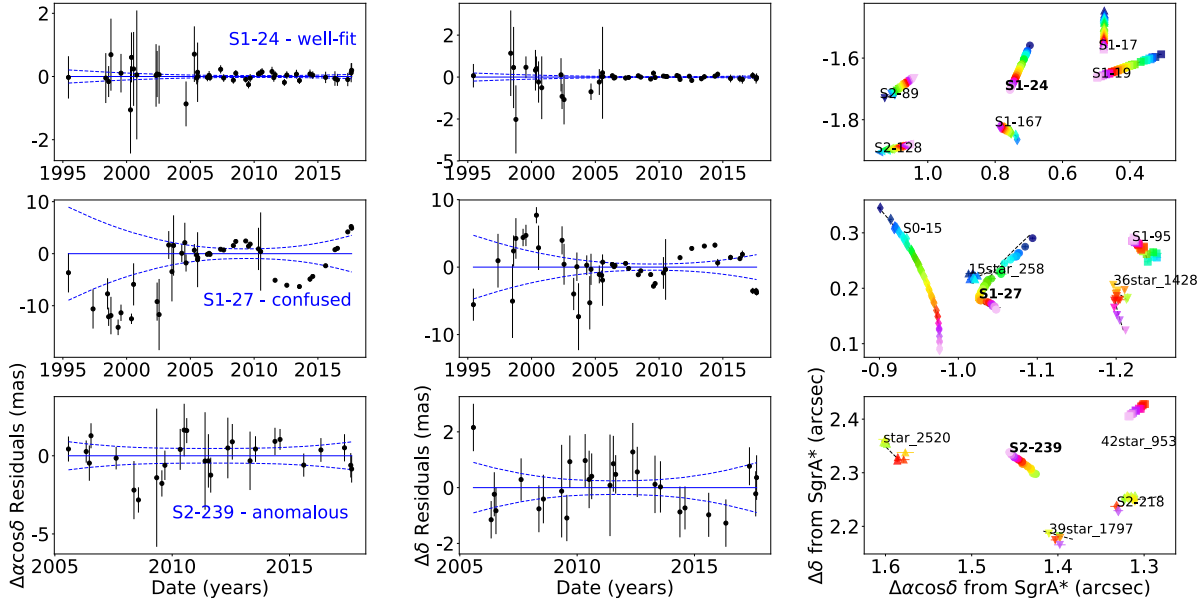


Figure 2.12: Examples of non-SMBH-acceleration sources from 3 different categories: “well-fit” (*top* row), “confused” (*middle* row) and “anomalous” (*bottom* row), which is defined by how well the acceleration fit is and whether there are confusion stars nearby. Here, the first two columns plot the residuals from acceleration fit in $\alpha \cos \delta$ and δ direction, and the third column plots the nearby stars.

with a first-order linear motion model. The median velocity for those 1048 linear moving stars is 0.05 mas yr^{-1} in each direction.

SMBH-acceleration and Non-SMBH-acceleration sources

The sample of 103 stars with significant ($> 5\sigma$) accelerations are analyzed to determine whether the best-fit accelerations are consistent with the expected acceleration from the SMBH (i.e. negative radial acceleration). The accelerations are projected into the radial (a_r) and tangential (a_t) directions w.r.t. the SMBH. Significant accelerating sources are defined when one of the following conditions is satisfied: (1) radial acceleration is 5 sigma larger than its uncertainty: $|a_r/\sigma_{a_r}| > 5$. (2) tangential acceleration is 5 sigma larger than its uncertainty: $|a_t/\sigma_{a_t}| > 5$.

Then, we define a sample of significant *SMBH-acceleration* sources when all of the following conditions are satisfied: (1) significant negative acceleration: $a_r/\sigma_{a_r} < -5$, (2) tangential acceleration is consistent with zero: $|a_t/\sigma_{a_t}| < 3$, (3) negative radial acceleration is 3σ smaller than allowed: $(a_{r,max} - a_r)/\sigma_{a_r} < 3$, where $a_{r,max} = -GM/r^2$ is the maximum allowed radial velocity from the SMBH, M is the mass of the SMBH and r is the 2D projected distance.

Since we don't have the line-of-sight distance, r will be the lower limit of the real 3D distance, so $a_{r,max}$ will be the upper limit of the allowed radial acceleration a_r . These criteria result in 27 significant SMBH-acceleration sources.

The remaining 76 accelerating sources are all *non-SMBH-acceleration* sources, including significant tangential accelerations, significant positive radial accelerations and too-large negative radial accelerations. The non-SMBH-acceleration sources are likely due to a number of factors including unrecognized confusion and binarity. A potential binary candidate is analyzed in §2.5.

The large number of non-SMBH-acceleration sources suggests that there may be some contaminants in the SMBH-acceleration sample. To determine the degree of contamination, we check all 103 stars by eye, and divide them into 3 categories. 1) "well-fit": stars that show no time-coherent residuals from the acceleration model. 2) "confused": stars that are not fit well by an acceleration model, but show potential confusion from neighboring stars. 3) "anomalous": stars that are not fit very well by acceleration, and show no potential confusion around them.

Figure 2.12 gives an example for each category from non-SMBH-acceleration sources. The first two columns are residuals from the acceleration fit and the third column shows the proper motion of nearby stars. To determine the category an accelerating source belongs to, we first look at the quality of the acceleration fit. In the figure, we can see S1-24 is fit well by an acceleration model given that the residuals are randomly distributed with time; so S1-24 is a "well-fit" non-SMBH-acceleration source. If a star is not fit well by an acceleration model, like S1-27 and S2-239, we consider alternative explanations: confusion from nearby stars, bad measurements, or other physical explanations such as astrometric wobble due to binarity or microlensing. To exclude confusion, we check whether there are nearby stars that are not detected in all epochs (note that in §2.2, confusion events are removed for cases when both stars are detected in more than 10 epochs, so stars detected in fewer epochs will be missed as potential confusion sources). S1-27 has a nearby star 15star_258 which is only detected in a few epochs, so S1-27 is potentially confused with 15star_258 when 15star_258 is not detected². In comparison, S2-239 is very isolated. Therefore, S1-27 is classified as a "confused" non-SMBH-acceleration source and S2-239 is an "anomalous" non-SMBH-acceleration source.

In summary, Table 2.4 shows the number of accelerating sources in different categories. Figure 2.13 plots the spatial distribution of stars from: orbit(33), acceleration (103) or linear motion (1048), where acceleration stars are further divided into the categories as showed in Table 2.4. The 24 well-fit SMBH-acceleration sources are particularly interesting and are explained in detail in §2.4. The 76 non-SMBH-acceleration sources require further analysis, in which S0-27 is used as an example to explore the potential binarity in §2.5.

²Notice that all stars in the final sample have names such as "SRR-NN", where RR is the radius this star belong to. Stars which are not in our final sample have temporary names such as "NNstar_NN".

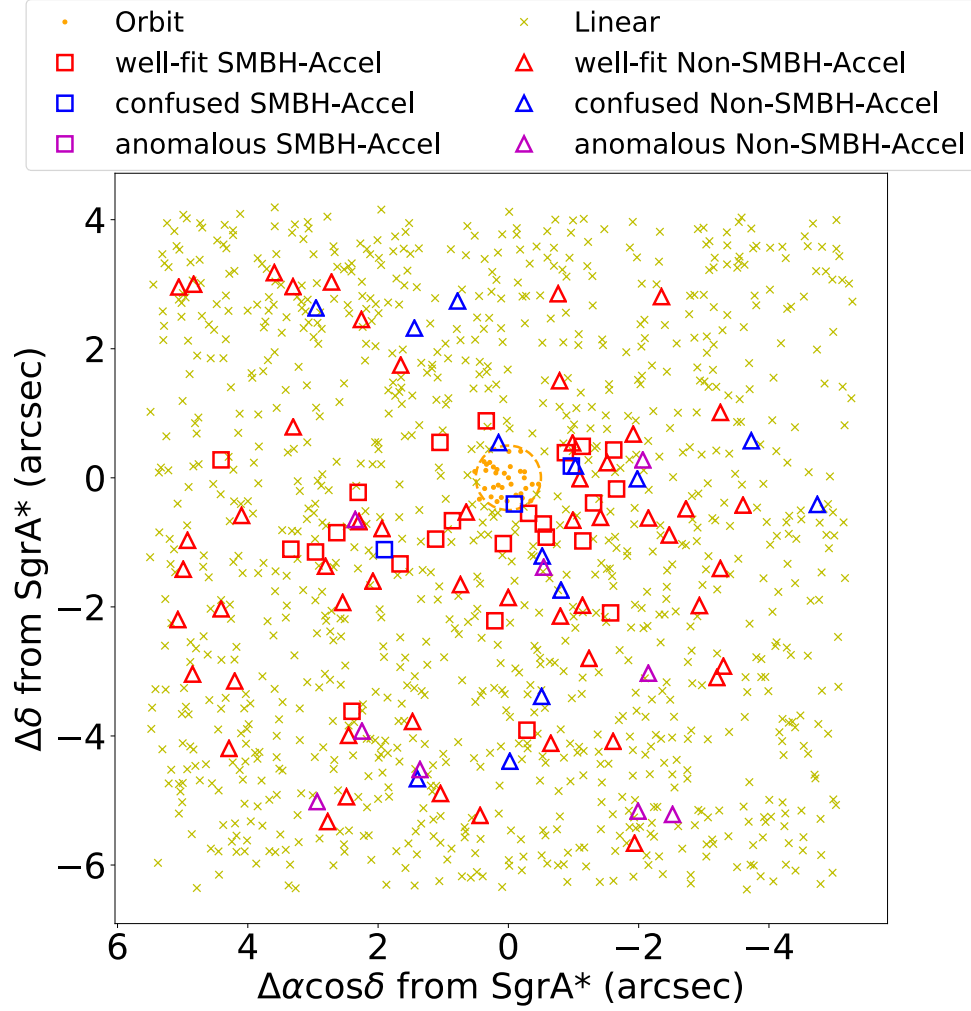


Figure 2.13: The spatial distribution of our final sample: 33 orbital stars, 1048 linear moving stars, 27 significant SMBH-acceleration sources and 76 significant non-SMBH-acceleration sources, where the significant SMBH-acceleration and non-SMBH-acceleration sources are further divided into well-fit, confused, and anomalous as described in §2.3. Stars within $0.5''$ from Sgr A* are orbital stars (*orange points*) and the orange circle shows the boundary of $0.5''$. Linearly moving stars are shown as *yellow crosses*. Stars with significant acceleration (more than 5σ) are then divided into SMBH-acceleration sources (*squares*) and non-SMBH-acceleration sources (*triangles*). These are further sub-divided into well-fit (*red*), confused (*blue*), and anomalous (*magenta*) stars.

2.4 Results

The improvements in the astrometric methodology described in §2.2 deliver astrometry that is more accurate and more precise (σ_V reduced by 40%) compared to [12]. A more detailed comparison is described in §2.4. Based on this analysis, we detect 24 high-quality SMBH-acceleration stars, which are discussed in §2.4. Finally, by fitting S0-2's orbit, we find the SMBH's position offset has been reduced by a factor of 2 in $\alpha \cos \delta$ direction and linear drift has been reduced by a factor of 4 in δ direction in §2.4.

Improved Astrometry

Stars are divided into 3 categories based on the order of the kinematic model needed to fit the on-sky positions: orbit, acceleration or linear motion (§2.3). To evaluate the aggregated goodness of the linear or acceleration fit, Figure 2.14 shows the median fitting residual as a function of time. Given the complexity of matching in the dense region around the SMBH, the orbit stars within the central 0.5'' from Sgr A* are not included. Faint stars are more likely to be confused or biased in the measurement, so we only include stars brighter than 16 magnitude. These two cuts yield a sample of 553 stars that are fit by linear motion or acceleration models based on the categories they belong to. From the Figure 2.14, we can see that the residuals from speckle epochs are 8 times larger than AO epochs, which is expected given the larger positional uncertainty in speckle epochs. In fact, Speckle and AO epochs both have residuals around 1σ relative to their positional uncertainty.

To evaluate the quality of our astrometry measurements, we also compare the χ^2 distribution from linear fits and the distribution of velocity uncertainties, σ_V , with [12] (referred as B16 here after). To make a fair comparison, we require stars to be detected in more than 20 epochs in the B16 alignment, which gives us a total of 596 stars in the comparison sample. The left panel from Figure 2.15 clearly shows the astrometry in this work is more precise (smaller σ_V) and more accurate (smaller χ^2). While the increased time baseline of our data set (up to 2013 vs. 2017) partially contributes to the increased measurement precision, the methodology changes in our work increase both the precision and accuracy. From the middle panel, it is clear that almost all stars brighter than 15 magnitude have a smaller σ_V in our analysis. The median σ_V is reduced by 40% from 0.017 mas yr⁻¹ to 0.010 mas yr⁻¹ in the $\alpha \cos \delta$ direction and from 0.019 mas yr⁻¹ to 0.010 mas yr⁻¹ in the δ direction for those bright stars. For faint stars, larger σ_V is expected due to their larger position uncertainties and will give a more robust uncertainty measurement. From the right panel, it is apparent that, for stars of all magnitudes, our analysis typically yields smaller χ^2 values. In summary, the majority of the stars favors our work relative to B16.

Figure 2.16 plots the final reduced χ^2 distribution for the *good stars* defined in §2.2. 328 stars are detected in the final sample among 352 *good stars*. With the improvements made on §2.2, the final χ^2 distribution for those 328 stars from 56 years of observation agrees with the predicted χ^2 distribution.

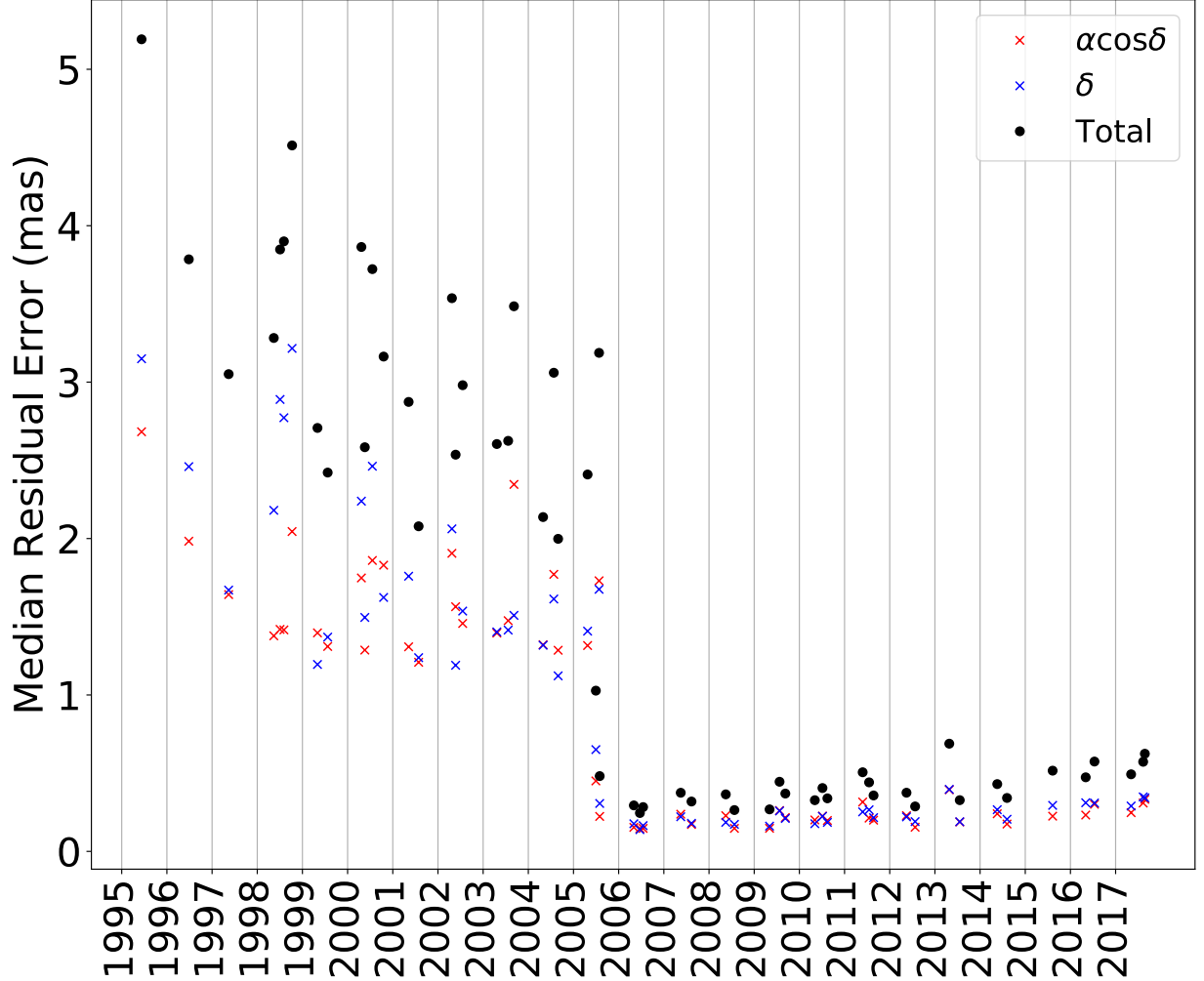


Figure 2.14: The median residual from each stars' best-fit motion model for different epochs. Here, the model is either linear or acceleration based on which category the star belongs to. The sample plotted here includes stars outside of $0.5''$, within $4''$ from Sgr A*, and brighter than 16 magnitude in K' . The red and blue crosses show the median residual in the $\alpha \cos \delta$ and δ direction separately, while the black dot is the total residual.

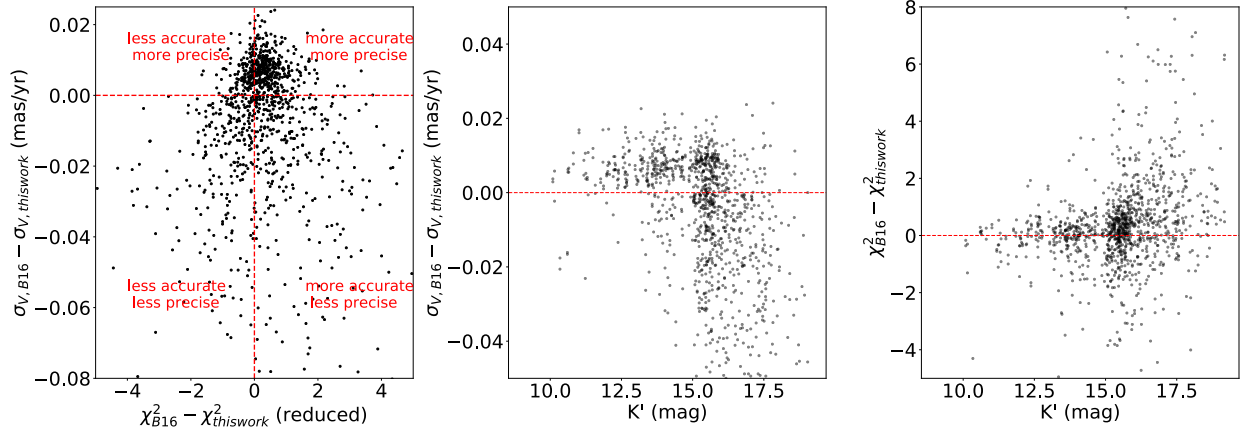


Figure 2.15: The comparison of χ^2 and σ_V for linear moving stars between B16 and this work. A smaller χ^2 indicates a more accurate measurement and a smaller σ_V indicates a more precise measurement. The *left* panel plots the Δ_{χ^2} and Δ_{σ_V} between B16 and this work. The *middle* and *right* panel show how Δ_{χ^2} and Δ_{σ_V} depend on magnitude.

New Accelerating Sources

From §2.3, we have identified 24 well-fit SMBH-acceleration sources outside of $r = 0.5''$, which are summarized in Table 2.5. According to [23], 12 (50%) of them are young stars. This is a large fraction considering that only 7% of stars are young stars among our total 1184 stars. This may suggest that young stars are very centrally concentrated compared to the old stars.

[39] published 6 young accelerating sources, 4 of which are also found accelerating in our analysis: S1-3, IRS16C, S1-14 and IRS16SW. For those 4 young accelerating sources, the average a_r uncertainty has been reduced by a factor of 2 from $0.39 \text{ kms}^{-1}\text{yr}^{-1}$ to $0.16 \text{ kms}^{-1}\text{yr}^{-1}$. The radial acceleration for S1-3, IRS16C, and IRS16SW agree within 2σ between our analysis and that of [39]. S1-14 is discrepant by 3σ , but the a_r uncertainties are very large in [39]. This large discrepancy comes from several aspects, among which the most important reasons are short time baseline and underestimated acceleration uncertainties in [39]. They only used data until 2011 (6 years AO observation), while we use data until 2017 (12 years AO observation), doubling the AO time baseline. Furthermore, the acceleration uncertainties are severely underestimated in [39] because they did not use the jackknife method as we do in §2.3. The other 2 stars that are accelerating in [39] are S0-15 and S1-12. S0-15 is also accelerating in our sample, but because of its poor acceleration fit and the presence of nearby stars (Figure 2.17), we categorize it as a confused source. S1-12 is not accelerating by more than 5σ in our analysis, so it is characterized as a linear moving star. [84] also reported the following 5 accelerating sources: S0-70, S0-36, S1-3, S1-2 and S1-13, which are all also in our sample as listed in Table 2.5.

We plot our sample of 24 accelerating sources in Figure 2.18, among which 15 accelerating

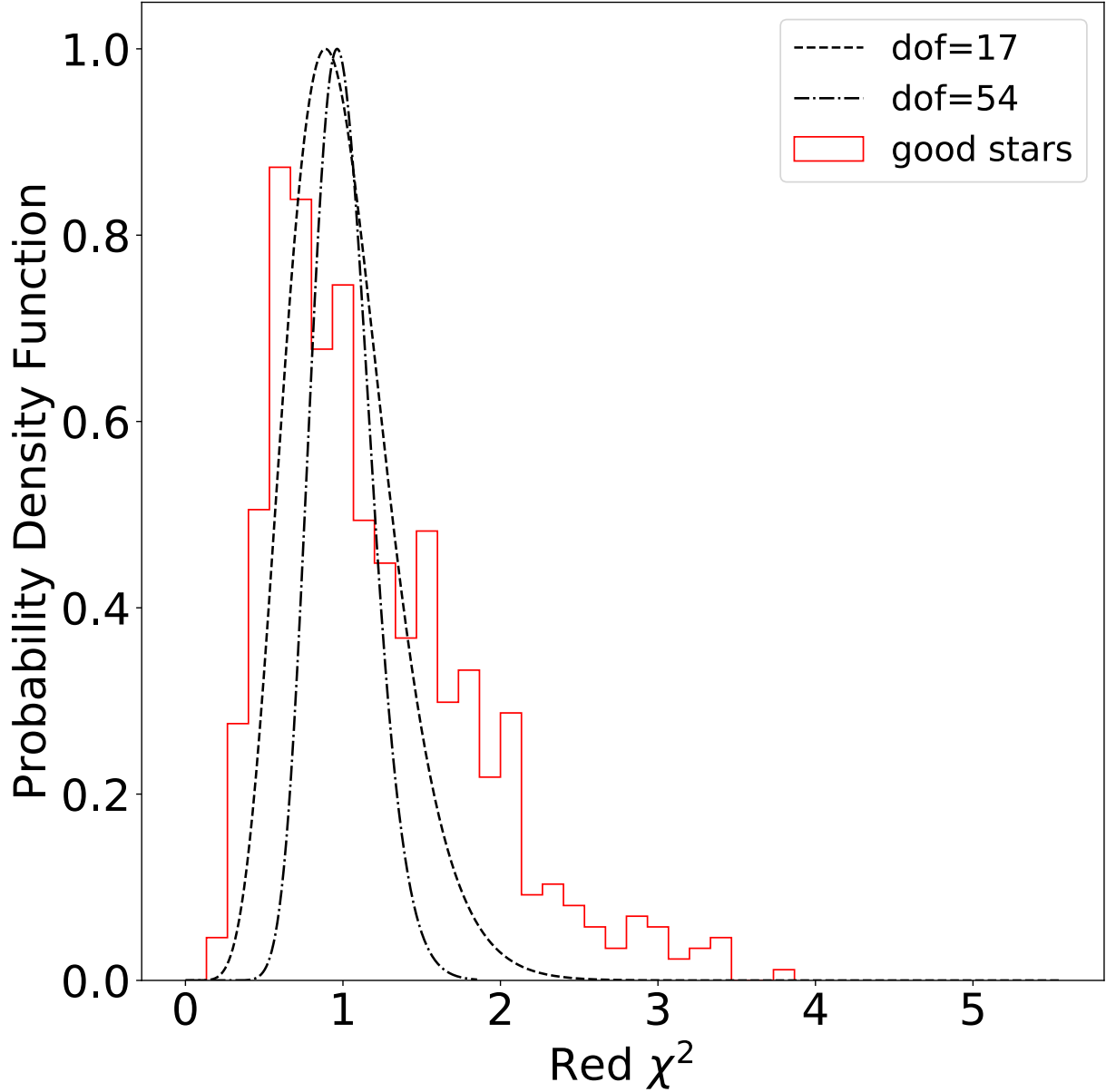


Figure 2.16: The final reduced χ^2 probability density function for *good stars*. Here reduced χ^2 comes from linear/acceleration fit based on the category defined in §2.3. Since stars are detected in different number of epochs and use different model fit (linear/acceleration), *dof* would vary between stars. To account for that effect, we plot the standard reduced χ^2 distribution for both the minimum *dof* and maximum *dof* in black lines. Reduced χ^2 from *good stars* (§2.2) is plotted in red histogram.

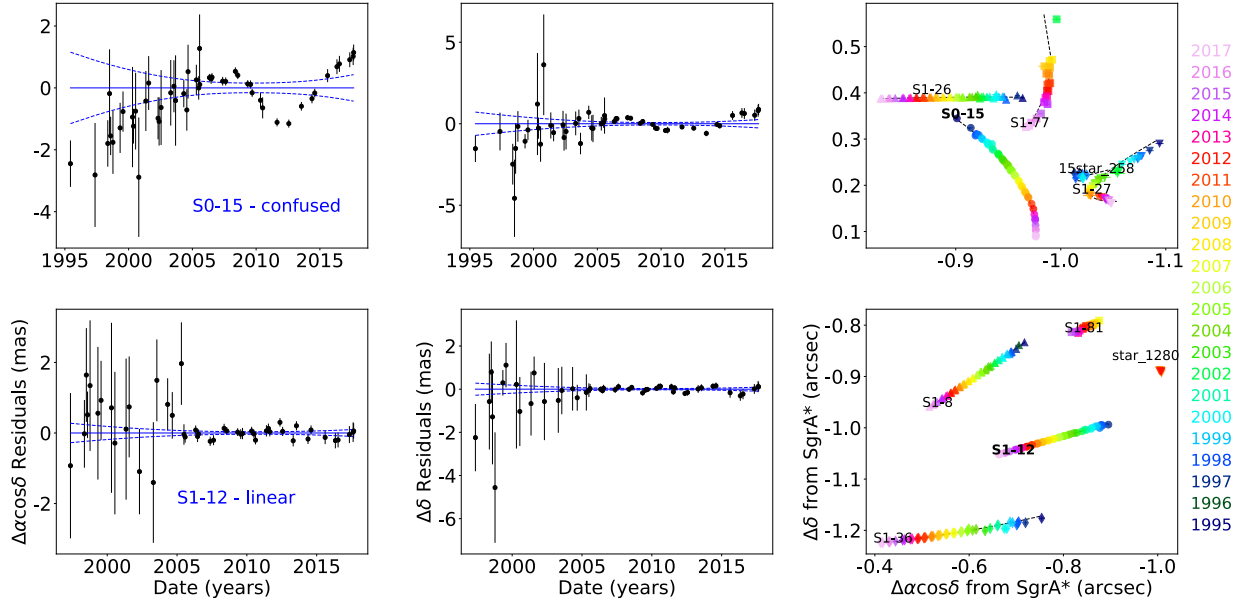


Figure 2.17: The *top* row plots the acceleration fit for star S0-15 while the *bottom* row plots the linear fit for star S1-12. The first two columns are residual from acceleration/linear fit and the third panel shows the moving track of S0-15/S1-12 and nearby stars. S0-15 and S1-12 are both accelerating sources in [39], but not in our analysis. There is a clear correlation between the residuals and time in both $\alpha\cos\delta$ and δ direction for S0-15. Potential confusion might be the reason for this poor fit, as S0-15 and S1-27 gets very close after 2007. S1-12 is not accelerating by more than 5σ in a_r and its linear fitting is already satisfying, so we label S1-12 as a linear moving star.

stars are reported for the first time. Accelerations are detected for sources at $4''$, 3 times further than previously published accelerations in [39].

Since non-SMBH-acceleration sources could exist in all directions, we determine how many potentially non-SMBH-acceleration sources exist in our 24 SMBH-acceleration sample. First, we define well-fit SMBH-acceleration sources as those where the star's tangential acceleration agrees with 0 within 3σ : $-3\sigma_{a_t} < a_t < 3\sigma_{a_t}$ and the star's radial acceleration is between 0 and $a_{r,max}$ within 3 sigma: $(a_{r,max} - 3\sigma_{a_r}) < a_r < 3\sigma_{a_r}$. In the 2D space of a_t vs. a_r , a star in the SMBH-acceleration sample would have a 1σ error ellipse that encloses $a_t = 0$ and would intersect with a line going from $a_{r,max}$ to $a_r = 0$. Since $a_{r,max}$ varies between different stars, we divided a_r and a_t by $a_{r,max}$ to normalize them.

Figure 2.19 shows $a_t/a_{r,max}$ versus $a_r/a_{r,max}$, where each stars is drawn as an ellipse with the semi-major axis of $3\sigma_{a_t}/a_{r,max}$ and $3\sigma_{a_r}/a_{r,max}$ respectively. The SMBH-acceleration stars intersect the segment on the horizontal axis from 0 to 1 and are shown in red, and the non-SMBH-acceleration stars are shown in blue. To determine the potential contaminants

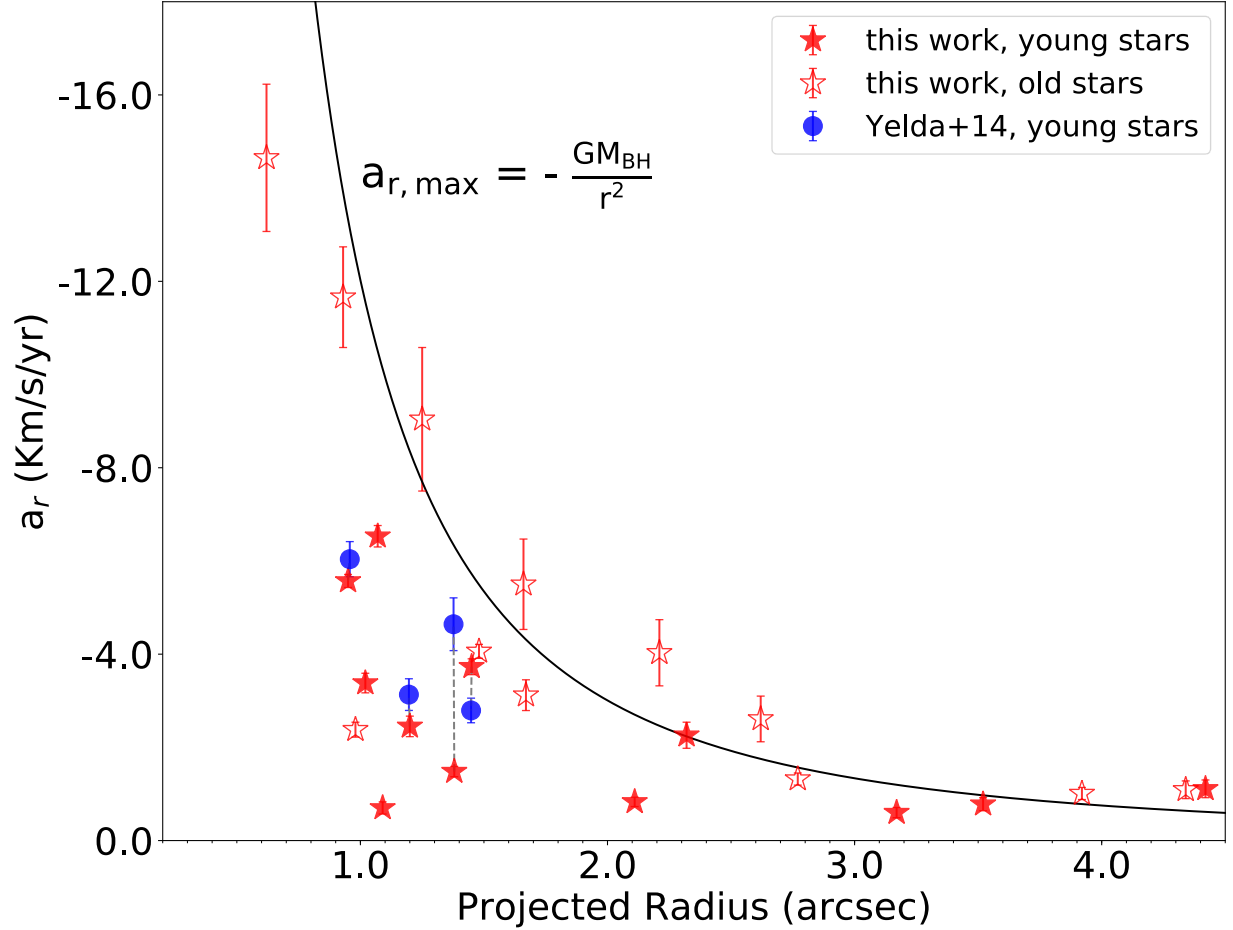


Figure 2.18: The radial acceleration a_r as a function of projected radius to Sgr A* for 24 well-fit SMBH-acceleration sources from this paper. The filled red stars are young stars and the empty red stars are old stars. Four of them (S1-3, IRS16C, S1-14 and IRS16SW) are previously published in [39], which are also plotted in blue circles. The 4 common stars are connected in dashed grey lines between our analysis and [39]. A solid black line shows the maximum allowed a_r from Sgr A* calculated by $a_{r, \max} = -GM_{\text{BH}}/r^2$, where r is the projected distance to the SMBH.

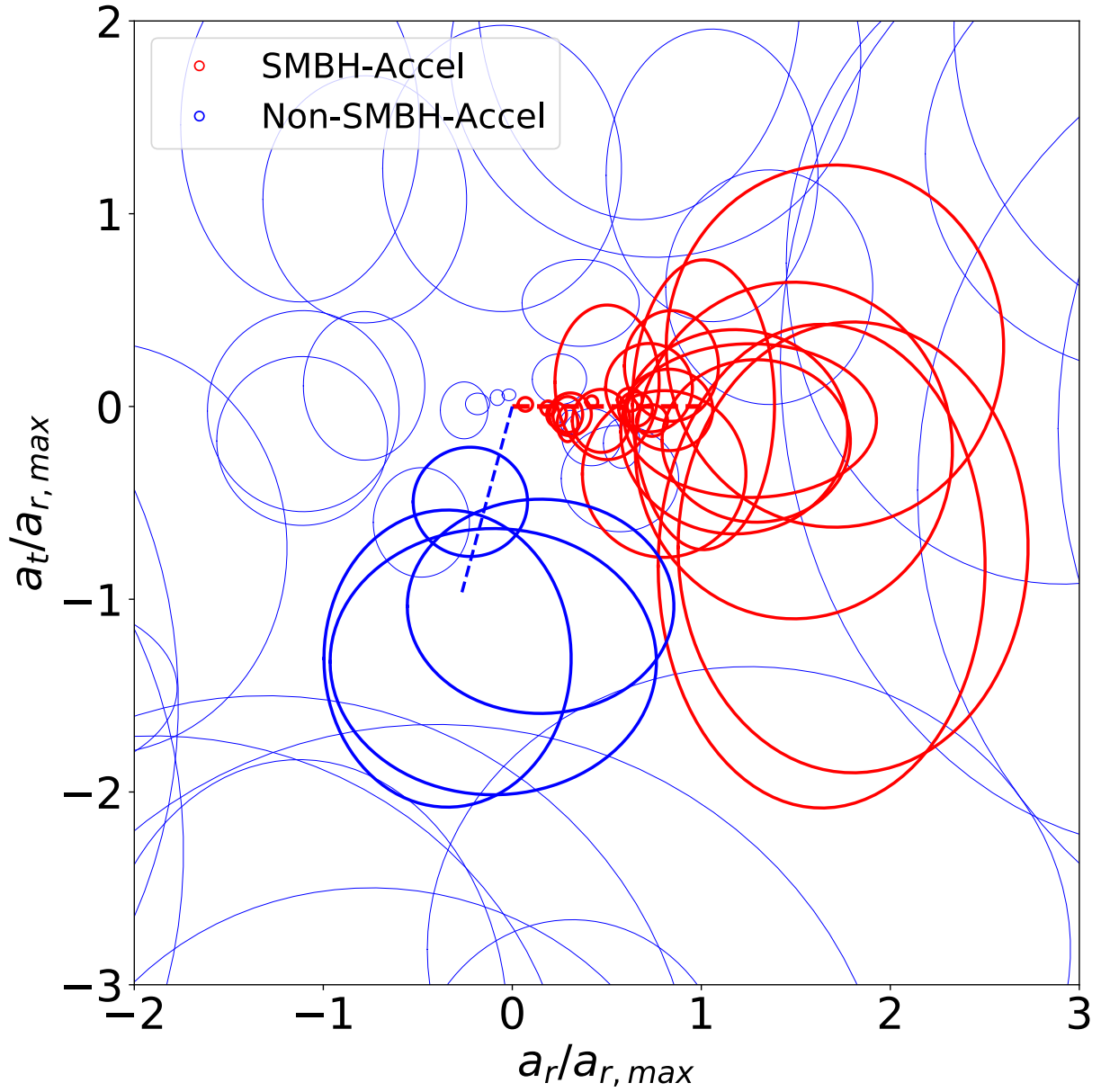


Figure 2.19: Radial (a_r) and tangential (a_t) accelerations are used to determine how the non-SMBH-acceleration sources contaminate the SMBH-acceleration sample. Each star's ellipse is centered on $[a_r/a_{r,max}, a_t/a_{r,max}]$ with the semi-major axis of $3\sigma_{a_r}/a_{r,max}$ and $3\sigma_{a_t}/a_{r,max}$. SMBH accelerations (*red ellipses*) will intersect with the red dashed line segment of length 1. To determine the non-SMBH-acceleration contaminants, we draw segments with random orientations, originating from $[0,0]$ with length of 1 and calculate how many non-SMBH-acceleration sources intersect with the segment, where non-SMBH-acceleration sources are plotted in blue ellipses. A random example is shown as a blue dashed line that intersects 5 non-SMBH-acceleration sources (thick blue ellipses).

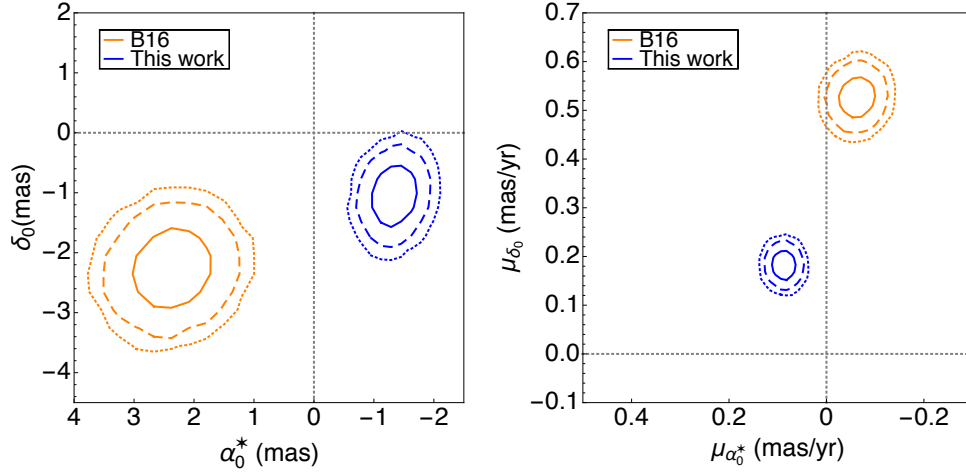


Figure 2.20: The posterior joint probability distributions of the position (α_0^*, δ_0) and proper motion $(\mu_{\alpha_0^*}, \mu_{\delta_0})$ of the SMBH based on S0-2’s orbit fit. The upper panel shows the SMBH position and the lower panel shows the SMBH’s linear drift. The orange contour is from [12], while the blue contour is from our work. The contours show 1σ , 2σ , and 3σ uncertainties. Our cross-epoch alignment reduces the SMBH offset and linear drift in both directions.

from non-SMBH-acceleration stars, we randomly draw segments with different orientations originating from $(0,0)$ and with length of 1. An example is showed in the bold blue ellipses. The average number of intersected ellipse from non-SMBH-acceleration sources gives the contamination rate. We find that among our 24 SMBH-acceleration sources, 3.5 ± 1.1 could potentially come from non-SMBH-acceleration contaminants.

To determine if the contaminants are correlated with radius, we perform similar tests for stars within $2.5''$ and beyond $2.5''$ Sgr A*. Of the 17 SMBH-acceleration sources within $2.5''$, we find that 2.8 ± 1.2 may come from non-SMBH-acceleration contaminants, or 16%. Of the 7 SMBH-acceleration sources beyond $2.5''$, we find that 0.6 ± 0.8 may come from non-SMBH-acceleration contaminants, or 9%. Therefore the contamination rate is not strongly radially dependent. Unfortunately, we cannot identify the specific contaminated sources at this time and future observations are needed.

Systematic Bias for Sgr A*

One of the goals of the methodology developed in this work is to construct a more stable reference frame for imaging observations of the Galactic Center. Currently, systematic uncertainties arising from the construction of the reference frame are assessed using orbital fits of short-period stars by including the astrometric position and velocity of the central SMBH as free parameters in the fit. A fit using S0-2’s observations in [12] has shown an offset in the position of the SMBH of 2.5 mas and a linear drift of 0.55 mas/yr. In this section,

we assess the improvements induced by our new methodology on S0-2’s orbital fit. We use astrometric observations from the cross-epoch alignment presented in §2.2. In addition, we use S0-2’s radial velocity measurements obtained using spectroscopic observations reported in [17], [9], [12], [85] and similar measurements from the VLT reported in [86], [87], [10], [11] (a summary of all S0-2’s radial velocity measurements can be found in [85]). The orbital fits are performed using Bayesian inference with the *MultiNest* sampler [88, 89]. The model used for the fit includes 13 parameters: the mass of the central SMBH, the distance to our Galactic Center R_0 , the positions (α_0^* and δ_0) and velocities ($\mu_{\alpha_0^*}$, μ_{δ_0} and v_{z_0}) of the SMBH and the six orbital parameters of S0-2. Hereafter we will use α^* to represent $\alpha \cos \delta$.

Figure 2.20 plots the posterior joint probability distributions of α_0^* and δ_0 and of $\mu_{\alpha_0^*}$ and μ_{δ_0} obtained using these observations. For comparison, we also present the posterior probability distributions obtained using S0-2 astrometry from [12]. Our new alignment methodology reduces the SMBH offset in both direction by a factor 2. The linear drift is also severely reduced in the δ -direction by a factor 4. This analysis shows that the new methodology to align the different astrometric observations presented in this work and in [79] improves the quality of the orbital fit of S0-2 significantly.

2.5 Discussion

Clockwise Disk of Young Stars

The age of the young stars (4-6 Myr) around the SMBH is much smaller than the relaxation time scale (~ 1 Gyr) in the GC, so their dynamical structure will greatly help us distinguish different star formation mechanisms [90]. Observations have found that around 20% of the young stars move in a well-defined clockwise disk [25, 16, 18, 27, 39], while the rest off-disk stars are more randomly distributed, which may suggest an in-situ formation theory. This 20% fraction might only be a lower limit because of stellar binaries [40]. An accurate disk membership derivation is required to divide disk and off-disk stars correctly, and later compare useful properties, like the initial mass function, between them.

The disk membership is derived from the standard Keplerian orbital elements, including 6 kinematic variables (α^* , δ , z , μ_{α^*} , μ_{δ} , v_z). [39] assigned the disk membership for 116 young stars using data from 1995 to 2011. With a much longer time baseline from 1995 to 2017, which doubles the AO observation time compared to [39], and an improved astrometric analysis, our work will give a much more accurate estimate of the Keplerian orbital parameters and thus improve the disk membership assignment.

Among the 6 kinematic variables, only the line of sight velocity v_z comes from the spectroscopic measurements, while the rest 5 parameters all come from the astrometric measurement. The projected position (α^* , δ) and proper motion (μ_{α^*} , μ_{δ}) can be directly derived from §2.3. For 48 linear moving stars which are reported both in [39] and our final sample, the median velocity uncertainty is reduced from $0.073 \text{ mas yr}^{-1}$ to $0.019 \text{ mas yr}^{-1}$, by almost a factor of 4. The most difficult part is to measure the line-of-sight distance (z). Fortunately,

the absolute value of line-of-sight distance can be derived using Eq. 9 from [39], if we can measure significant a_r . Even if we do not have significant measurements of a_r , stars with 3σ acceleration upper limits smaller than $a_{r,max}$ can provide lower limits on the line-of-sight distance. From §2.4, we have already detected 12 significant a_r measurements for young stars, 2 times more than [39], and our acceleration uncertainty is reduced by a factor of 2. Therefore the 5 kinematic parameters (α^* , δ , μ_{α^*} , μ_{δ} , z) can all be better estimated with our improved astrometry.

Potential Binaries

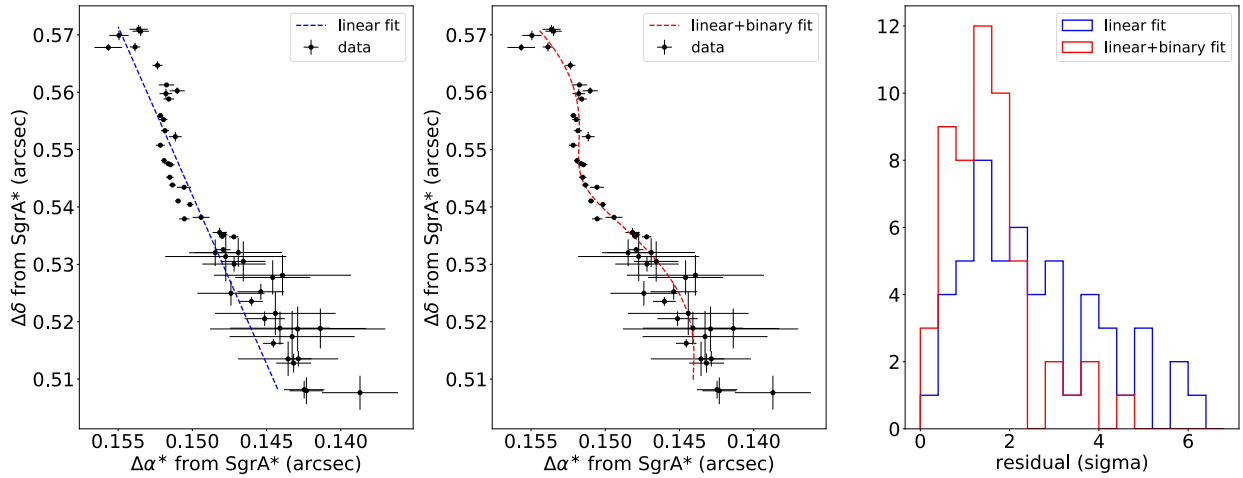


Figure 2.21: S0-27’s sky position over time with two different models overlaid. The *left* panel uses a linear motion model and the *middle* panel uses a linear+binary motion model. The fitting residuals between the model and the observations are plotted in the *right* panel in unit of sigma. The periodic feature on the proper motion makes S0-27 a potential binary candidate.

In §2.3, we identified 76 non-SMBH-acceleration sources with significant acceleration inconsistent with the gravitational force of only the SMBH. These non-SMBH-acceleration sources could be explained by binarity, microlensing, or unrecognized confusion. Here we use S0-27 as an example to explore the potential for binarity. We use two models to fit its proper motion.

Model A includes only linear motion with 4 free parameters:

- μ_{α^*} : The proper motion of the star in X direction
- μ_{δ} : The proper motion of the star in Y direction
- α_0^* : The fiducial position of the star in X direction

- δ_0 : The fiducial position of the star in Y direction.

Then, the astrometric positions α^* and δ are given by:

$$\begin{aligned}\alpha^*(t) &= \alpha_0^* + \mu_{\alpha^*}(t - t_0) \\ \delta(t) &= \delta_0 + \mu_{\delta}(t - t_0)\end{aligned}\tag{2.2}$$

Here t_0 is fiducial time, chosen to be 1990.

Model B includes linear motion plus binarity, which adds 6 parameters, with 11 free parameters in total: μ_{α^*} , μ_{δ} , α_0^* , δ_0 , ω , Ω , i , e , t_p , P , a . Here we follow the models in [91], which is summarized as follows:

- μ_{α^*} : The proper motion of the system in X direction
- μ_{δ} : The proper motion of the system in Y direction
- α_0^* : The fiducial position of the system in X direction
- δ_0 : The fiducial position of the system in Y direction
- ω : The argument of periastron of the primary star's orbit in degrees.
- Ω : The longitude of the ascending node of the secondary star's orbit in degrees.
- i : The inclination of the system in degrees.
- e : The eccentricity of the Keplerian in orbit.
- t_p : The time of periastron passage in year.
- P : The period of the Keplerian orbit in year.
- a : The photometric semi-major axis in mas.

Then astrometric positions x and y are measured by:

$$\begin{aligned}\alpha^*(t) &= \alpha_0^* + \mu_{\alpha^*}(t - t_0) + BX(t) + GY(t) \\ \delta(t) &= \delta_0 + \mu_{\delta}(t - t_0) + AX(t) + FY(t)\end{aligned}\tag{2.3}$$

where A, B, F, and G are the Thiele-Innes constants [92]; X(t) and Y(t) are the elliptical rectangular coordinates.

We use PyMultiNest [93, 88, 89] to explore the parameter space as it efficiently handles degeneracies inherent to binary orbit fitting. The priors and posteriors for the two models are presented in Table 2.6 and Table 2.7.

The best-fit models for S0-27 are shown in Figure 2.21, where the left panel is from the model A and the middle panel is from model B. Model B, with linear motion plus binarity,

is a much better fit as it reduces the Bayesian information criterion (BIC) from 944 to 395, and increases the log of the likelihood from -463 to -172.

From the linear+binary model, we are able to constrain the semi-major axis a and the period P . If we assume the primary component is a black hole, the photometric semi-major axis a would equal to the secondary component's semi-major axis. If we further assume the primary component is much more massive than the secondary component, then we can use a as the total semi-major axis of the binary system. Using Kepler's law, the total mass of this binary system can be calculated using the following equation:

$$\frac{M}{[M_{\odot}]} = \left(\frac{a}{[mas]} \frac{d}{[kpc]} \frac{1}{|\cos i|} \right)^3 \left(\frac{[yr]}{P} \right)^2 \quad (2.4)$$

With the best fit solutions in Table 2.7, the total mass of the system is $\sim 265 \pm 40 M_{\odot}$. This is significantly larger than what we expect for a binary system. However we made several simplistic assumptions that, if broken, would lower the mass. In the future, improved constraints on the total system mass will be derived by combining our astrometric measurements with multi-epoch radial velocity measurements. As astrometric monitoring of the Galactic Center continues, more candidate astrometric binaries are likely to be detected and we will be able to constrain the binary fraction at the Galactic Center.

2.6 Summary

The Galactic Center astrometric precision and accuracy has been increased by improving the cross-epoch alignment of starlists with the following major changes: (1) A magnitude dependent additive error σ_{add} is implemented for all AO epochs to create a standard χ^2 distribution. (2) A higher order local distortion map is made for 8 non-06-14 alignment data sets in 2005, 2015, 2016 and 2017. (3) Potential confusion events are removed based on stars' proper motion. (4) Artifact edge sources coming from elongated PSF wings are excluded from our final sample. (5) We use jackknife to derive robust proper motion uncertainties.

These new astrometric methods produce both more precise and more accurate stellar proper motions (σ_V reduced by 40%) as compared with our previous work [e.g. 12]. Among the final sample of 1184 stars, we have identified 24 significantly accelerating sources with 3.5 potential contaminants, among them 15 are reported for the first time. We have constructed a much more stable reference frame - the position and velocity of Sgr A* derived from S0-2's orbit is both more precise and more accurate, by more than a factor of 2.

This improved astrometry will help answer many open questions in the GC. For example, with a better measurement of proper motion, especially significant acceleration, the young stars can be classified as disk and off-disk stars more easily, which will help with understanding the star formation history in the GC. Tests of General Relativity with S0-2 will be significantly improved. With longer time baseline, we will even be able to find potential binaries and microlensing candidates based on their astrometric measurements.

Table 2.2: Summary of AO Imaging Observations

Date		Frames		FWHM	Strehl	N _{stars}	K _{lim} ^a	σ_{pos} ^b	Data Source ^c
(U.T.)	(Decimal)	Obtained	Used	(mas)	ratio		(mag)	(mas)	
2005 Jun 30	2005.495	10	10	62	0.29	794	16.2	0.39	LGSAO; Ref. 8
2005 Jul 31	2005.580	59	31	57	0.22	1753	19.0	0.25	LGSAO; Ref. 7
2006 May 3	2006.336	127	107	58	0.32	1951	19.2	0.05	LGSAO; Ref. 7
2006 Jun 20-21	2006.470	289	156	57	0.35	2438	19.5	0.08	LGSAO; Ref. 7
2006 Jul 17	2006.541	70	64	58	0.34	2165	19.3	0.09	LGSAO; Ref. 7
2007 May 17	2007.374	101	76	58	0.35	2492	19.5	0.09	LGSAO; Ref. 7
2007 Aug 10,12	2007.612	139	78	58	0.30	1877	19.1	0.08	LGSAO; Ref. 7
2008 May 15	2008.371	138	134	54	0.29	2080	19.4	0.06	LGSAO; Ref. 9
2008 Jul 24	2008.562	179	104	58	0.32	2175	19.3	0.04	LGSAO; Ref. 9
2009 May 1,2,4	2009.340	311	149	57	0.35	2297	19.4	0.04	LGSAO; Ref. 9
2009 Jul 24	2009.561	146	75	62	0.25	1699	18.9	0.09	LGSAO; Ref. 9
2009 Sep 9	2009.689	55	43	62	0.31	1920	19.1	0.11	LGSAO; Ref. 9
2010 May 4-5	2010.342	219	158	63	0.28	2027	19.2	0.06	LGSAO; Ref. 9
2010 Jul 6	2010.511	136	117	62	0.29	1950	19.1	0.08	LGSAO; Ref. 9
2010 Aug 15	2010.620	143	127	61	0.26	1819	19.1	0.07	LGSAO; Ref. 9
2011 May 27	2011.401	164	114	66	0.25	1557	18.9	0.13	LGSAO; Ref. 9
2011 Jul 18	2011.543	212	167	59	0.26	2017	19.3	0.07	NGSAO; Ref. 9
2011 Aug 23-24	2011.642	218	196	60	0.32	2354	19.5	0.05	LGSAO; Ref. 9
2012 May 15,18	2012.371	290	201	59	0.30	2256	19.4	0.05	LGSAO; Ref. 10
2012 Jul 24	2012.562	223	162	58	0.33	2317	19.5	0.06	LGSAO; Ref. 11
2013 Apr 26-27	2013.318	267	140	68	0.22	1264	18.4	0.09	LGSAO; Ref. 11
2013 Jul 20	2013.550	238	193	58	0.33	1788	19.1	0.08	LGSAO; Ref. 11
2014 May 19	2014.380	173	147	64	0.28	1468	18.7	0.08	LGSAO; Ref. 12
2014 Aug 6	2014.596	137	127	56	0.33	1760	19.1	0.08	LGSAO; Ref. 12
2015 Aug 9-11	2015.606	288	203	58	0.33	1887	19.1	0.06	LGSAO; Ref. 12
2016 May 3	2016.338	253	166	60	0.31	1655	18.9	0.06	LGSAO; Ref. 12
2016 Jul 13	2016.532	207	144	60	0.26	1378	18.5	0.07	LGSAO; Ref. 12
2017 May 4,5	2017.343	469	179	59	0.31	1674	19.0	0.06	LGSAO; Ref. 12
2017 Aug 9-11	2017.610	213	111	55	0.32	1476	18.7	0.09	LGSAO; Ref. 12
2017 Aug 23,24,26	2017.647	216	112	61	0.29	1216	18.0	0.06	LGSAO; Ref. 12

^a K_{lim} is the magnitude at which the cumulative distribution function of the observed K magnitudes reaches 90% of the total sample size;

^b Positional error taken as error on the mean from the three sub-images in each epoch and includes stars ;

^c Data originally reported in (6) Ghez et al. (2005)[67], (7) Ghez et al. (2008)[9], (8) Lu et al. (2009)[27], (9) Yelda et al. (2014)[39], (10) Meyer et al. (2012) [68], (11) Boehle et al. (2016)[12], and (12) this work;

Table 2.3: Local Distortion Summary

Date	Distortion in X (pixel)	Distortion in Y (pixel)
2005 Jun 30	0.028 ± 0.025	0.050 ± 0.039
2005 Jul 31	0.029 ± 0.016	0.051 ± 0.021
2015 Aug 10	0.033 ± 0.017	0.030 ± 0.022
2016 May 03	0.038 ± 0.021	0.031 ± 0.025
2016 Jul 13	0.047 ± 0.019	0.034 ± 0.025
2017 May 05	0.025 ± 0.020	0.030 ± 0.024
2017 Aug 11	0.028 ± 0.023	0.030 ± 0.026
2017 Aug 24	0.025 ± 0.022	0.027 ± 0.025

Both the distortion value and distortion uncertainty value reported here is the median value without edges. Edges are defined when distortion or distortion uncertainty is larger than 0.3 pixel, which comes from the lack of sample stars on the edge.

Table 2.4: Accelerating Categories Summary

Category	SMBH-Accel	Non-SMBH-Accel
well-fit	24	54
confused	3	13
anomalous	0	9

Table 2.5: Significant SMBH-acceleration Sources

Name	Mag	Radius	a_r	a_t	a_r from [39]	a_r from [84]	Young ^a
		(arcsec)	(km s ⁻¹ yr ⁻¹)	(km s ⁻¹ yr ⁻¹)	(km s ⁻¹ yr ⁻¹)	(km s ⁻¹ yr ⁻¹)	
S0-70	17.8	0.62	-14.65 ± 1.58	2.33 ± 1.70	—	-22.64 ± 3.02	—
S0-36	15.9	0.93	-11.66 ± 1.08	0.25 ± 0.99	—	-13.21 ± 2.26	—
S1-3	12.0	0.95	-5.57 ± 0.14	-0.31 ± 0.14	-6.04 ± 0.38	-2.64 ± 0.38	y
S1-26	15.4	0.98	-2.38 ± 0.16	0.12 ± 0.16	—	—	—
S1-2	14.6	1.02	-3.38 ± 0.21	1.03 ± 0.36	—	-1.89 ± 0.38	y
S1-4	12.3	1.07	-6.53 ± 0.23	-0.37 ± 0.21	—	—	y
S1-8	14.0	1.09	-0.70 ± 0.13	-0.11 ± 0.12	—	—	y
IRS16C	9.9	1.20	-2.45 ± 0.22	0.35 ± 0.26	-3.13 ± 0.34	—	y
S1-92	16.6	1.25	-9.04 ± 1.54	1.01 ± 1.36	—	—	—
S1-14	12.6	1.38	-1.48 ± 0.11	0.27 ± 0.13	-4.64 ± 0.57	—	y
IRS16SW	10.0	1.45	-3.73 ± 0.17	0.08 ± 0.17	-2.79 ± 0.26	—	y
S1-13	13.9	1.48	-4.06 ± 0.15	0.40 ± 0.15	—	-2.64 ± 0.38	—
S1-47	15.5	1.66	-5.50 ± 0.97	0.32 ± 0.58	—	—	—
S1-51	14.9	1.67	-3.12 ± 0.33	-0.42 ± 0.33	—	—	—
S2-6	11.8	2.11	-0.83 ± 0.10	0.11 ± 0.10	—	—	y
S2-127	15.7	2.21	-4.03 ± 0.71	2.04 ± 1.03	—	—	—
S2-22	12.8	2.32	-2.26 ± 0.28	-0.02 ± 0.56	—	—	y
S2-219	15.8	2.62	-2.61 ± 0.49	0.40 ± 0.51	—	—	—
S2-75	14.3	2.77	-1.32 ± 0.13	-0.33 ± 0.15	—	—	—
S3-5	16.1	3.17	-0.60 ± 0.11	-0.15 ± 0.16	—	—	y
S3-10	13.6	3.52	-0.78 ± 0.14	0.34 ± 0.14	—	—	y
S3-370	13.5	3.92	-1.01 ± 0.13	0.14 ± 0.11	—	—	—
S4-139	14.4	4.34	-1.09 ± 0.19	-0.20 ± 0.20	—	—	—
S4-169	13.5	4.42	-1.11 ± 0.19	0.45 ± 0.24	—	—	y

^a Young stars are published in [23];

Table 2.6: Linear Model

Params	Prior	Best Fit
μ_{α^*} (mas/yr)	flat prior in [0,1]	0.56 ± 0.01
μ_{δ} (mas/yr)	flat prior in [2,4]	3.28 ± 0.01
α_0^* (")	flat prior in [0.12, 0.15]	0.1396 ± 0.0002
δ_0 (")	flat prior in [0.45, 0.50]	0.4808 ± 0.0003

Table 2.7: Linear + Binary Model

Params	Prior	Best Fit
μ_{α^*} (mas/yr)	flat prior in [0,1]	0.61 ± 0.01
μ_{δ} (mas/yr)	flat prior in [2,4]	3.15 ± 0.02
α_0^* (")	flat prior in [0.12, 0.15]	0.1388 ± 0.0003
δ_0 (")	flat prior in [0.45, 0.50]	0.4834 ± 0.0004
ω (degree)	flat prior in [0, 360]	42 ± 14
		$222 \pm 16^*$
Ω (degree)	flat prior in [0, 360]	136 ± 3
		$316 \pm 3^*$
i (degree)	flat prior in $P(i) = \sin(i)$ in [0,180]	111 ± 3
e	flat prior in $P(e) = e$ in [0,1]	0.40 ± 0.08
t_p (year)	flat prior in [2000, 2020]	2010.2 ± 0.4
P (year)	flat prior in [0, 30]	12.7 ± 0.6
a (mas)	flat prior in [0, 4]	1.55 ± 0.08

* ω and Ω have two solutions because they are degenerate with each other;

Chapter 3

Stellar Populations in the Central 0.5 pc of Our Galaxy: The Dynamical Sub-structures

3.1 Observations and Data Reduction

The kinematic analysis of the young stars at the Galactic Center requires both proper motion and radial velocity (RV) measurements in order to determine their orbital planes and membership on the disk. Details of the observations, data reduction, and image analysis are presented in [94] and [23]. Here, we briefly summarize the analysis methods most relevant to this work.

The photometry for those young stars are from deep wide mosaic image [63]. We applied the latest extinction map from [95] to correct extinction of $A_{ks}=2.7$.

Sample Selection

In this work, we included all spectroscopically identified young (early-type) stars with well measured radial velocities (RV) and proper motions. To get a complete young stars list, we combined our new Galactic Center OSIRIS Wide-field Survey (GCOWS) with previous GCOWS observations and with other spectral types from the literature. Our GCOWS survey sample includes stars that are spectroscopically identified as early-type from the GCOWS survey and that have sufficient Signal-to-Noise ratio to measure their radial velocities (RV). Details of the GCOWS survey are presented in [23] and [19], where they published young stars covering the central region and eastern field in the GC (green boxes in Figure 3.1). In this work, we have added new observations in the South and North (magenta boxes in Figure 3.1). This resulted in 7 more young stars with good quality RVs - S10-261, S10-48, S11-176, S11-21, S11-246, S12-76, S5- 106 (see Section 3.1 for observation details).

Then, we combined our list of young stars with those identified in the literature from [16],

[18] and [20]. S2-66 is claimed to be young in [16], but later proven to be old in [23], so this star was excluded. By combining those detections, we are able to collect RVs for 149 young stars in Section 3.1. Unfortunately, [16] and [18] did not report their spectral completeness curve, so we can not use stars that are only found in their paper, but we can still use their RVs for stars identified in GCOWS or [20], which leaves us with 91 stars with both RVs and their completeness correction curves.

Among those stars, we are able to extract 88 stars' proper motions from a combination of Keck AO observations in the inner region and HST observations in the outer region (see Section 3.1 for details). The spatial distribution of our young star sample is plotted in Figure 3.1. Although seven new young stars were identified from the new GCOWS observations, S10-261 does not have a measured proper motion, so only six new stars are included in our sample, as shown in Figure 3.1. In summary, we have 88 young stars in our sample, extending out to $14''$ (~ 0.5 pc), down to a 90% limiting magnitude of $K_{\text{lim}} = 15.3$.

Radial Velocities

As mentioned in Section 3.1, RVs used in this work come from two sources: (1) Our GCOWS survey from Keck observations [23, 19]. (2) Other published RV data for Galactic Center, including [16], [18] and [20].

The GCOWS survey consists of observations with the Keck OSIRIS spectrograph behind laser-guide star adaptive optics. We obtain diffraction-limited medium spectral-resolution ($R \sim 4000$) spectra with the Kn3 filter (2.121-2.220 μm). We use two different plate scales: 35 mas in the central fields where the stellar densities are the highest and 50 mas for the outer fields with relatively lower stellar densities. Additional details of the GCOWS survey are presented in [23] and [19]. Overall, our GCOWS survey sample includes 50 stars that are spectroscopically identified as early-type from the GCOWS survey and that have sufficient Signal-to-Noise ratio to measure their RVs.

We derive radial velocities for all Keck OSIRIS data (both previously reported and new) using full spectral fitting with a synthetic spectral grid. We use the spectral fitting code StarKit [96] to fit the radial velocity along with their physical properties such as the effective temperature, surface gravity, metallicity, and rotational velocity. By fitting the physical parameters simultaneously, we can capture the effect of correlations between the parameters. We use the BOSZ spectral grid [97] to generate the spectra for our Bayesian inference model. Additional discussion of this method is given in [98, 4]. In general, the statistical uncertainties dominate the radial velocity measurement, but for the brightest sources, the systematic uncertainties dominate at the level of 11 km s^{-1} , which is mainly due to the residual OH line subtraction. See [4] for a complete discussion of radial velocity systematic uncertainties.

Both [16] and [18] used AO-assisted, near-infrared integral field spectrometer SPIFFI / SINFONI on ESO VLT. Since [18] is claimed to be an improvement from [16], we will always adopt the RVs from [18] if it is reported differently between them. [20] used integral-field spectrograph KMOS on VLT.

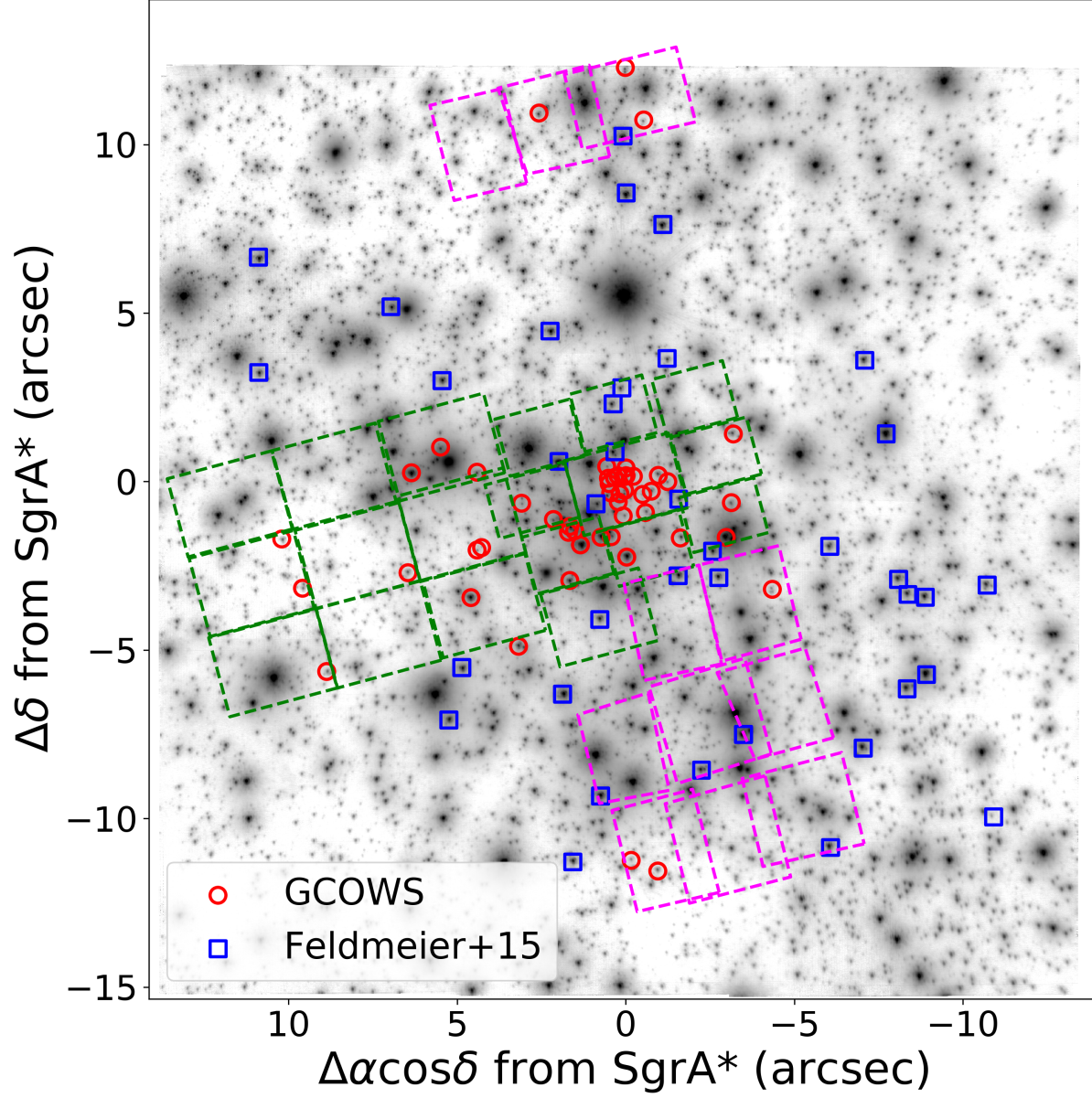


Figure 3.1: The spatial distribution of our final sample of 88 stars. Red circles are the 50 stars from our GCOWS observations and blue squares are the 38 stars from [20]. The dashed lines show our GCOWS sky coverage, where the green boxes are previously published fields in [19] and magenta boxes are new areas first published in this paper.

Based on the magnitudes and positions of stars, we match the star lists reported in other literature with our high resolution image. However, due to different spatial resolution between our observation and other published observation, not all stars can be matched. For example, star 3308 from [20] is matched to a clump of 3 stars in our image, and it is hard to determine which one they reported. All RVs are reported in Table 3.1 and 3.2.

For stars detected in GCOWS, we always use our measurements. Most stars only have one detection, which will be used as their final RV measurements. Some stars in the central region have multiple detections, which are marked with asterisk mark in Table 3.1. For those multiply detected stars, we will use the weighted mean RV if it is detected less than 5 times or has a relatively constant RV. However if RV measurements show significant variations with time (S0-1, S0-2, S0-3, S0-8, S0-15, S0-19, S1-2), they will be fitted with a full Keplerian orbit in Section 3.2.

For stars that are not detected in our database, we will use their literature RVs. For stars reported multiple times in the literature, we use the weighted mean RV, where the weight, w computed as:

$$w = \frac{1}{\sigma_{\text{RV}}^2} \quad (3.1)$$

All stars with RVs from literature agree with each other within 2 sigma, except star S7-236, for which we adopt the most recent RV from [20].

Spectroscopic Completeness

In order to properly correct for incompleteness in our selection sample, we utilize results from star-planting simulations. The sample was selected from two sources: [20] and GCOWS, and their completeness is described in the following context.

For stars in GCOWS observation, the completeness C is a production of imaging completeness C_{img} and spectral completeness C_{spec} . Imaging completeness C_{img} is estimated using star-planting simulations (see details in Appendix C.1 in [19]), and it is 90% complete down to $K_p=16$, as shown in Figure 3.2.

For spectral completeness C_{spec} , we follow a similar process as in [19], where each star from GCOWS survey is assigned a probability of being young P_E . So for young stars, $P_E = 1$; for old stars, $P_E = 0$; for unknown type stars, P_E is simulated based on the Bayesian evidence for the early-type and late-type hypotheses using the known type stars as training sample. However, not all young stars have spectra with good enough quality to measure their RVs. Therefore, the completeness used in this work is defined as "completeness for young stars with measured RVs".

Stars are divided into 8 magnitude bins based on $K_{p,\text{ext}}$ from 9 to 17 with 1 magnitude interval, and the spectral completeness curve is an linear interpolation. In each magnitude bin, the completeness C_{spec} is calculated using the following equation:

$$C_{\text{spec}} = \frac{N_{\text{youngRV}} + N_{\text{youngWR}}}{N_{\text{young}} + \sum_{\text{unk}} P_E} \quad (3.2)$$

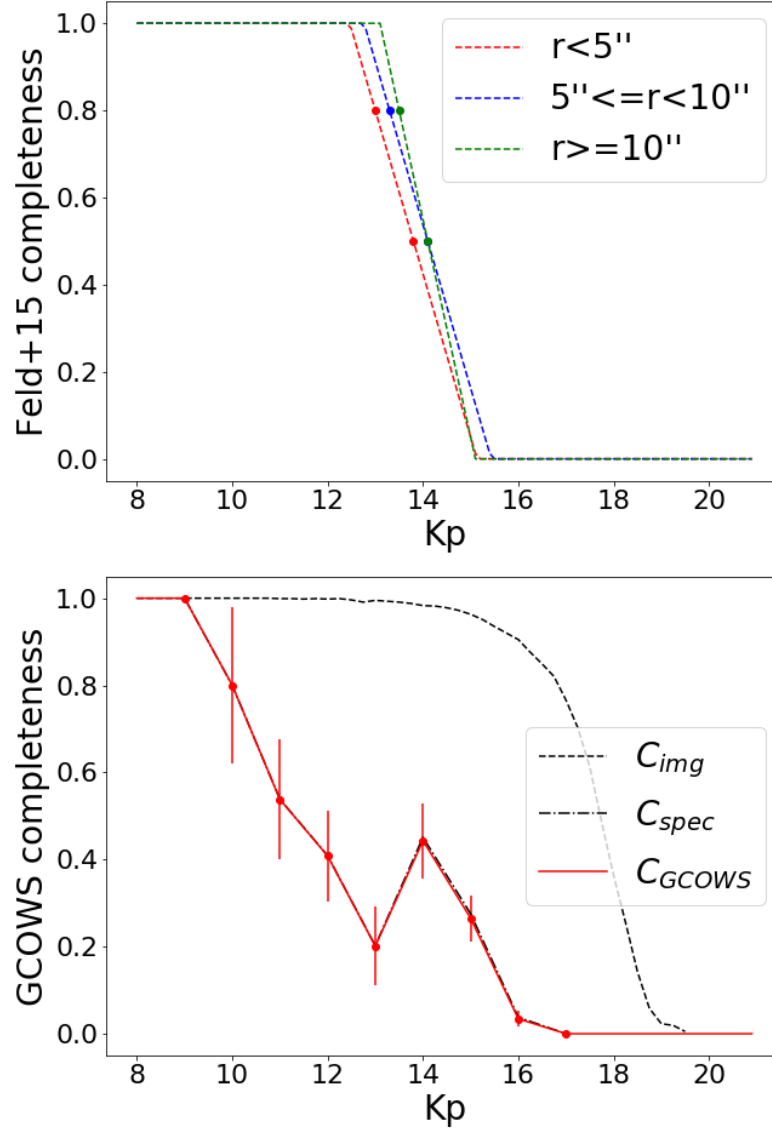


Figure 3.2: Completeness curve as a function of magnitude K_p for our data (top) and [20] (bottom). Completeness for [20] is a linear interpolation from the 80% completeness and 50% completeness point they reported in their paper. And completeness also depends on radius, where inner region is less complete because of crowdedness. Completeness from our observation in red line is a product of imaging completeness and spectral completeness, but mostly determined by spectral completeness. The dip in $K_p = 13$ is probably a result of the fact that stars are transitioning to main sequence at that stage, so most spectra we get in that magnitude bin are featureless without usable RV.

Where N_{youngRV} is the number of young stars with well-measured RV, N_{youngWR} is the number of WR young stars, N_{young} is the total number of young stars (including all spectrally identified young stars, no matter whether they have well-measured RV or not) and $\sum_{\text{unk}} P_E$ is the sum of probability of being young for all unknown type stars. WR stars are all very bright and have high SNR spectra, but currently we could not fit their emission lines due to the lack of a good model. So we decided to include WR stars in the numerator, since the missing RV from them are not because of incompleteness. The total completeness for young stars with RVs is shown in red line in Figure 3.2. Usually completeness will decrease towards fainter magnitude, but a dip in completeness curve appears at $Kp = 13$. This is because young stars are at pre-main sequence turn off point at this magnitude, so they are obscured by dust and gas, making them harder to see.

For stars from [20], they reported 80% and 50% completeness at different radial bins. An linear interpolation is derived based on those two data point. Notice that they report their completeness is extinction corrected magnitude Kp_{ext} , but we assume it is a proper approximation to completeness curve in the observed magnitude system. The completeness curve for [20] is shown in Figure 3.2.

When determining which completeness to use for each star, first we need to determine whether this star is within our GCOWS field. If a star is within our GCOWS field, we will use the higher completeness fraction between GCOWS and [20]. Otherwise, completeness from [20] will be applied. An example of completeness map at $K_p = 15$ is shown in Figure 3.3.

Position, Proper motion and Acceleration

The projected positions and proper motions are derived either from our long time baseline observation with the 10 m W. M. Keck Observatory or HST. We have been taking images of the GC at different resolution and field of view (FOV) to achieve specific scientific goals.

(1) The central $10'' \times 10''$ region of the GC (approximately centered around Sgr A*) has been monitored with diffraction-limited, near-infrared imaging cameras from Keck observatory since 1995. For those stars, we have the longest time baseline and highest resolution, which gives precise proper motions and even significant accelerations for well measured stars. The detailed analysis is presented in [94].

(2) To measure the proper motions of the young stars at larger radii, we used a widely dithered mosaic images with a $22'' \times 22''$ FOV as described in [79]. The astrometric uncertainties for mosaic data are larger in general than the central $10''$ data, because of its shorter time baseline and shallower detection limit.

(3) For stars at even larger radius ($R > 7.5''$), we use proper motions measured from the HST WFC3-IR observations. This data set contains 10 epochs of F153M observations centered on SgrA* that were obtained between 2010 – 2020 (2010.5: GO 11671/PI Ghez, 2011.6: GO 11671/PI Ghez, 2012.6: GO 12318/PI Ghez, 2014.1: GO 13049/PI Do, 2018.1: GO 15199, PI Do, 2019.2: GO 15498/PI Do, 2019.6: GO 16004/PI Do, 2019.7: GO 16004/PI Dom 2019.8: GO 16004/PI Do, 2020.2: GO 15894/PI Do). While the HST resolution is ~ 2.5

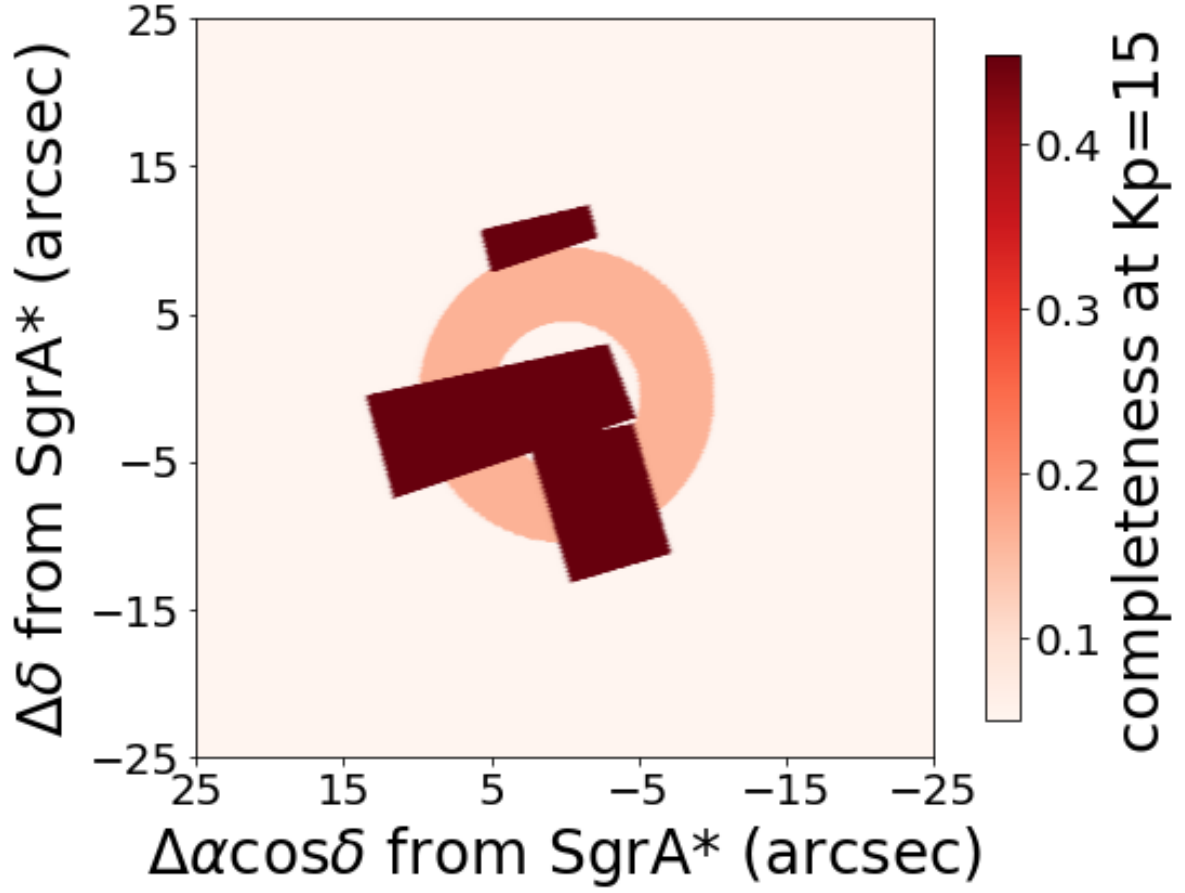


Figure 3.3: An example of completeness map at $k_p=15$ with color showing the completeness. For stars at this magnitude, completeness is higher in our curve, so GCOWS completeness is adopted for observed GCOWS regions (darker polygons) while completeness from [20] is adopted elsewhere (three concentric circles).

times lower than is achieved with the Keck observations (FWHM $\sim 0.17''$ versus FWHM $\sim 0.06''$), its FOV of $120'' \times 120''$ is much larger than can be realistically achieved with current AO systems. The astrometry from each epoch is first transformed into the *Gaia* absolute reference frame [99] and then further transformed into the AO reference frame via 2nd order polynomial transformations. The resulting HST catalog achieves an average precision of 0.33 mas and 0.07 mas/yr for the positions and proper motions of the stars in the sample. A detailed description of the HST catalog will be provided in a future paper (Hosek et al., in prep).

In summary, among 91 young stars from sample on Section 3.1, we are able to find proper motion for 88 stars. The proper motion for the final sample of 54 stars are from category (1), 20 are from category (2), 14 are from category (3), as showed in Table 3.4.

Table 3.1: RVs Summary

Name	RV _{GCOWS} ^a (km s ⁻¹)	t0 _{GCOWS}	ID _{P06}	RV _{P06} ^b (km s ⁻¹)	RV _{B09} ^c (km s ⁻¹)	ID _{F15} (km s ⁻¹)	RV _{F15} ^d (km s ⁻¹)
S0-20	261 ± 25*	2008.4	E3	-280 ± 50	—	—	—
S0-2	-473 ± 27*	2002.4	E1	-1060 ± 25	—	—	—
S0-1	-1054 ± 19*	2005.5	E4	-1033 ± 25	—	—	—
S0-8	-389 ± 40*	2006.5	E7	-390 ± 70	—	—	—
S0-16	139 ± 28*	2011.5	E2	300 ± 80	—	—	—
S0-3	-709 ± 15*	2006.5	E6	-570 ± 40	—	—	—
S0-5	631 ± 14*	2005.5	E9	610 ± 40	—	—	—
S0-19	164 ± 46*	2006.5	E5	280 ± 50	—	—	—
S0-11	-59 ± 29*	2006.5	E12	-20 ± 150	—	—	—
S0-7	72 ± 72*	2006.5	E11	160 ± 60	—	—	—
S0-4	-2 ± 19*	2006.5	E10	15 ± 30	—	—	—
S0-30	-16 ± 71	2008.4	—	—	—	—	—
S0-9	135 ± 20*	2005.5	—	—	—	—	—
S0-31	-132 ± 30*	2006.5	E13	-890 ± 31	—	—	—
S0-14	-82 ± 21*	2006.5	E14	-14 ± 40	-75 ± 25	2233	-28 ± 104
S1-3	—	—	E15	68 ± 40	1 ± 33	562	110 ± 72
S0-15	-738 ± 33*	2006.5	E16	-424 ± 70	-557 ± 26	—	—
S1-2	39 ± 32*	2006.5	E17	26 ± 30	—	—	—
S1-8	57 ± 48*	2006.5	E18	-364 ± 40	—	1619	102 ± 94
S1-4	—	—	—	—	—	668	-221 ± 179
S1-33	44 ± 23*	2006.5	—	—	—	—	—
S1-22	—	—	E25	-224 ± 50	-235 ± 100	900	-229 ± 35
S1-19	-169 ± 16	2007.5	—	—	—	—	—
S1-24	125 ± 10	2008.4	E26	206 ± 30	206 ± 30	331	221 ± 45

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Table 3.1 – continued from previous page

Name	RV _{GCOWS} ^a (km s ⁻¹)	t0 _{GCOWS}	ID _{P06}	RV _{P06} ^b (km s ⁻¹)	RV _{B09} ^c (km s ⁻¹)	ID _{F15}	RV _{F15} ^d (km s ⁻¹)
IRS 16CC	–	–	E27	241 ± 25	241 ± 25	64	256 ± 12
S2-4	218 ± 6	2008.4	E28	286 ± 20	286 ± 20	443	229 ± 36
S2-6	152 ± 8	2008.4	E30	216 ± 20	216 ± 20	–	–
IRS 33N	23 ± 5	2007.5	E33	68 ± 20	63 ± 20	294	105 ± 61
S2-50	-135 ± 32	2008.4	–	–	–	–	–
S2-17	57 ± 4	2008.4	E34	100 ± 20	100 ± 20	109	149 ± 27
S2-21	-92 ± 12	2007.5	–	–	–	1534	-83 ± 42
S2-19	–	–	E36	41 ± 20	41 ± 20	941	157 ± 55
S2-58	77 ± 16	2007.5	–	–	–	–	–
S2-74	–	–	E38	36 ± 20	36 ± 20	1474	145 ± 72
S3-3	23 ± 32	2015.6	–	–	–	–	–
S3-96	-3 ± 24	2009.4	E42	40 ± 40	40 ± 40	–	–
S3-19	–	–	E43	-114 ± 50	-114 ± 50	507	-46 ± 57
S3-26	–	–	E45	63 ± 30	63 ± 30	725	117 ± 37
S3-30	10 ± 17*	2008.4	E47	91 ± 30	56 ± 20	–	–
IRS 13E1	-10 ± 4	2009.4	E46	71 ± 20	71 ± 20	–	–
S3-190	-263 ± 22	2007.5	–	–	–	–	–
S3-331	–	–	E52	-167 ± 20	-167 ± 20	1892	-153 ± 60
S3-374	–	–	E53	29 ± 20	20 ± 20	847	52 ± 23
S4-71	–	–	E55	76 ± 20	60 ± 50	785	104 ± 170
S4-169	158 ± 44	2010.3	E57	196 ± 40	196 ± 40	–	–
S4-262	41 ± 22	2010.3	–	–	–	–	–
S4-314	154 ± 51	2010.3	–	–	–	–	–
S4-364	–	–	E62	-134 ± 40	-134 ± 40	516	-46 ± 64
S5-106	-29 ± 17	2019.4	–	–	–	–	–
S5-237	42 ± 18	2010.3	–	–	–	–	–
S5-183	-146 ± 17	2010.6	–	–	-135 ± 30	372	-148 ± 23
S5-191	127 ± 56	2010.6	–	–	140 ± 50	–	–
S6-89	–	–	–	–	-135 ± 70	567	-120 ± 77
S6-96	–	–	–	–	-35 ± 50	973	133 ± 70
S6-81	-22 ± 5	2010.3	E67	8 ± 20	8 ± 20	205	32 ± 16
S6-63	–	–	E69	153 ± 50	110 ± 50	227	154 ± 28
S7-30	-35 ± 37	2010.4	–	–	–	–	–
S7-5	–	–	–	–	–	596	120 ± 52
S7-10	–	–	E73	-92 ± 40	-92 ± 40	445	-61 ± 21
S7-216	–	–	–	–	60 ± 50	96	136 ± 17
S7-236	–	–	–	–	-170 ± 70	838	78 ± 64
S8-5	–	–	–	–	–	209	87 ± 31

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Table 3.1 – continued from previous page

Name	RV _{GCOWS} ^a (km s ⁻¹)	t0 _{GCOWS}	ID _{P06}	RV _{P06} ^b (km s ⁻¹)	RV _{B09} ^c (km s ⁻¹)	ID _{F15}	RV _{F15} ^d (km s ⁻¹)
S8-4	–	–	E75	-138 ± 40	-138 ± 40	230	-114 ± 20
S8-196	–	–	–	–	190 ± 50	728	277 ± 166
S8-10	–	–	–	–	–	721	-22 ± 44
S8-8	–	–	–	–	–	610	-217 ± 211
S8-126	–	–	–	–	–	718	9 ± 88
S9-143	–	–	–	–	40 ± 100	958	46 ± 68
S9-6	–	–	–	–	–	366	257 ± 135
S9-221	–	–	–	–	–	757	184 ± 89
S10-50	35 ± 43*	2010.4	–	–	–	–	–
S10-4	–	–	E84	-250 ± 40	-250 ± 40	273	-165 ± 25
S10-32	146 ± 11	2010.4	–	–	–	–	–
S10-185	–	–	–	–	–	617	202 ± 102
S10-34	-107 ± 38	2010.6	–	–	–	–	–
S10-232	–	–	–	–	–	511	231 ± 124
S10-238	–	–	–	–	–	166	154 ± 102
S10-48	-285 ± 45	2013.4	E86	-205 ± 50	–	–	–
S11-147	–	–	–	–	–	1103	53 ± 92
S11-21	-81 ± 13	2016.6	E87	-120 ± 30	-160 ± 70	–	–
S11-176	148 ± 69	2014.4	–	–	–	–	–
S11-8	–	–	–	–	–	890	31 ± 45
S11-214	–	–	–	–	–	853	257 ± 124
S11-246	132 ± 18	2014.4	–	–	–	–	–
S12-76	-159 ± 32	2013.4	E89	-100 ± 40	–	–	–
S12-178	–	–	–	–	–	722	268 ± 30
S12-5	–	–	–	–	–	483	-180 ± 42
S14-196	–	–	–	–	–	2048	130 ± 52
S7-180	–	–	–	–	120 ± 70	1554	222 ± 31
S8-70	-137 ± 66	2010.4	–	–	–	–	–
S10-261	77 ± 19	2014.4	–	–	–	–	–
S0-26	–	–	E8	30 ± 90	–	–	–
S1-1	–	–	–	–	536 ± 30	–	–
IRS 16C	–	–	E20	125 ± 30	158 ± 40	–	–
IRS 16NW	–	–	E19	-44 ± 20	-15 ± 50	–	–
S1-12	–	–	E21	-24 ± 30	8 ± 22	–	–
S1-14	–	–	E22	-434 ± 50	-400 ± 100	–	–
IRS 16SW	–	–	E23	320 ± 40	470 ± 50	–	–
S1-21	–	–	E24	-344 ± 50	-277 ± 50	–	–

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Name	RV _{GCOWS} ^a (km s ⁻¹)	t0 _{GCOWS}	ID _{P06}	RV _{P06} ^b (km s ⁻¹)	RV _{B09} ^c (km s ⁻¹)	ID _{F15}	RV _{F15} ^d (km s ⁻¹)
IRS 29N	–	–	E31	-190 ± 90	-190 ± 90	–	–
S2-7	–	–	E29	-94 ± 50	–	–	–
IRS 16SW-E	–	–	E32	366 ± 70	366 ± 70	–	–
S2-22	–	–	–	–	49 ± 50	–	–
S2-16	–	–	E35	-100 ± 70	-100 ± 70	–	–
IRS 16NE	–	–	E39	-10 ± 20	-10 ± 20	–	–
S3-5	–	–	E40	327 ± 100	327 ± 100	–	–
IRS 33E	–	–	E41	170 ± 20	170 ± 20	–	–
S3-25	–	–	E44	-114 ± 40	-84 ± 6	–	–
S3-10	–	–	E50	281 ± 20	305 ± 70	–	–
IRS 13E4	–	–	E48	56 ± 70	56 ± 70	–	–
IRS 13E3b	–	–	E49	87 ± 20	–	–	–
IRS 13E2	–	–	E51	40 ± 40	40 ± 40	–	–
S4-36	–	–	E54	-154 ± 25	-154 ± 25	–	–
IRS 34W	–	–	E56	-290 ± 30	-290 ± 30	–	–
IRS 7SE	–	–	E59	-150 ± 100	-150 ± 100	–	–
S4-258	–	–	E60	330 ± 80	330 ± 80	–	–
IRS 34NW	–	–	E61	-150 ± 30	-150 ± 30	–	–
S5-34	–	–	–	–	-40 ± 70	–	–
IRS 1W	–	–	E63	35 ± 20	35 ± 20	–	–
S5-235	–	–	–	–	-115 ± 50	–	–
S5-236	–	–	–	–	155 ± 50	–	–
S5-187	–	–	–	–	10 ± 50	–	–
S5-231	–	–	E64	40 ± 25	24 ± 25	–	–
IRS 9W	–	–	E65	140 ± 50	140 ± 50	–	–
S6-90	–	–	E66	-350 ± 50	-350 ± 50	–	–
S6-95	–	–	E68	-305 ± 100	-305 ± 100	–	–
S6-93	–	–	E70	-80 ± 100	-80 ± 100	–	–
S6-100	–	–	E71	-300 ± 150	–	–	–
S6-82	–	–	E72	86 ± 100	86 ± 100	–	–
S7-161	–	–	–	–	-120 ± 50	–	–
S7-16	–	–	–	–	160 ± 50	–	–
S7-19	–	–	–	–	-65 ± 50	–	–
S7-20	–	–	–	–	-45 ± 50	–	–
S7-228	–	–	–	–	150 ± 30	–	–
S8-15	–	–	–	–	-130 ± 50	–	–
S8-7	–	–	–	–	30 ± 100	–	–
S8-181	–	–	E74	70 ± 70	70 ± 70	–	–

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Table 3.1 – continued from previous page

Name	RV _{GCOWS} ^a (km s ⁻¹)	t0 _{GCOWS}	ID _{P06}	RV _{P06} ^b (km s ⁻¹)	RV _{B09} ^c (km s ⁻¹)	ID _{F15}	RV _{F15} ^d (km s ⁻¹)
S9-20	–	–	E76	180 ± 80	180 ± 80	–	–
S9-23	–	–	E77	-155 ± 50	-185 ± 50	–	–
S9-13	–	–	–	–	-160 ± 50	–	–
S9-1	–	–	E78	-230 ± 100	-230 ± 100	–	–
S9-114	–	–	E79	160 ± 30	160 ± 50	–	–
S9-283	–	–	E81	30 ± 70	30 ± 70	–	–
S9-9	–	–	E80	130 ± 100	130 ± 100	–	–
S10-136	–	–	E82	-70 ± 70	-70 ± 70	–	–
S10-5	–	–	E83	-180 ± 70	-180 ± 70	–	–
S10-7	–	–	E85	-150 ± 40	-150 ± 40	–	–
S11-5	–	–	E88	-65 ± 40	-65 ± 40	–	–
S13-3	–	–	E90	-190 ± 40	–	–	–

^a RV measurements from GCOWS. Stars with multiple measurements in our database is marked with asterisk and only one of the measurements is showed here;

^b RV measurements from [16];

^c RV measurements from [18];

^d RV measurements from [20];

Table 3.2: Final RVs

Name	RV _{use} (km s ⁻¹)	ref ^a	WR (km s ⁻¹)	Data Available ^b (km s ⁻¹)
S0-20	398 ± 12	1	-	RV + C + PM
S0-2	multiple RVs	1	-	RV + C + PM
S0-1	multiple RVs	1	-	RV + C + PM
S0-8	multiple RVs	1	-	RV + C + PM
S0-16	151 ± 6	1	-	RV + C + PM
S0-3	multiple RVs	1	-	RV + C + PM
S0-5	687 ± 5	1	-	RV + C + PM
S0-19	multiple RVs	1	-	RV + C + PM
S0-11	-22 ± 4	1	-	RV + C + PM
S0-7	98 ± 5	1	-	RV + C + PM
S0-4	-168 ± 5	1	-	RV + C + PM
S0-30	-16 ± 71	1	-	RV + C + PM

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Table 3.2 – continued from previous page

Name	RV _{use} (km s ⁻¹)	ref ^a	WR (km s ⁻¹)	Data Available ^b (km s ⁻¹)
S0-9	95 ± 4	1	-	RV + C + PM
S0-31	-120 ± 10	1	-	RV + C + PM
S0-14	-32 ± 2	1	-	RV + C + PM
S1-3	20 ± 30	3,4	-	RV + C + PM
S0-15	multiple RVs	1	-	RV + C + PM
S1-2	multiple RVs	1	-	RV + C + PM
S1-8	-112 ± 7	1	-	RV + C + PM
S1-4	-221 ± 179	4	-	RV + C + PM
S1-33	26 ± 7	1	-	RV + C + PM
S1-22	-230 ± 33	3,4	-	RV + C + PM
S1-19	-169 ± 16	1	-	RV + C + PM
S1-24	125 ± 10	1	-	RV + C + PM
IRS 16CC	253 ± 11	2,3,4	-	RV + C + PM
S2-4	218 ± 6	1	-	RV + C + PM
S2-6	152 ± 8	1	-	RV + C + PM
IRS 33N	23 ± 5	1	-	RV + C + PM
S2-50	-135 ± 32	1	-	RV + C + PM
S2-17	57 ± 4	1	-	RV + C + PM
S2-21	-92 ± 12	1	-	RV + C + PM
S2-19	54 ± 19	2,3,4	-	RV + C + PM
S2-58	77 ± 16	1	-	RV + C + PM
S2-74	44 ± 19	2,3,4	-	RV + C + PM
S3-3	23 ± 32	1	-	RV + C + PM
S3-96	-3 ± 24	1	-	RV + C + PM
S3-19	-84 ± 38	2,3,4	-	RV + C + PM
S3-26	84 ± 23	2,3,4	-	RV + C + PM
S3-30	6 ± 15	1	-	RV + C + PM
IRS 13E1	-10 ± 4	1	-	RV + C + PM
S3-190	-262 ± 22	1	-	RV + C + PM
S3-331	-166 ± 19	2,3,4	-	RV + C + PM
S3-374	34 ± 15	3,4	-	RV + C + PM
S4-71	64 ± 48	3,4	-	RV + C + PM
S4-169	158 ± 44	1	-	RV + C + PM
S4-262	41 ± 22	1	-	RV + C + PM
S4-314	154 ± 51	1	-	RV + C + PM
S4-364	-109 ± 34	2,3,4	-	RV + C + PM
S5-106	-29 ± 17	1	-	RV + C + PM
S5-237	42 ± 18	1	-	RV + C + PM

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Table 3.2 – continued from previous page

Name	RV _{use} (km s ⁻¹)	ref ^a	WR (km s ⁻¹)	Data Available ^b (km s ⁻¹)
S5-183	-146 ± 17	1	-	RV + C + PM
S5-191	127 ± 56	1	-	RV + C + PM
S6-89	-128 ± 52	3,4	-	RV + C + PM
S6-96	22 ± 41	3,4	-	RV + C + PM
S6-81	-22 ± 5	1	-	RV + C + PM
S6-63	144 ± 24	3,4	-	RV + C + PM
S7-30	-35 ± 37	1	-	RV + C + PM
S7-5	120 ± 52	4	-	RV + C + PM
S7-10	-68 ± 19	2,3,4	-	RV + C + PM
S7-216	128 ± 16	3,4	-	RV + C + PM
S7-236	78 ± 64	4	-	RV + C + PM
S8-5	87 ± 31	4	-	RV + C + PM
S8-4	-119 ± 18	2,3,4	-	RV + C + PM
S8-196	197 ± 48	3,4	-	RV + C + PM
S8-10	-22 ± 44	4	-	RV + C + PM
S8-8	-217 ± 211	4	-	RV + C + PM
S8-126	9 ± 88	4	-	RV + C + PM
S9-143	44 ± 56	3,4	-	RV + C + PM
S9-6	257 ± 135	4	-	RV + C + PM
S9-221	184 ± 89	4	-	RV + C + PM
S10-50	70 ± 23	1	-	RV + C + PM
S10-4	-189 ± 21	2,3,4	-	RV + C + PM
S10-32	146 ± 11	1	-	RV + C + PM
S10-185	202 ± 102	4	-	RV + C + PM
S10-34	-107 ± 38	1	-	RV + C + PM
S10-232	231 ± 124	4	-	RV + C + PM
S10-238	154 ± 102	4	-	RV + C + PM
S10-48	-285 ± 45	1	-	RV + C + PM
S11-147	53 ± 92	4	-	RV + C + PM
S11-21	-81 ± 13	1	-	RV + C + PM
S11-176	148 ± 69	1	-	RV + C + PM
S11-8	31 ± 45	4	-	RV + C + PM
S11-214	257 ± 124	4	-	RV + C + PM
S11-246	132 ± 18	1	-	RV + C + PM
S12-76	-159 ± 32	1	-	RV + C + PM
S12-178	268 ± 30	4	-	RV + C + PM
S12-5	-180 ± 42	4	-	RV + C + PM
S14-196	130 ± 52	4	-	RV + C + PM

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Table 3.2 – continued from previous page

Name	RV _{use} (km s ⁻¹)	ref ^a	WR (km s ⁻¹)	Data Available ^b (km s ⁻¹)
S7-180	205 ± 28	3,4	-	RV + C
S8-70	-137 ± 66	1	-	RV + C
S10-261	77 ± 19	1	-	RV + C
S0-26	30 ± 90	2	-	RV + PM
S1-1	536 ± 30	3	-	RV + PM
IRS 16C	158 ± 40	3	Y	RV + PM
IRS 16NW	-15 ± 50	3	Y	RV + PM
S1-12	8 ± 22	3	-	RV + PM
S1-14	-400 ± 100	3	-	RV + PM
IRS 16SW	470 ± 50	3	Y	RV + PM
S1-21	-277 ± 50	3	-	RV + PM
IRS 29N	-190 ± 90	2,3	Y	RV + PM
S2-7	-94 ± 50	2	-	RV + PM
IRS 16SW-E	366 ± 70	2,3	Y	RV + PM
S2-22	49 ± 50	3	-	RV + PM
S2-16	-100 ± 70	2,3	Y	RV + PM
IRS 16NE	-10 ± 20	2,3	Y	RV + PM
S3-5	327 ± 100	2,3	Y	RV + PM
IRS 33E	170 ± 20	2,3	Y	RV + PM
S3-25	-84 ± 6	3	-	RV + PM
S3-10	305 ± 70	3	-	RV + PM
IRS 13E4	56 ± 70	2,3	Y	RV + PM
IRS 13E3b	87 ± 20	2	Y	RV + PM
IRS 13E2	40 ± 40	2,3	Y	RV + PM
S4-36	-154 ± 25	2,3	-	RV + PM
IRS 34W	-290 ± 30	2,3	Y	RV + PM
IRS 7SE	-150 ± 100	2,3	Y	RV + PM
S4-258	330 ± 80	2,3	Y	RV + PM
IRS 34NW	-150 ± 30	2,3	Y	RV + PM
S5-34	-40 ± 70	3	-	RV + PM
IRS 1W	35 ± 20	2,3	-	RV + PM
S5-235	-115 ± 50	3	-	RV + PM
S5-236	155 ± 50	3	-	RV + PM
S5-187	10 ± 50	3	-	RV + PM
S5-231	24 ± 25	3	-	RV + PM
IRS 9W	140 ± 50	2,3	Y	RV + PM
S6-90	-350 ± 50	2,3	Y	RV + PM

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Table 3.2 – continued from previous page

Name	RV_{use} (km s ⁻¹)	ref ^a	WR (km s ⁻¹)	Data Available ^b (km s ⁻¹)
S6-95	-305 ± 100	2,3	Y	RV + PM
S6-93	-80 ± 100	2,3	Y	RV + PM
S6-100	-300 ± 150	2	Y	RV + PM
S6-82	86 ± 100	2,3	-	RV + PM
S7-161	-120 ± 50	3	-	RV + PM
S7-16	160 ± 50	3	-	RV + PM
S7-19	-65 ± 50	3	-	RV + PM
S7-20	-45 ± 50	3	-	RV + PM
S7-228	150 ± 30	3	-	RV + PM
S8-15	-130 ± 50	3	-	RV + PM
S8-7	30 ± 100	3	-	RV + PM
S8-181	70 ± 70	2,3	Y	RV + PM
S9-20	180 ± 80	2,3	Y	RV + PM
S9-23	-185 ± 50	3	-	RV + PM
S9-13	-160 ± 50	3	-	RV + PM
S9-1	-230 ± 100	2,3	Y	RV + PM
S9-114	160 ± 50	3	Y	RV + PM
S9-283	30 ± 70	2,3	Y	RV + PM
S9-9	130 ± 100	2,3	Y	RV + PM
S10-136	-70 ± 70	2,3	Y	RV + PM
S10-5	-180 ± 70	2,3	Y	RV + PM
S10-7	-150 ± 40	2,3	-	RV + PM
S11-5	-65 ± 40	2,3	Y	RV + PM
S13-3	-190 ± 40	2	-	RV + PM

^a Reference for final RV_{use} : 1 - database, 2 - P06[16], 3 - B09[18], 4 - F15[20];

^b The available data for each star: RV - Radial Velocity, C - spectral completeness, PM - proper motion. See details in section 3.1, section 3.1 and section 3.1. Our final sample of 88 stars have RV, proper motion and spectral completeness;

Table 3.3: Proper Motions Summary (1)

Name	Kp	Kp _{ext} ^a	ar (km s ⁻¹ yr ⁻¹)
S0-20	15.7	16.1	-167.68 ± 17.25
S0-2	14.1	14.5	-111.17 ± 12.27
S0-1	14.7	15.2	-103.22 ± 2.09
S0-8	15.8	16.1	-77.39 ± 2.75
S0-16	15.3	15.8	-41.50 ± 15.75
S0-3	14.5	14.9	-43.15 ± 1.09
S0-5	15.1	15.6	-62.51 ± 0.82
S0-19	15.5	15.9	-59.35 ± 1.77
S0-11	15.1	15.5	0.50 ± 0.25
S0-7	15.4	15.8	-1.47 ± 0.46
S0-4	14.3	14.7	-26.62 ± 0.29
S0-30	16.1	16.7	3.19 ± 4.84
S0-9	14.2	14.8	-0.33 ± 0.22
S0-31	15.1	15.5	-0.31 ± 0.36
S0-14	13.7	14.3	0.09 ± 0.11
S1-3	12.1	12.5	-5.57 ± 0.14
S0-15	13.6	14.1	-11.81 ± 0.39
S1-2	14.8	15.4	-3.38 ± 0.21
S1-8	14.2	14.7	-0.70 ± 0.13
S1-4	12.6	13.0	-6.53 ± 0.23
S1-33	14.9	15.5	0.65 ± 0.26
S1-22	12.6	13.1	0.42 ± 0.18
S1-19	13.6	14.3	-0.41 ± 0.26
S1-24	11.5	12.1	0.69 ± 0.08
IRS 16CC	10.7	11.0	-3.64 ± 2.44
S2-4	12.1	12.6	-0.24 ± 0.11
S2-6	12.0	12.5	-0.83 ± 0.10
IRS 33N	11.3	12.0	0.19 ± 0.13
S2-50	15.4	15.9	-2.46 ± 0.70
S2-17	10.7	11.4	0.01 ± 0.30
S2-21	13.4	13.9	-0.47 ± 0.15
S2-19	12.6	13.0	-0.90 ± 0.26
S2-58	14.1	14.4	0.50 ± 0.20
S2-74	13.2	13.6	-0.07 ± 0.17
S3-3	15.1	15.4	0.52 ± 0.15
S3-96	14.3	14.9	1.62 ± 1.56

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Table 3.3 – continued from previous page

Name	Kp	Kp _{ext} ^a	ar (km s ⁻¹ yr ⁻¹)
S3-19	11.9	12.8	-1.62 ± 1.05
S3-26	12.3	13.0	0.42 ± 0.17
S3-30	12.4	13.3	0.08 ± 0.17
IRS 13E1	10.6	11.3	0.27 ± 0.17
S3-190	14.0	14.5	-0.32 ± 0.28
S3-331	13.5	13.8	0.71 ± 0.55
S3-374	12.3	12.9	-0.14 ± 0.15
S4-71	12.3	13.2	0.27 ± 0.20
S4-169	13.5	13.7	-1.11 ± 0.19
S4-262	16.2	16.7	0.35 ± 0.21
S4-314	15.3	15.8	-0.64 ± 0.29
S4-364	11.7	12.1	–
S5-106	15.5	16.2	-1.51 ± 0.53
S5-237	13.2	13.6	-3.92 ± 2.66
S5-183	11.6	12.2	-0.46 ± 0.11
S5-191	12.8	13.6	0.42 ± 0.18
S6-89	12.0	12.8	-3.43 ± 2.14
S6-96	12.8	13.1	–
S6-81	11.0	11.6	–
S6-63	11.2	12.1	1.53 ± 1.11
S7-30	13.9	14.7	–
S7-5	12.2	12.9	0.16 ± 1.01
S7-10	11.4	11.7	–
S7-216	10.7	11.0	–
S7-236	12.5	12.8	–
S8-5	11.1	11.9	–
S8-4	11.0	11.3	–
S8-196	12.4	12.7	–
S8-10	12.2	12.5	–
S8-8	12.1	12.8	–
S8-126	12.5	13.2	–
S9-143	12.6	12.8	–
S9-6	11.4	12.2	–
S9-221	12.1	12.3	–
S10-50	14.7	15.6	–
S10-4	11.2	11.5	–
S10-32	14.4	14.9	–
S10-185	12.2	12.4	–

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Table 3.3 – continued from previous page

Name	Kp	Kp _{ext} ^a	ar (km s ⁻¹ yr ⁻¹)
S10-34	14.5	15.4	–
S10-232	11.9	12.4	–
S10-238	10.9	11.3	–
S10-48	15.1	15.3	–
S11-147	12.8	13.2	–
S11-21	13.5	13.8	–
S11-176	15.3	16.0	–
S11-8	12.6	12.8	–
S11-214	12.4	13.1	–
S11-246	14.2	14.9	–
S12-76	14.5	15.1	–
S12-178	12.2	13.0	–
S12-5	11.6	11.9	–
S14-196	15.8	16.5	–
S7-180	13.4	13.5	–
S8-70	15.3	15.7	–
S10-261	14.6	15.4	–
S0-26	15.3	15.7	21.29 ± 17.591
S1-1	13.1	13.3	-0.20 ± 0.09
IRS 16C	9.9	10.3	-2.45 ± 0.22
IRS 16NW	10.1	10.6	–
S1-12	13.6	14.0	-0.16 ± 0.11
S1-14	12.7	13.4	-1.48 ± 0.11
IRS 16SW	10.1	10.6	-3.73 ± 0.17
S1-21	13.2	13.7	-0.15 ± 0.24
IRS 29N	10.5	11.1	1.33 ± 0.84
S2-7	14.2	14.6	0.14 ± 0.26
IRS 16SW-E	11.1	11.5	-1.37 ± 0.26
S2-22	12.9	13.2	-2.26 ± 0.28
S2-16	12.0	12.4	0.44 ± 0.38
IRS 16NE	9.1	9.2	-0.62 ± 1.54
S3-5	12.0	12.4	-0.60 ± 0.11
IRS 33E	10.2	11.0	-0.38 ± 0.31
S3-25	13.9	14.2	-0.36 ± 0.18
S3-10	12.1	12.6	-0.78 ± 0.14
IRS 13E4	11.7	12.4	1.82 ± 0.71
IRS 13E3b	12.4	13.1	-2.96 ± 8.22

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Table 3.3 – continued from previous page

Name	Kp	Kp _{ext} ^a	ar (km s ⁻¹ yr ⁻¹)
IRS 13E2	10.7	11.3	-0.03 ± 0.18
S4-36	12.6	12.8	0.39 ± 0.36
IRS 34W	11.4	11.6	0.64 ± 0.42
IRS 7SE	13.1	13.5	2.26 ± 0.83
S4-258	12.6	13.0	-0.29 ± 0.26
IRS 34NW	13.3	13.6	-0.75 ± 0.70
S5-34	13.6	14.3	-0.49 ± 0.32
IRS 1W	10.9	11.3	6.99 ± 5.12
S5-235	13.2	13.6	–
S5-236	13.1	13.4	–
S5-187	13.2	14.1	-0.62 ± 0.64
S5-231	12.0	12.4	–
IRS 9W	12.1	12.9	-0.25 ± 0.40
S6-90	12.3	12.6	–
S6-95	13.2	13.1	–
S6-93	12.8	13.4	–
S6-100	13.9	14.1	–
S6-82	13.5	14.1	–
S7-161	13.6	13.8	–
S7-16	12.5	13.4	–
S7-19	13.2	13.6	–
S7-20	13.3	13.6	–
S7-228	11.8	12.2	–
S8-15	13.0	13.2	–
S8-7	11.9	12.7	–
S8-181	11.6	11.8	–
S9-20	13.2	14.0	–
S9-23	13.6	13.7	–
S9-13	13.1	13.1	–
S9-1	12.7	12.9	–
S9-114	10.8	11.2	–
S9-283	12.5	12.9	–
S9-9	11.7	12.5	–
S10-136	13.0	13.1	–
S10-5	11.9	12.2	–
S10-7	12.7	13.1	–
S11-5	11.8	12.4	–
S13-3	11.9	12.2	–

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Table 3.3 – continued from previous page

Name	Kp	Kp _{ext} ^a	ar
			(km s ⁻¹ yr ⁻¹)

^a Extinction corrected Kp to Aks=2.7e;

Table 3.4: Proper Motions Summary (2)

Name	x0	y0	vx	vy	t0	ref ^a
	(")	(")	(mas yr ⁻¹)	(mas yr ⁻¹)		
S0-20	0.162 ± 0.00665	0.096 ± 0.00824	32.32 ± 2.17	-8.66 ± 2.22	2011.3	1
S0-2	-0.029 ± 0.00213	0.160 ± 0.00390	-6.50 ± 0.91	-2.13 ± 1.54	2010.2	1
S0-1	0.111 ± 0.00076	-0.294 ± 0.00296	19.00 ± 0.32	-7.58 ± 1.39	2009.9	1
S0-8	-0.179 ± 0.00153	0.099 ± 0.00065	22.49 ± 1.02	-28.73 ± 0.34	2010.3	1
S0-16	0.230 ± 0.00586	0.171 ± 0.00234	16.51 ± 2.12	13.17 ± 1.63	2008.0	1
S0-3	0.340 ± 0.00115	0.114 ± 0.00034	7.62 ± 0.58	-1.71 ± 0.16	2009.8	1
S0-5	0.170 ± 0.00083	-0.359 ± 0.00175	-3.48 ± 0.39	2.49 ± 0.80	2009.8	1
S0-19	-0.007 ± 0.00037	0.400 ± 0.00201	8.52 ± 0.10	16.10 ± 0.98	2009.7	1
S0-11	0.490 ± 0.00006	-0.066 ± 0.00010	-3.53 ± 0.02	-2.66 ± 0.05	2011.3	1
S0-7	0.516 ± 0.00007	0.099 ± 0.00006	5.86 ± 0.02	0.66 ± 0.02	2010.6	1
S0-4	0.442 ± 0.00068	-0.329 ± 0.00052	10.06 ± 0.33	-10.77 ± 0.22	2009.7	1
S0-30	-0.513 ± 0.00012	-0.379 ± 0.00009	-10.61 ± 0.06	3.27 ± 0.09	2009.1	1
S0-9	0.226 ± 0.00005	-0.605 ± 0.00006	8.99 ± 0.02	-5.39 ± 0.02	2010.1	1
S0-31	0.565 ± 0.00007	0.445 ± 0.00006	6.37 ± 0.02	1.05 ± 0.02	2009.7	1
S0-14	-0.756 ± 0.00003	-0.287 ± 0.00003	2.45 ± 0.01	-1.47 ± 0.01	2010.4	1
S1-3	0.339 ± 0.00008	0.882 ± 0.00019	-14.01 ± 0.03	1.40 ± 0.08	2009.1	1
S0-15	-0.965 ± 0.00043	0.182 ± 0.00007	-2.43 ± 0.15	-11.76 ± 0.03	2009.8	1
S1-2	0.077 ± 0.00005	-1.019 ± 0.00009	13.17 ± 0.02	-0.29 ± 0.04	2010.1	1
S1-8	-0.581 ± 0.00003	-0.920 ± 0.00003	9.26 ± 0.01	-5.69 ± 0.01	2010.3	1
S1-4	0.858 ± 0.00019	-0.664 ± 0.00014	11.31 ± 0.06	2.87 ± 0.08	2007.7	1
S1-33	-1.245 ± 0.00005	0.002 ± 0.00005	-0.15 ± 0.02	5.17 ± 0.01	2011.0	1
S1-22	-1.567 ± 0.00004	-0.524 ± 0.00003	8.25 ± 0.01	-3.26 ± 0.01	2010.2	1
S1-19	0.435 ± 0.00007	-1.639 ± 0.00008	8.54 ± 0.02	-3.38 ± 0.02	2010.3	1
S1-24	0.736 ± 0.00002	-1.655 ± 0.00002	2.69 ± 0.01	-6.58 ± 0.01	2010.3	1
IRS 16CC	1.996 ± 0.00016	0.548 ± 0.00023	-1.51 ± 0.09	6.53 ± 0.07	2002.2	1
S2-4	1.521 ± 0.00004	-1.458 ± 0.00003	8.14 ± 0.01	2.80 ± 0.01	2010.3	1
S2-6	1.663 ± 0.00004	-1.334 ± 0.00004	8.04 ± 0.01	2.10 ± 0.01	2010.2	1
IRS 33N	-0.026 ± 0.00002	-2.245 ± 0.00002	3.64 ± 0.01	-5.97 ± 0.01	2010.6	1

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Table 3.4 – continued from previous page

Name	x0 (")	y0 (")	vx (mas yr ⁻¹)	vy (mas yr ⁻¹)	t0	ref ^a
S2-50	1.701 ± 0.00020	-1.506 ± 0.00021	1.98 ± 0.05	2.04 ± 0.05	2010.0	1
S2-17	1.340 ± 0.00004	-1.878 ± 0.00003	9.19 ± 0.01	-0.20 ± 0.01	2009.6	1
S2-21	-1.612 ± 0.00004	-1.676 ± 0.00003	9.54 ± 0.01	-3.85 ± 0.01	2010.7	1
S2-19	0.377 ± 0.00003	2.307 ± 0.00004	-8.18 ± 0.01	0.78 ± 0.01	2010.5	1
S2-58	2.145 ± 0.00007	-1.125 ± 0.00006	-0.81 ± 0.02	6.43 ± 0.02	2010.4	1
S2-74	0.111 ± 0.00001	2.779 ± 0.00003	-8.67 ± 0.01	1.10 ± 0.01	2010.6	1
S3-3	3.093 ± 0.00002	-0.640 ± 0.00005	3.68 ± 0.01	4.11 ± 0.02	2010.9	1
S3-96	-3.131 ± 0.00016	-0.625 ± 0.00012	-0.02 ± 0.06	5.20 ± 0.07	2010.5	1
S3-19	-1.555 ± 0.00005	-2.795 ± 0.00004	8.00 ± 0.02	-1.76 ± 0.02	2009.0	1
S3-26	-2.569 ± 0.00003	-2.071 ± 0.00002	5.99 ± 0.01	1.47 ± 0.01	2010.7	1
S3-30	1.660 ± 0.00003	-2.933 ± 0.00003	-0.80 ± 0.01	4.00 ± 0.01	2010.7	1
IRS 13E1	-2.979 ± 0.00003	-1.660 ± 0.00002	-3.65 ± 0.01	-2.66 ± 0.01	2010.5	1
S3-190	-3.191 ± 0.00004	1.410 ± 0.00005	-2.80 ± 0.01	-3.06 ± 0.02	2010.6	1
S3-331	-1.224 ± 0.00005	3.655 ± 0.00008	5.96 ± 0.02	4.23 ± 0.02	2010.4	1
S3-374	-2.756 ± 0.00004	-2.855 ± 0.00002	-0.28 ± 0.01	-4.39 ± 0.01	2010.8	1
S4-71	0.771 ± 0.00003	-4.095 ± 0.00004	0.08 ± 0.01	-4.60 ± 0.01	2010.8	1
S4-169	4.411 ± 0.00004	0.278 ± 0.00004	-2.53 ± 0.01	3.97 ± 0.01	2010.4	1
S4-262	4.276 ± 0.00002	-1.964 ± 0.00004	-1.40 ± 0.01	-5.44 ± 0.01	2011.4	1
S4-314	4.409 ± 0.00004	-2.031 ± 0.00005	-0.13 ± 0.02	2.26 ± 0.02	2010.5	1
S4-364	2.248 ± 0.00016	4.464 ± 0.00021	6.86 ± 0.04	-2.97 ± 0.06	2011.1	2
S5-106	-4.351 ± 0.00008	-3.200 ± 0.00008	-3.77 ± 0.02	-2.04 ± 0.02	2011.5	1
S5-237	5.493 ± 0.00025	1.022 ± 0.00031	-1.31 ± 0.14	6.19 ± 0.12	2012.2	1
S5-183	4.591 ± 0.00002	-3.444 ± 0.00002	-4.52 ± 0.01	-2.13 ± 0.01	2010.4	1
S5-191	3.181 ± 0.00002	-4.889 ± 0.00003	-1.52 ± 0.01	-3.78 ± 0.01	2010.7	1
S6-89	5.450 ± 0.00031	2.994 ± 0.00016	2.88 ± 0.10	-6.45 ± 0.06	2011.3	1
S6-96	-6.049 ± 0.00050	-1.923 ± 0.00052	-0.80 ± 0.12	7.28 ± 0.15	2010.3	2
S6-81	6.353 ± 0.00020	0.272 ± 0.00023	-2.69 ± 0.05	5.06 ± 0.05	2010.3	2
S6-63	1.872 ± 0.00007	-6.305 ± 0.00014	5.85 ± 0.02	1.66 ± 0.04	2011.1	1
S7-30	6.460 ± 0.00020	-2.702 ± 0.00024	-3.01 ± 0.04	-3.92 ± 0.06	2011.5	2
S7-5	4.860 ± 0.00010	-5.522 ± 0.00008	2.45 ± 0.05	5.24 ± 0.04	2010.5	1
S7-10	-1.117 ± 0.00024	7.622 ± 0.00031	-4.47 ± 0.04	-2.30 ± 0.07	2011.3	2
S7-216	-7.747 ± 0.00031	1.456 ± 0.00039	2.20 ± 0.08	5.71 ± 0.10	2015.5	3
S7-236	-7.142 ± 0.00018	3.562 ± 0.00021	-2.54 ± 0.04	-4.04 ± 0.02	2015.5	3
S8-5	-3.480 ± 0.00018	-7.506 ± 0.00025	-1.10 ± 0.06	4.55 ± 0.07	2011.1	2
S8-4	-0.021 ± 0.00024	8.564 ± 0.00025	-0.42 ± 0.06	3.56 ± 0.06	2011.3	2
S8-196	-8.111 ± 0.00021	-2.914 ± 0.00031	1.14 ± 0.05	-1.55 ± 0.07	2015.5	3
S8-10	6.962 ± 0.00018	5.190 ± 0.00025	3.28 ± 0.04	-3.07 ± 0.08	2011.1	2
S8-8	5.257 ± 0.00023	-7.069 ± 0.00017	0.29 ± 0.08	0.75 ± 0.06	2011.5	2

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Table 3.4 – continued from previous page

Name	x0 (")	y0 (")	vx (mas yr ⁻¹)	vy (mas yr ⁻¹)	t0	ref ^a
S8-126	-2.240 ± 0.00034	-8.564 ± 0.00024	-4.73 ± 0.15	2.53 ± 0.08	2011.1	2
S9-143	-8.389 ± 0.00018	-3.378 ± 0.00030	0.59 ± 0.04	-2.85 ± 0.07	2015.5	3
S9-6	0.760 ± 0.00028	-9.326 ± 0.00033	1.50 ± 0.10	1.62 ± 0.10	2010.8	2
S9-221	-8.936 ± 0.00014	-3.453 ± 0.00029	-3.12 ± 0.02	-1.72 ± 0.06	2015.5	3
S10-50	9.583 ± 0.00052	-3.182 ± 0.00034	-0.78 ± 0.13	-4.04 ± 0.08	2011.8	2
S10-4	0.071 ± 0.00032	10.251 ± 0.00029	-1.93 ± 0.07	1.13 ± 0.07	2011.3	2
S10-32	10.208 ± 0.00023	-1.689 ± 0.00031	2.72 ± 0.07	3.47 ± 0.07	2011.6	2
S10-185	-8.399 ± 0.00054	-6.164 ± 0.00023	-3.01 ± 0.15	-1.11 ± 0.03	2015.5	3
S10-34	8.878 ± 0.00039	-5.622 ± 0.00034	0.74 ± 0.10	3.47 ± 0.07	2011.2	2
S10-232	-7.056 ± 0.00026	-7.934 ± 0.00029	2.01 ± 0.06	-3.20 ± 0.06	2015.5	3
S10-238	-8.927 ± 0.00017	-5.720 ± 0.00022	2.81 ± 0.03	1.75 ± 0.03	2015.5	3
S10-48	-0.525 ± 0.00042	10.729 ± 0.00048	1.92 ± 0.08	0.81 ± 0.13	2011.5	2
S11-147	-10.751 ± 0.00038	-3.053 ± 0.00059	-0.20 ± 0.09	3.79 ± 0.15	2015.5	3
S11-21	2.559 ± 0.00022	10.932 ± 0.00034	-1.82 ± 0.05	-2.45 ± 0.08	2011.4	2
S11-176	-0.185 ± 0.00034	-11.214 ± 0.00053	-1.19 ± 0.09	2.21 ± 0.13	2015.5	3
S11-8	10.876 ± 0.00014	3.246 ± 0.00021	-1.70 ± 0.03	-2.98 ± 0.05	2010.7	2
S11-214	1.539 ± 0.00018	-11.284 ± 0.00022	-1.10 ± 0.04	-0.43 ± 0.03	2015.5	3
S11-246	-0.992 ± 0.00024	-11.567 ± 0.00029	-3.30 ± 0.06	-1.56 ± 0.06	2015.5	3
S12-76	0.018 ± 0.00049	12.285 ± 0.00101	2.84 ± 0.11	1.43 ± 0.28	2011.3	2
S12-178	-6.096 ± 0.00016	-10.875 ± 0.00021	-0.01 ± 0.03	-2.06 ± 0.02	2015.5	3
S12-5	10.891 ± 0.00014	6.671 ± 0.00029	0.17 ± 0.04	-0.04 ± 0.09	2011.3	2
S14-196	-10.924 ± 0.00073	-9.937 ± 0.00105	0.33 ± 0.19	2.98 ± 0.26	2015.5	3
S7-180	–	–	–	–	–	–
S8-70	–	–	–	–	–	–
S10-261	–	–	–	–	–	–
S0-26	0.334 ± 0.00152	0.217 ± 0.00034	-3.08 ± 1.10	9.02 ± 0.26	2007.9	1
S1-1	1.038 ± 0.00004	0.032 ± 0.00003	5.90 ± 0.01	1.24 ± 0.01	2009.5	1
IRS 16C	1.050 ± 0.00013	0.549 ± 0.00012	-8.85 ± 0.04	7.49 ± 0.03	2009.6	1
IRS 16NW	0.078 ± 0.00005	1.219 ± 0.00004	6.16 ± 0.01	0.74 ± 0.01	2009.8	1
S1-12	-0.744 ± 0.00003	-1.031 ± 0.00002	11.44 ± 0.01	-2.97 ± 0.01	2010.6	1
S1-14	-1.308 ± 0.00005	-0.389 ± 0.00004	5.58 ± 0.02	-8.03 ± 0.02	2011.8	1
IRS 16SW	1.116 ± 0.00016	-0.949 ± 0.00016	7.61 ± 0.06	2.84 ± 0.05	2010.5	1
S1-21	-1.638 ± 0.00004	0.089 ± 0.00003	4.30 ± 0.01	-5.37 ± 0.01	2010.5	1
IRS 29N	-1.548 ± 0.00011	1.361 ± 0.00011	4.99 ± 0.04	-5.53 ± 0.04	2011.3	1
S2-7	0.926 ± 0.00003	1.851 ± 0.00003	-6.72 ± 0.01	1.48 ± 0.01	2010.4	1
IRS 16SW-E	1.903 ± 0.00005	-1.113 ± 0.00010	5.95 ± 0.02	3.76 ± 0.03	2011.3	1

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Table 3.4 – continued from previous page

Name	x0 (")	y0 (")	vx (mas yr ⁻¹)	vy (mas yr ⁻¹)	t0	ref ^a
S2-22	2.306 ± 0.00010	-0.225 ± 0.00005	-1.60 ± 0.05	6.10 ± 0.03	2007.0	1
S2-16	-1.074 ± 0.00004	2.057 ± 0.00005	-8.89 ± 0.01	-1.16 ± 0.02	2010.5	1
IRS 16NE	2.895 ± 0.00028	0.953 ± 0.00030	2.92 ± 0.08	-9.30 ± 0.10	2010.3	1
S3-5	2.961 ± 0.00002	-1.149 ± 0.00003	2.74 ± 0.01	4.59 ± 0.01	2010.5	1
IRS 33E	0.710 ± 0.00004	-3.138 ± 0.00005	6.84 ± 0.01	-1.31 ± 0.02	2010.7	1
S3-25	1.405 ± 0.00003	2.954 ± 0.00003	-7.02 ± 0.01	0.35 ± 0.01	2010.3	1
S3-10	3.339 ± 0.00003	-1.106 ± 0.00002	-0.33 ± 0.01	5.16 ± 0.01	2010.5	1
IRS 13E4	-3.256 ± 0.00026	-1.404 ± 0.00027	-5.94 ± 0.06	1.16 ± 0.07	2012.5	1
IRS 13E3b	-3.196 ± 0.00120	-1.531 ± 0.00177	-2.16 ± 0.32	-1.11 ± 0.46	2009.9	1
IRS 13E2	-3.206 ± 0.00003	-1.730 ± 0.00002	-6.49 ± 0.01	0.71 ± 0.01	2010.4	1
S4-36	-3.695 ± 0.00005	1.777 ± 0.00007	-4.88 ± 0.02	-4.65 ± 0.02	2010.1	1
IRS 34W	-4.070 ± 0.00006	1.551 ± 0.00006	-2.14 ± 0.02	-5.65 ± 0.02	2010.2	1
IRS 7SE	2.991 ± 0.00007	3.464 ± 0.00016	5.94 ± 0.02	-0.11 ± 0.05	2009.8	1
S4-258	-4.402 ± 0.00005	-1.632 ± 0.00002	-4.34 ± 0.02	1.62 ± 0.01	2010.1	1
IRS 34NW	-3.780 ± 0.00006	2.823 ± 0.00008	-5.43 ± 0.02	-4.04 ± 0.03	2010.4	1
S5-34	-4.336 ± 0.00005	-2.744 ± 0.00005	-3.31 ± 0.02	-2.40 ± 0.01	2010.5	1
IRS 1W	5.252 ± 0.00066	0.636 ± 0.00252	-1.22 ± 0.41	6.66 ± 0.85	2012.8	1
S5-235	2.779 ± 0.00020	4.532 ± 0.00020	-0.79 ± 0.04	-4.06 ± 0.05	2011.5	2
S5-236	-5.535 ± 0.00036	-1.287 ± 0.00041	4.97 ± 0.11	1.48 ± 0.14	2010.7	2
S5-187	-1.711 ± 0.00008	-5.546 ± 0.00011	-0.92 ± 0.03	-4.11 ± 0.04	2009.7	1
S5-231	5.813 ± 0.00018	0.115 ± 0.00018	-0.14 ± 0.04	6.51 ± 0.04	2012.0	2
IRS 9W	2.899 ± 0.00003	-5.589 ± 0.00004	5.40 ± 0.01	3.45 ± 0.02	2010.3	1
S6-90	-3.951 ± 0.00026	4.906 ± 0.00027	0.28 ± 0.06	-3.80 ± 0.07	2011.3	2
S6-95	-2.401 ± 0.00023	5.998 ± 0.00035	5.34 ± 0.05	0.07 ± 0.09	2011.3	2
S6-93	4.467 ± 0.00020	4.960 ± 0.00022	5.08 ± 0.04	-1.31 ± 0.06	2011.6	2
S6-100	1.548 ± 0.00036	6.525 ± 0.00125	-4.46 ± 0.11	3.20 ± 0.50	2010.8	2
S6-82	6.722 ± 0.00022	-0.458 ± 0.00047	1.67 ± 0.07	4.90 ± 0.16	2011.3	2
S7-161	-7.420 ± 0.00018	0.033 ± 0.00070	-2.08 ± 0.04	-3.61 ± 0.19	2015.5	3
S7-16	1.629 ± 0.00019	-7.227 ± 0.00019	2.08 ± 0.06	3.62 ± 0.06	2011.3	2
S7-19	-3.774 ± 0.00022	6.512 ± 0.00033	5.27 ± 0.06	3.04 ± 0.08	2011.3	2
S7-20	-3.683 ± 0.00025	6.956 ± 0.00036	5.10 ± 0.05	2.40 ± 0.09	2011.2	2
S7-228	-7.756 ± 0.00018	1.721 ± 0.00055	2.68 ± 0.04	2.67 ± 0.15	2015.5	3
S8-15	-1.611 ± 0.00026	8.027 ± 0.00033	-2.91 ± 0.05	-3.12 ± 0.09	2011.3	2
S8-7	-3.673 ± 0.00019	-7.422 ± 0.00025	4.55 ± 0.05	-0.44 ± 0.07	2011.1	2
S8-181	-7.660 ± 0.00020	-3.613 ± 0.00029	-1.41 ± 0.05	-3.47 ± 0.06	2015.5	3
S9-20	4.315 ± 0.00036	-8.032 ± 0.00023	2.35 ± 0.13	0.84 ± 0.07	2011.7	2
S9-23	-1.285 ± 0.00037	9.134 ± 0.00052	-2.64 ± 0.11	-2.84 ± 0.17	2011.6	2
S9-13	-3.012 ± 0.00022	8.823 ± 0.00040	2.67 ± 0.05	3.27 ± 0.09	2011.0	2

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Table 3.4 – continued from previous page

Name	x0 (")	y0 (")	vx (mas yr ⁻¹)	vy (mas yr ⁻¹)	t0	ref ^a
S9-1	9.444 ± 0.00017	0.269 ± 0.00018	-2.55 ± 0.03	-2.99 ± 0.09	2010.4	2
S9-114	-6.519 ± 0.00016	-6.882 ± 0.00024	2.54 ± 0.03	1.64 ± 0.04	2015.5	3
S9-283	-9.622 ± 0.00064	-2.555 ± 0.00030	1.91 ± 0.19	-1.14 ± 0.07	2015.5	3
S9-9	5.649 ± 0.00016	-8.192 ± 0.00018	-1.11 ± 0.04	-2.00 ± 0.06	2010.7	2
S10-136	-8.665 ± 0.00037	-5.268 ± 0.00046	-1.79 ± 0.10	4.04 ± 0.12	2015.5	3
S10-5	-1.578 ± 0.00030	10.027 ± 0.00040	-1.17 ± 0.07	-1.73 ± 0.10	2011.3	2
S10-7	9.709 ± 0.00016	4.409 ± 0.00037	-0.56 ± 0.04	-4.34 ± 0.12	2010.8	2
S11-5	1.367 ± 0.00023	11.695 ± 0.00029	-0.52 ± 0.04	2.22 ± 0.07	2011.4	2
S13-3	11.902 ± 0.00012	5.936 ± 0.00029	1.09 ± 0.02	2.93 ± 0.09	2011.4	2

^a Reference for final proper motion: 1 - central 10 2 - maser mosaic, 3 - HST;

Photometry and Extinction

To assure our final sample shares a common photometric system, we adopt the Kp magnitude for each star from the deep wide mosaic image analysis reported in [63] which covers all 88 stars in our sample. The Kp magnitude for each star is reported in Kp column in Table 3.3. Then we applied the latest extinction map from [95] to differentially de-redden all stars to a common $A_{Ks} = 2.7$, and the extinction corrected magnitude is reported in Kp_{ext} in Table 3.3.

3.2 Orbit Analysis

For stars with measured $(x_0, y_0, v_x, v_y, v_z, a_R)$, the 6 Keplerian orbit parameters (inclination i , angle to the ascending node Ω , time of periaapse passage T_0 , longitude of periaapse ω , period P , and eccentricity e) can be analytically determined if the central potential is known [27]. Section 3.2 describes the Monte Carlo (MC) process used to estimate stars' orbital parameters given prior estimates on the central potential. For stars that show time evolving RVs (S0-1, S0-2, S0-3, S0-8, S0-15, S0-19, S1-2), their orbits are better constrained by simultaneously fitting the astrometry and RV measurements as a function of time [4]. Section 3.2 describes the more computationally expensive orbit fitting procedure used for these 7 stars.

MC analysis for stars with $(x_0, y_0, v_x, v_y, v_z, a_R)$

Orbital parameters are determined for 81 stars in our sample using a MC analysis as described in details in [27] and [39]. These stars have radial velocity v_z measured from Section 3.1 and proper motion x_0, y_0, v_x, v_y measured from Section 3.1. The absolute value of line-of-site

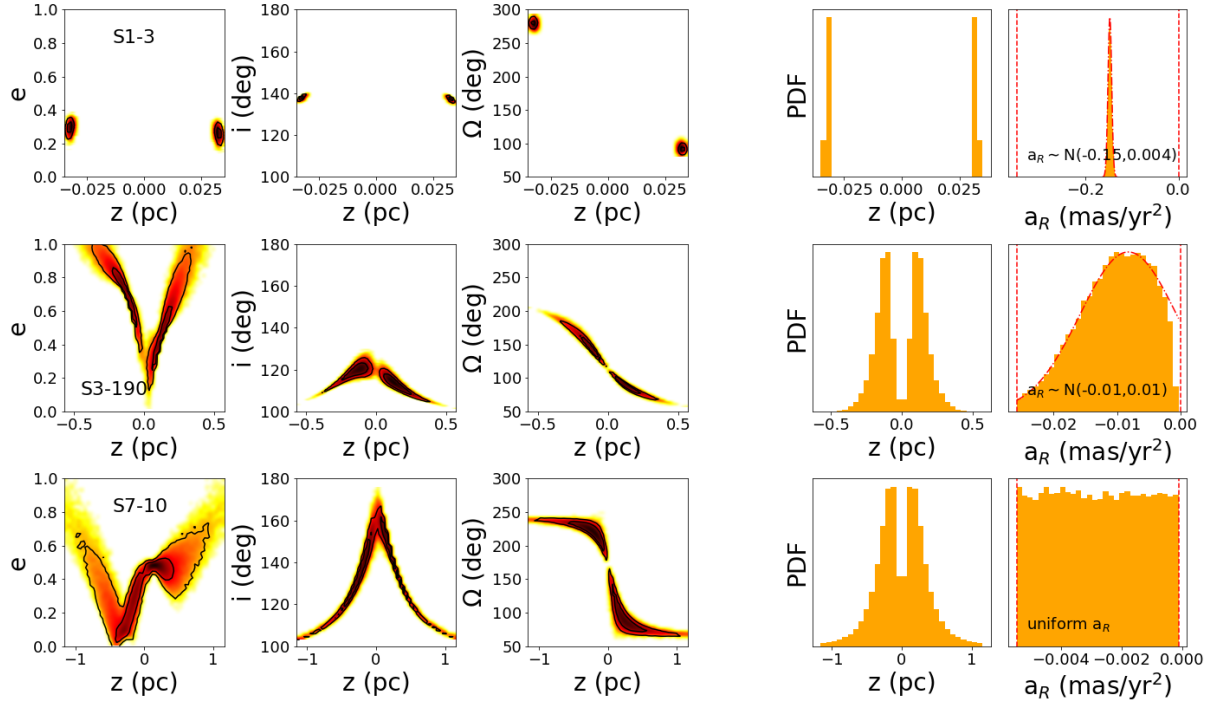


Figure 3.4: Examples of MC analysis for S1-3 (top), S3-190 (middle) and S7-10 (bottom). The left three columns shows the density map for orbital parameters: e , i , Ω . The right two columns shows the MC simulation input prior z and a_R in orange histogram. For S1-3 and S3-190, their measured a_R and σ_{a_R} (red dash line) are within the upper limit and lower limit (vertical red dash line), so the simulated a_R are draw from Gaussian distribution $N \sim (a_R, \sigma_{a_R})$. For S7-10, no a_R is measured, so it is drawn from a uniform distribution between a_{Rmin} and a_{Rmax} . Simulated z is then derived from simulated a_R using Equation 3.3.

distance $|z|$ can be calculated from the following equation if a_R is known.

$$a_R = -\frac{GM_{tot}(r)R}{r^3} \quad (3.3)$$

Here, r is the 3D distance and R is the 2D projected distance from Sgr A*, where $r^2 = R^2 + z^2$. We note that there is a sign ambiguity in the line-of-sight distance.

We follow a similar MC simulation to derive orbital parameters as described in [27] and [39]. We sample those 6 kinematic parameters 10^5 times from their measured values and uncertainties with an Gaussian distribution. The probability density functions (PDFs) for 6 orbital parameters are then constructed based on those artificial data sets.

The mass distribution giving rise to the central potential, M_{tot} , is a combination of the

SMBH mass and an extended mass profile:

$$M_{tot}(r) = M_{BH} + M_{ext}(r) \quad (3.4)$$

The adopted SMBH properties include a mass of $M_{BH} = (3.975 \pm 0.058) \times 10^6 M_\odot$ and a distance of $R_0 = (7.959 \pm 0.059) kpc$, based on the analysis of S0-2's orbit [4]. We used the extended mass profiles from [100] and found that adding extended mass has limited impact on the orbit analysis. We nonetheless adopt an extended mass profile with

$$M_{ext}(r) = \int_0^r 4\pi \frac{\rho_0}{1 + (r/R_b)^2} r^2 dr \quad (3.5)$$

where $\rho_0 \simeq 2.1 \times 10^6 M_\odot$ and the break radius is $R_b = 8.9''$. We have also tried other extended mass profiles, like [101], and the results are similar.

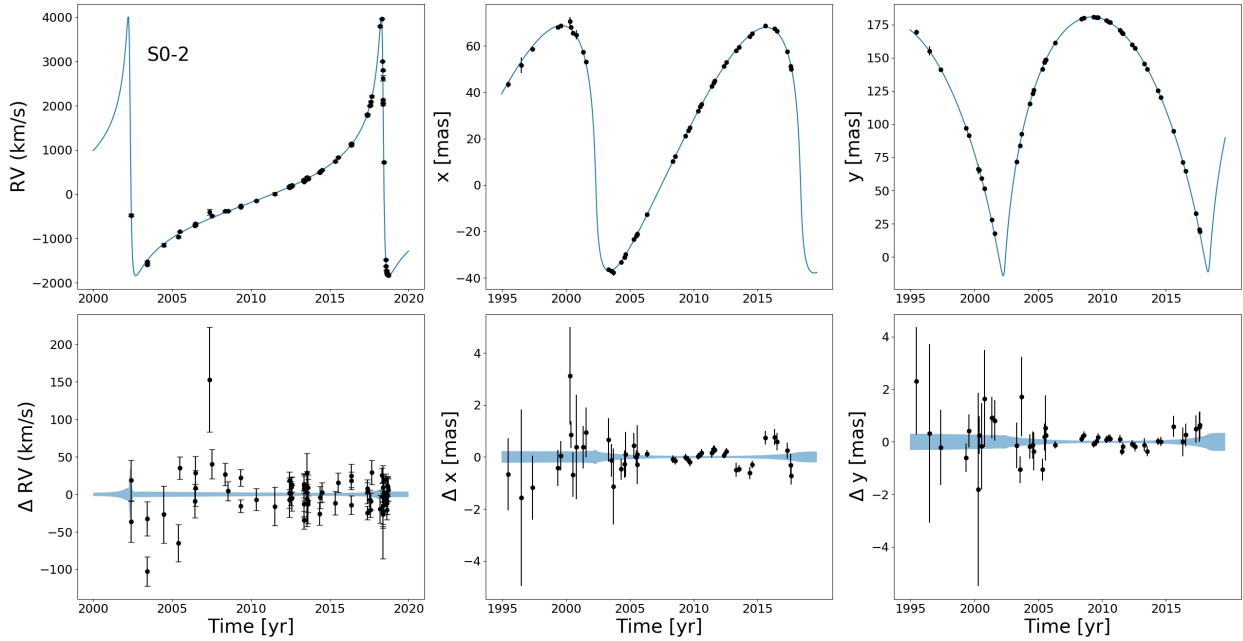


Figure 3.5: Examples of a full Keplerian orbit analysis for S0-2. The top row shows the line-of-sight velocity v_z and projected position from SgrA x , y as a function of time, and our model in blue line fit both very well. The bottom row shows the offset between data and model as a function of time.

Among the 81 stars, 47 stars have precise and accurate astrometry from [94] with well-measured proper motions and constraints on the projected acceleration a_R as shown on Table 3.3. However, not all measured a_R values and uncertainties are physically allowed.

The physically allowed a_R range corresponds to a_R at $z = 0$ pc and $z = 0.8$ oc:

$$\begin{aligned} a_{Rmax}(R) &= -\frac{GM_{tot}(R)}{R^3} \\ a_{Rmin}(R) &= -\frac{GM_{tot}(r)R}{r^3}, r = \sqrt{R^2 + z^2}, z = 0.8pc \end{aligned} \quad (3.6)$$

The maximum z is chosen to be 0.8 pc because very few young stars are detected outside 0.4 pc and the most distant young star detected in our sampler is at 0.6 pc. Assuming a relative symmetric profile, then the maximum line-of-sight distance should not exceed 0.8 pc. If a star's measured a_R ($\pm 2\sigma_{a_R}$) overlaps with the allowed range, we will draw its z in each MC trial from a Gaussian distribution with the mean set to the measured a_R , the standard deviation set to the measured σ_{a_R} , and truncated to the allowed a_R range. This is the case for 37 stars. For the rest of stars without a significant and physical a_R , we use a uniform distribution in the range of allowed a_R . In Figure 3.4, we use three stars as an example to show how a_R is simulated in the MC analysis. The last column in that plot shows the distribution of a_R and its limit (a_{Rmax} and a_{Rmin}) in red dash line, and the orange histogram is the simulated a_R .

Even for those 37 stars with measured a_R within the allowed range, the significance of their accelerations varies. This is due to many factors, including the location of the star, the brightness of the star, the number of epochs the star is detected etc. A more precise acceleration will result in more precise orbital parameters. In Figure 3.4, we show the MC analysis results for 3 stars: S1-3, S3-190 and S7-10. Both S1-3 and S3-190 have a measurement-based a_R prior, while S7-10's prior is a uniform a_R over the allowed range. S1-3 has a $\sim 40\sigma$ significant acceleration, but S3-190 only has 1σ significant acceleration. As a result, S1-3 has a much more constrained orbital parameters compared to S3-190. For S7-10 with a uniform a_R prior, its orbital parameters are even less constrained. The z distribution for S7-10 is decreasing with radius even when a_R is evenly distributed, which agrees with observation and validates the uniform a_R prior.

Full Keplerain orbit fit for central stars

For stars with time-variant RV, we infer their orbital parameters by simultaneously fitting spectroscopic and astrometric measurements. S0-2 is fit in this way to test General Relativity [4]. We also drew 10^5 chains of estimated orbital parameters from the posterior distributions, so later we can easily combine stars fitted with single RV from Section 3.2 and stars fitted with multiple RVs from this section. S0-1, S0-2, S0-3, S0-8, S0-15, S0-19, S1-2 in our sample are fitted this way. We adopted the new observable-based prior paradigm from [102] that is based on uniformity in observables to improve orbital solutions for low-phase-coverage orbits. A full Keplerian orbit fit for S0-2 is shown in Fig 3.5. All 7 stars' fitting result are attached in appendix A.

3.3 Disk Membership Analysis

Detecting Stellar Disks

Each star's orbital plane can be uniquely described by a unit normal vector $\hat{L}$ that is perpendicular to the orbital plane. This normal vector $\hat{L}$ can be expressed by the inclination (i) and the angle to the ascending node (Ω) (see Equation 8 in [27]). Stars on a common plane share a common normal vector. In order to detect the stellar disk, we follow a similar nearest neighbor method as used in [27] and [39]. In this method, the sky is divided into 49152 pixels with equal solid angle area and for a given MC simulation the density at each (i, Ω) position is calculated using the following equation:

$$\Sigma = \frac{k}{2\pi(1 - \cos \theta_k)} \text{ stars sr}^{-1} \quad (3.7)$$

where θ_k is the angle to the k th nearest star and we use $k = 6$. Then the combined density map is an average over all 10^5 MC trials from Section 3.2. The resulting density maps are presented in Section 3.4.

Disk Membership Probabilities

With the disk area measured from Section 3.3, the disk membership probability L_{disk} can be estimated following [27]:

$$\begin{aligned} L_{disk} &= 1 - L_{non-disk} \\ &= 1 - \left(\frac{\int_{disk} PDF(i, \Omega) dSA}{\int_{peak} PDF(i, \Omega) dSA} \right) \\ &= \frac{\int_{disk} PDF(i, \Omega) dSA}{\int_{peak} PDF(i, \Omega) dSA} \end{aligned} \quad (3.8)$$

$$\int_{disk} dSA = \int_{peak} dSA \quad (3.9)$$

where SA is the solid angle measured at the counter when the disk peak of the combined density map drops to 50%, $\int_{disk} PDF(i, \Omega) dSA$ is the integration of each star's density map over the stellar disk region, and $\int_{peak} PDF(i, \Omega) dSA$ is the integration over its own peak region with the same SA. The final disk membership is shown in Section 3.4.

3.4 Results

Two Stellar Disks

The density map of normal vector $\hat{L}$ (Figure 3.6) shows two over-dense regions, indicating the presence of at least two distinct populations that share a common orbital plane. We

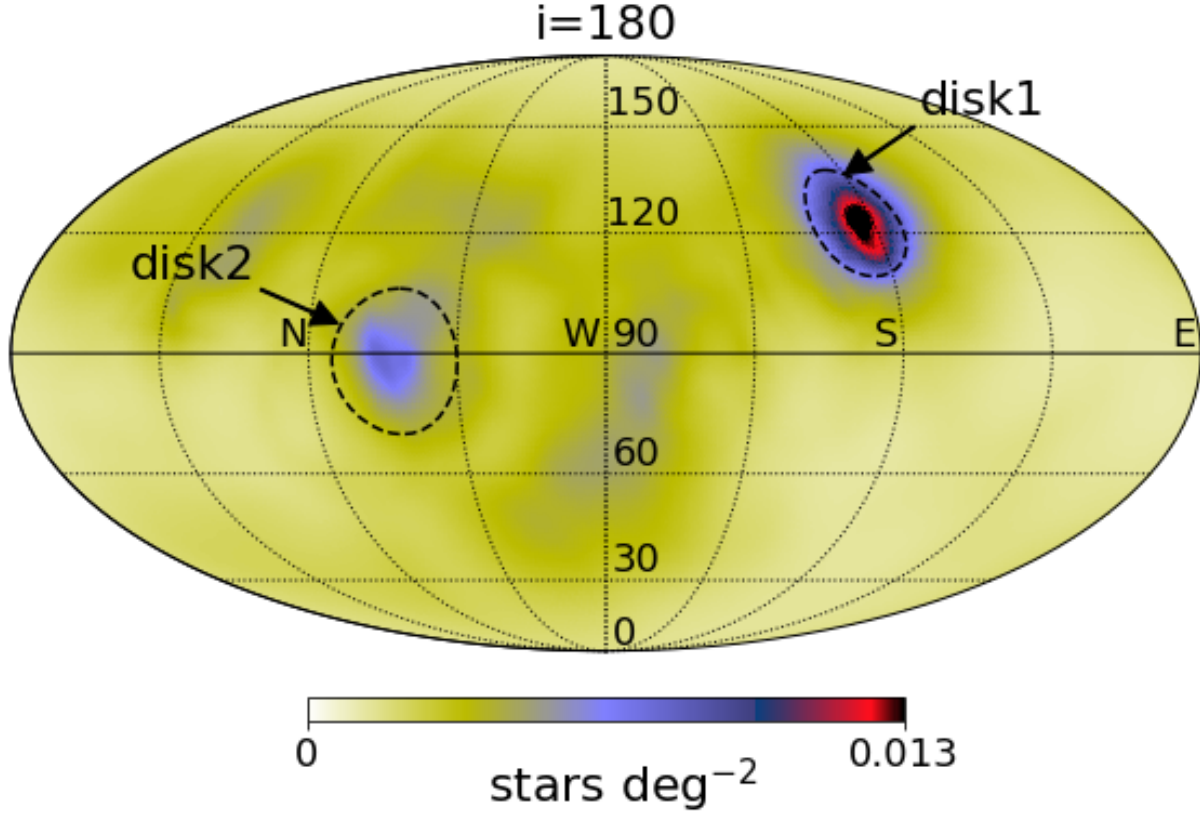


Figure 3.6: Density of normal vectors (i, Ω) (in stars deg^{-2}) of all 88 stars in our sample. Two significant over-dense region are marked as disk1 and disk2. Disk1 has a density of 0.014 stars deg^{-2} located at $(i, \Omega) = (123^\circ, 96^\circ)$. Disk2 has a density of 0.007 stars deg^{-2} located at $(i, \Omega) = (88^\circ, 244^\circ)$. The 1 sigma region is defined when the density drops to half of its peak, and is circled in black dash line.

Table 3.5: Disk Summary

Disk	i ($^\circ$)	Ω ($^\circ$)	Solid Angle (sr)	ρ_{disk} (stars/deg 2)	ρ_{iso} (stars/deg 2)	Significance
disk1	123 ± 14	95 ± 15	0.16	0.014 ± 0.009	0.003 ± 0.002	6
disk2	88 ± 18	244 ± 19	0.35	0.007 ± 0.002	0.002 ± 0.002	2

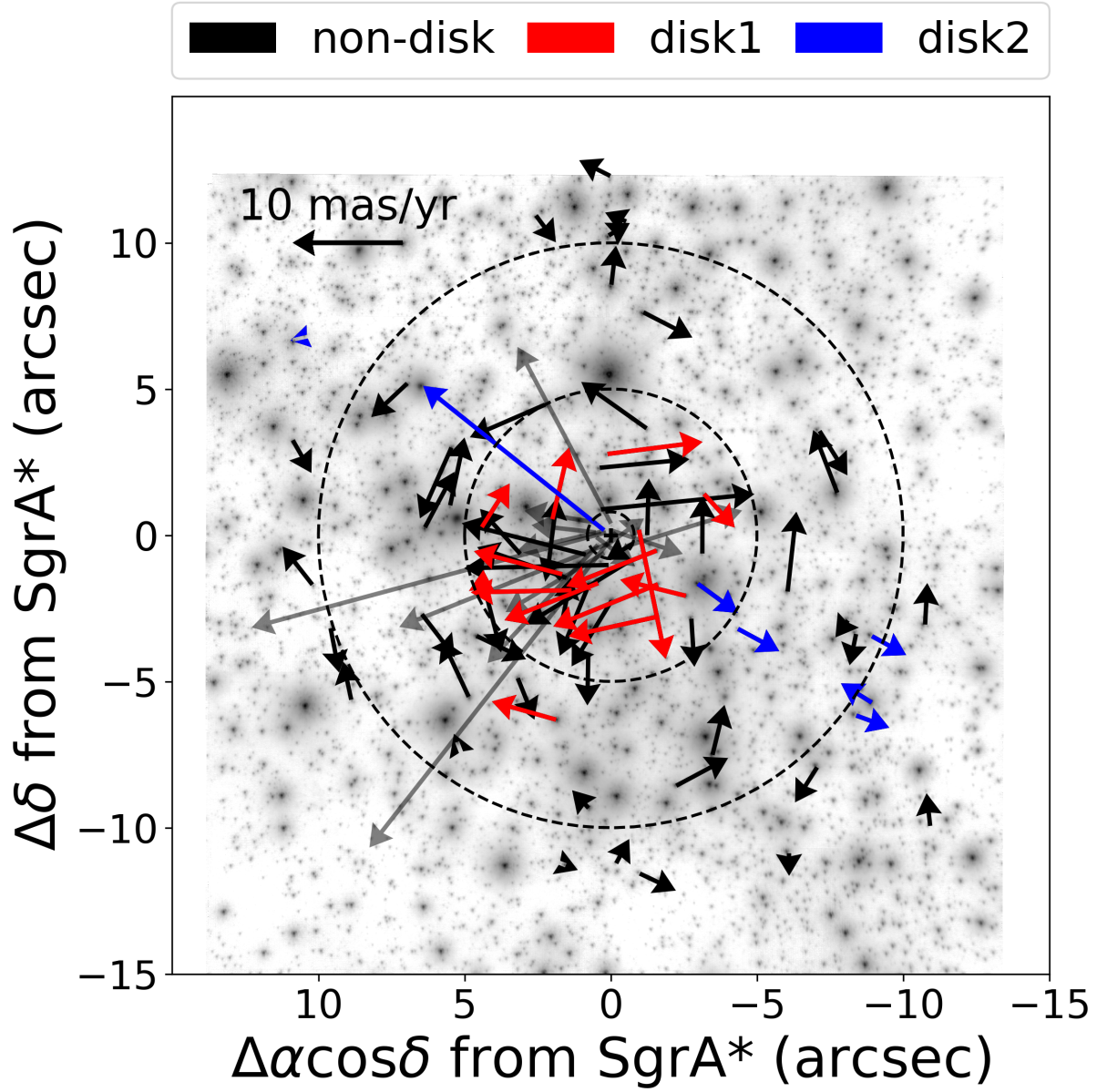


Figure 3.7: The proper motion direction for each star, where red arrows are disk1 candidates and blue arrows are disk2 candidates. The non-disk stars are shown in black arrows, and the inner region non-disk stars are plotted with low opacity due to their large velocity.

label the two peaks as disk1 and disk2. The well-known clockwise (CW) disk is the upper right disk1 located at $(i_{disk1}, \Omega_{i,disk1}) = (123^\circ, 95^\circ)$, consistent with the past measurements of the disk location [25, 103, 16, 27, 18, 39]. We found another almost edge-on disk which we call disk2, located at $(i_{disk2}, \Omega_{disk2}) = (88^\circ, 244^\circ)$. The uncertainty is defined when the intensity drops to 50% of its peak intensity and it is marked in black dash line in Figure 3.6. The uncertainties for *disk1* and *disk2* are $(\sigma_{i,disk1}, \sigma_{\Omega,disk1}) = (15^\circ, 14^\circ)$, $(\sigma_{i,disk2}, \sigma_{\Omega,disk2}) = (18^\circ, 19^\circ)$ separately.

To quantify the significance of both disks, we simulated an isotropic population with syntheic (x,y,vx,vy,vz) and extract orbits in the same way as on the real data. Each simulated star was assigned a position and velocity with random direction, with unchanged R and v_{tot} values [39]. For 6 stars from Section 3.2, because their orbital parameters are so well measured, we just drew 10^5 MC trials from their posterior distribution. The significance of disk1 and disk2 is defined by:

$$S = \frac{\rho_{disk} - \rho_{iso}}{\sigma_{iso}} \quad (3.10)$$

where ρ_{disk} is the density of the stellar disk at its peak, ρ_{iso} is the density of the isotropic simulation at the same position, and σ_{iso} is the uncertainty for isotropic simulation at the same position. A summary of both disks' peak density, significance, and the simulated isotropic density is presented in Table 3.5. The significance for disk1 is 6 and disk2 has a significance of 2. The reason we have a lower significance for disk1 compared to [39] is because we have a smaller sample size due to our requirement that only stars from surveys with published completeness curves can be included. As a result, our uncertainty in the isotropic simulation is 2 times larger.

We use Equation 3.8 to calculate the probability of each star laying on disk1 and disk2, as shown in Table 3.6. To view stars in disk1 and disk2 in a more direct way, we make a quiver plot showing the proper motion of each star in Figure 3.7. For illustrative purposes, we identify high-probability disk members as those with $P_{disk} > 0.2$; and disk1 and disk2 stars are shown in red and blue arrows, respectively.

Table 3.6: Disk Membership

Name	P _{disk1}	P _{disk2}	SA
S0-1	0.000	0.000	0.001
S0-11	0.000	0.050	0.081
S0-14	0.000	0.000	0.004
S0-15	0.950	0.000	0.005
S0-16	0.000	0.410	0.024
S0-19	0.000	0.000	0.002
S0-2	0.000	0.000	0.000
S0-20	0.000	0.000	0.049
S0-3	0.000	0.000	0.000

S0-30	0.000	0.000	0.106
S0-31	0.000	0.110	0.004
S0-4	0.000	0.000	0.004
S0-5	0.000	0.000	0.003
S0-7	0.000	0.000	0.001
S0-8	0.000	0.000	0.001
S0-9	0.000	0.000	0.003
S1-19	0.420	0.000	0.022
S1-2	0.100	0.000	0.012
S1-22	0.560	0.000	0.189
S1-24	0.000	0.000	0.003
S1-3	0.190	0.000	0.007
S1-33	0.000	0.000	0.122
S1-4	0.020	0.000	0.081
S1-8	0.000	0.000	0.003
S10-185	0.000	0.740	0.132
S10-232	0.000	0.000	0.578
S10-238	0.000	0.780	0.002
S10-32	0.150	0.000	0.119
S10-34	0.000	0.000	0.517
S10-4	0.000	0.000	0.087
S10-48	0.000	0.000	0.037
S10-50	0.000	0.000	0.400
S11-147	0.000	0.000	1.322
S11-176	0.000	0.000	0.237
S11-21	0.020	0.000	0.174
S11-214	0.000	0.000	0.131
S11-246	0.000	0.010	0.209
S11-8	0.000	0.000	0.784
S12-178	0.000	0.030	0.008
S12-5	0.000	1.000	0.003
S12-76	0.000	0.000	0.257
S14-196	0.000	0.000	0.410
S2-17	0.350	0.000	0.009
S2-19	0.190	0.000	0.071
S2-21	0.490	0.000	0.018
S2-4	0.160	0.000	0.007
S2-50	0.000	0.000	0.188
S2-58	0.000	0.000	0.188
S2-6	0.500	0.000	0.009
S2-74	0.500	0.000	0.044
S3-19	0.380	0.000	0.357

S3-190	0.340	0.000	0.084
S3-26	0.220	0.170	0.163
S3-3	0.000	0.000	0.467
S3-30	0.000	0.000	0.012
S3-331	0.000	0.000	0.103
S3-374	0.000	0.000	0.086
S3-96	0.000	0.000	0.351
S4-169	0.360	0.000	0.420
S4-262	0.000	0.000	0.117
S4-314	0.460	0.000	0.231
S4-364	0.000	0.000	0.296
S4-71	0.000	0.000	0.012
S5-106	0.000	0.340	0.029
S5-183	0.000	0.000	0.094
S5-191	0.000	0.000	0.612
S5-237	0.000	0.000	0.247
S6-63	0.280	0.000	0.215
S6-81	0.000	0.000	0.086
S6-89	0.000	0.000	0.398
S6-96	0.000	0.000	0.411
S7-10	0.100	0.030	0.285
S7-216	0.000	0.000	0.153
S7-236	0.000	0.000	0.873
S7-30	0.000	0.000	0.592
S7-5	0.030	0.000	0.499
S8-10	0.000	0.000	0.769
S8-126	0.000	0.000	1.139
S8-196	0.000	0.200	0.142
S8-4	0.000	0.000	0.032
S8-5	0.000	0.000	0.370
S8-8	0.010	0.000	0.271
S9-143	0.010	0.000	1.220
S9-221	0.000	0.760	0.044
S9-6	0.040	0.000	0.267
IRS 13E1	0.000	0.500	0.002
IRS 16CC	0.210	0.000	0.070
IRS 33N	0.000	0.000	0.004

Disk Properties

In this section, we are going to compare the distribution of eccentricity e , semi-major axis a (derived from period P) and disk thickness among different dynamical subgroups, and later use this distribution as the input prior for disk simulations in Section 3.4.

The way stars are divided into disk1, disk2 and non-disk stars are not by a hard likelihood cut. Instead we assign a weight to each MC trial among the 10^5 trials for every star (Section 3.2). For a particular MC trial with a set of Keplerian orbital parameters (i, Ω, e, a) , the weight of being on disk1 (W_{disk1}), disk2 (W_{disk2}) and non-disk ($W_{\text{non-disk}}$) are calculated by the following equation:

$$\begin{aligned} W_{\text{disk1}}(i, \Omega, e, a) &= N(i|\mu = i_{\text{disk1}}, \sigma = \sigma_{i,\text{disk1}}) \times \\ &\quad N(\Omega|\mu = \Omega_{\text{disk1}}, \sigma = \sigma_{\Omega,\text{disk1}}) \\ W_{\text{disk2}}(i, \Omega, e, a) &= N(i|\mu = i_{\text{disk2}}, \sigma = \sigma_{i,\text{disk2}}) \times \\ &\quad N(\Omega|\mu = \Omega_{\text{disk2}}, \sigma = \sigma_{\Omega,\text{disk2}}) \\ W_{\text{non-disk}}(i, \Omega, e, a) &= 1 - W_{\text{disk1}}(i, \Omega, e, a) \\ &\quad - W_{\text{disk2}}(i, \Omega, e, a) \end{aligned} \tag{3.11}$$

where $N(\mu, \sigma)$ is normal distribution with mean of μ and standard deviation of σ .

The weighted histogram of eccentricity e is plotted in Figure 3.8. For disk1 population, we see a peak at $e = 0.2$, which is consistent with previous calculation [39], but with larger uncertainty $\sigma_e = 0.2$, implying more high eccentric orbits in disk1. For disk2 population, very interestingly the probability of e is almost proportional to e itself. This is typical for a relaxed population, but based on how young those stars are, we think it is just a reflection of really high eccentricities. For non-disk population, the distribution is nearly flat between 0 and 1, a reasonable distribution for random population in the midst of relaxation. To summarize, we find that the eccentricity distributions are described by:

$$\begin{aligned} P_{\text{disk1}}(e) &\propto N(\mu = 0.2, \sigma = 0.2) & e \in [0, 1] \\ P_{\text{disk2}}(e) &\propto e & e \in [0, 1] \\ P_{\text{non-disk}}(e) &\propto \text{Uniform}(0, 1) \end{aligned} \tag{3.12}$$

The weighted histogram of semi-major axis a is plotted in Figure 3.9. At larger radius ($a > 0.3$ pc), the distribution of a is consistent between disk1, disk2 and non-disk within uncertainties. They all follow a power law function with a slope of -2 at larger a , which agrees with the observed surface-density profile [54, 16]. However at inner radius region, disk1 population density drops a lot faster compared to disk2 and non-disk population. See details in Equation 3.13.

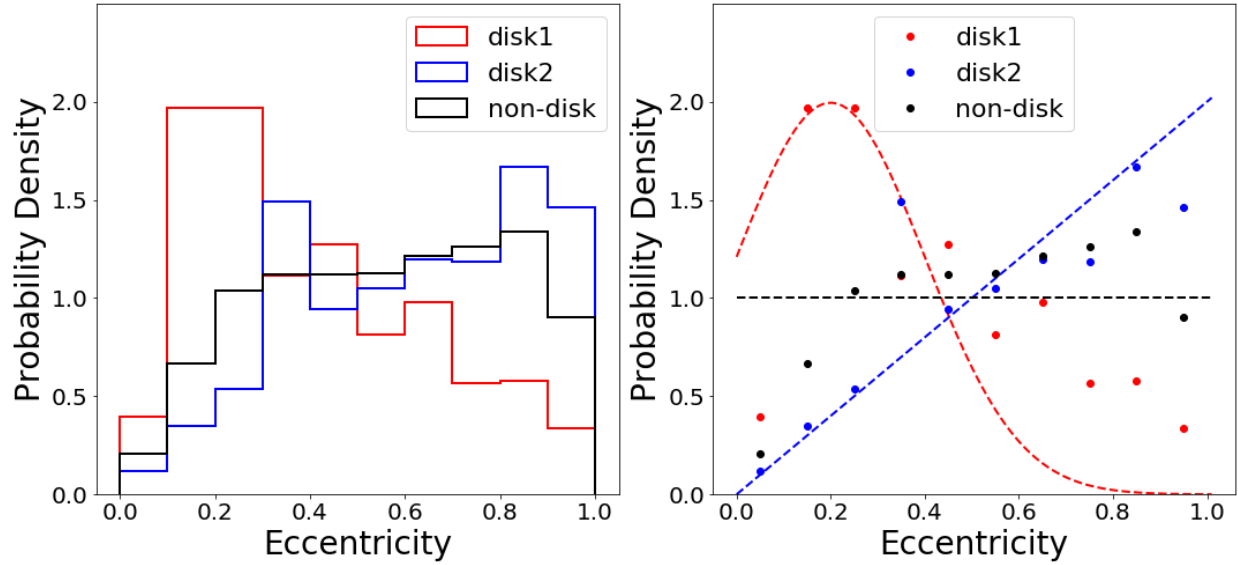


Figure 3.8: Left: The weighted histogram of eccentricity for disk1 (red), disk2 (blue) and non-disk (black) stars, with weighted calculated from Equ 3.11. Right: The approximate fit for their probability density function - disk1 population is fit by a Gaussian distribution with mean of 0.2 and std of 0.2; disk2 population is fit by a linear distribution $P(e) \propto e$; non-disk population is fit by a flat distribution between 0 and 1.

$$\begin{aligned}
 P_{\text{disk1}}(a) &\propto \begin{cases} a^{2.3} & 0.005\text{pc} \leq a < 0.12\text{pc} \\ a^{-2} & 0.12\text{pc} \leq a < 300\text{pc} \end{cases} \\
 P_{\text{disk2}}(a) &\propto \begin{cases} a^{-0.2} & 0.005\text{pc} \leq a < 0.38\text{pc} \\ a^{-2} & 0.38\text{pc} \leq a < 300\text{pc} \end{cases} \\
 P_{\text{non-disk}}(a) &\propto \begin{cases} a^{-0.35} & 0.005\text{pc} \leq a < 0.33\text{pc} \\ a^{-2} & 0.33\text{pc} \leq a < 300\text{pc} \end{cases}
 \end{aligned} \tag{3.13}$$

The thickness of the disk can be estimated using the velocity dispersion perpendicular to the disk plane. Here we follow a similar process as described in [27]. First each potential disk candidate's three dimensional velocity $\vec{v}$ is projected into the direction of $\hat{L}$. The uncertainty of both $\vec{v}$ and $\hat{L}$ are taken into consideration when calculating intrinsic velocity dispersion σ_{v_n} , where $v_n = \vec{v} \cdot \hat{L}$. Then the disk's scale height (h/r) can be derived from the ratio of σ_{v_n} and $\langle \vec{v} \rangle$. Finally this scale height can be related to disk thickness described in dispersion angle $h/r \propto \sqrt{1/2\Delta\theta}$. For disk1, we find $\sigma_{v_n} = 22 \text{ km s}^{-1}$, giving a scale height of 0.05 ± 0.01 and a dispersion angle of $4^\circ \pm 1^\circ$, slighter smaller than previous results [27, 39]. For

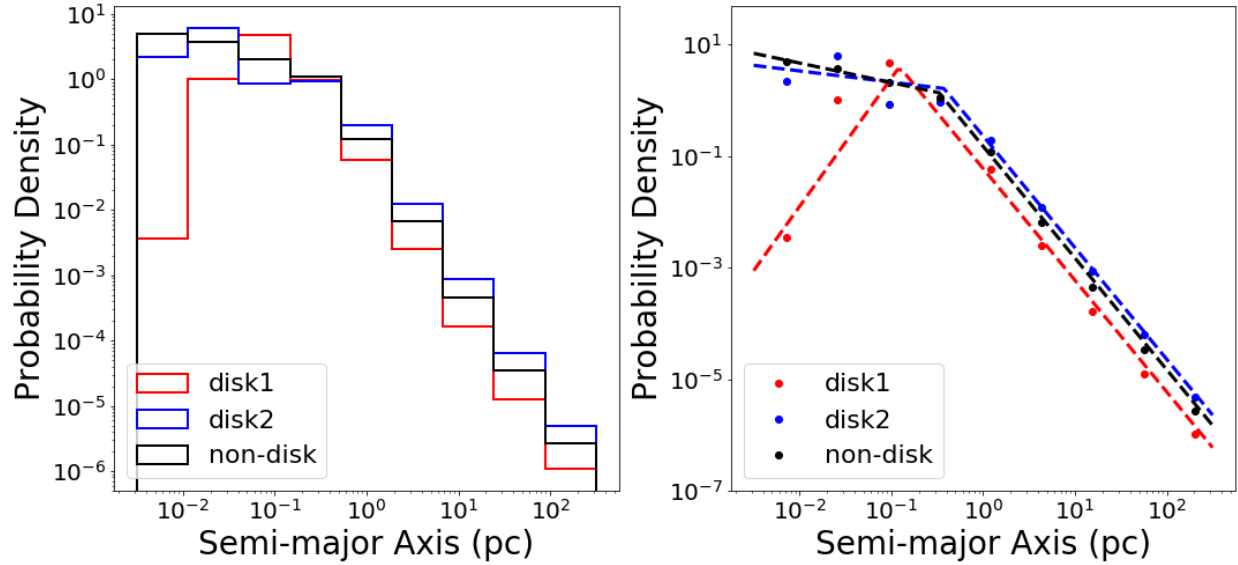


Figure 3.9: The weighted histogram of semi-major axis for disk1, disk2 and non-disk stars. They all follow a power-law function with similar slope at larger radius: a slope of -2, but the inner region slope slight varies.

disk2, we find $\sigma_{v_n} = 50 \text{ km s}^{-1}$, giving a scale height of 0.2 ± 0.08 and a dispersion angle of $16^\circ \pm 6^\circ$.

Disk Simulation

From Figure 3.6, we can see the spatial distribution within each disk plane does not appear symmetric, especially for disk2 stars. In order to understand where this asymmetry comes from, we need to simulate each dynamical sub-groups using properties from Section 3.4.

Table 3.7 summarizes the input simulation parameters. First, we use an open-source python package *PyPopStar* (Hosek et al. 2020, submitted) to generate a single age (3.2 Myr) star cluster with solar metallicity, located 8 kpc away. The total mass of the cluster is $10^5 M_\odot$, with minimum mass of $1 M_\odot$ and maximum mass of $100 M_\odot$; however, the simulations will be rescaled to match observed stellar densities. We use a traditional power law IMF with a slope of -2.35 [64]. We generate the synthetic photometry for each star in the Kp filter, assuming a fixed extinction value of $A_{K_s} = 2.7 \text{ mag}$, So this corrected magnitude can be directly compared with Kp_{ext} shown in Table 3.3. In this step, we assume the three dynamical subgroups have the same age, IMF, and extinction.

The cluster generated by *PyPopStar* gives mass and Kp_{ext} for every non-WR star in the system. The next step is to assign a position for each non-WR star based on their dynamical sub-group. We use the disk properties from Section 3.4 to simulate positions for disk1,

Table 3.7: Input for the Simulated Cluster

Parameters	disk1	disk2	non-disk
i	$N \propto (i_{\text{disk1}}, \sigma_{i,\text{disk1}})$	$N \propto (i_{\text{disk2}}, \sigma_{i,\text{disk2}})$	$P(i) \propto \cos(i)$
Ω	$N \propto (\Omega_{\text{disk1}}, \sigma_{\Omega,\text{disk1}})$	$N \propto (\Omega_{\text{disk2}}, \sigma_{\Omega,\text{disk2}})$	Uniform($0^\circ, 360^\circ$)
ω	Uniform($0^\circ, 360^\circ$)	Uniform($0^\circ, 360^\circ$)	Uniform($0^\circ, 360^\circ$)
e	Equ 3.12	Equ 3.12	Equ 3.12
a	Equ 3.13	Equ 3.13	Equ 3.13
t_0	Uniform(1995, 1995+P)	Uniform(1995, 1995+P)	Uniform(1995, 1995+P)
$M_{\min}$	$1M_\odot$	$1M_\odot$	$1M_\odot$
$M_{\max}$	$150M_\odot$	$150M_\odot$	$150M_\odot$
M_{cluster}	$10^5 M_\odot$	$10^5 M_\odot$	$10^5 M_\odot$
Age	3.2Myr	3.2Myr	3.2Myr
IMF	$\xi(m) \propto m^{-2.35}$	$\xi(m) \propto m^{-2.35}$	$\xi(m) \propto m^{-2.35}$
distance	8kpc	8kpc	8kpc

disk2 and non-disk stars. To account for differential extinction over the field of view, which introduces asymmetric features in the observed distributions, we redden the $K_{p_{\text{ext}}}$ back to observed K_p using the extinction map from [95] and then apply the completeness correction from Section 3.1.

The comparison of observed stellar density profile and simulated density profile is shown in Figure 3.10. We presented the density profile in polar coordinates, where North is at 0° and East is at 90° , so that it is easier to see the azimuthal structure in each dynamical subgroup. This system has the same convention as the Keplerian orbital element, Ω .

For the non-disk group, the substructure is mainly caused by the differential extinction, and our simulation reproduces the observation well within uncertainties. This indicates that the non-disk group is nearly isotropic. For disk1 and disk2, we expect an over-density in their disk plane, which can be seen in both the observed and simulated maps. However, for disk2, the Southwest side is much more dense than the Northeast side in the observed distribution, which cannot be explained by differential extinction and incompleteness alone. This indicates that the stars in disk2 are not uniformly distributed within the disk plane. Given the extreme asymmetry, disk2 may actually be a stream.

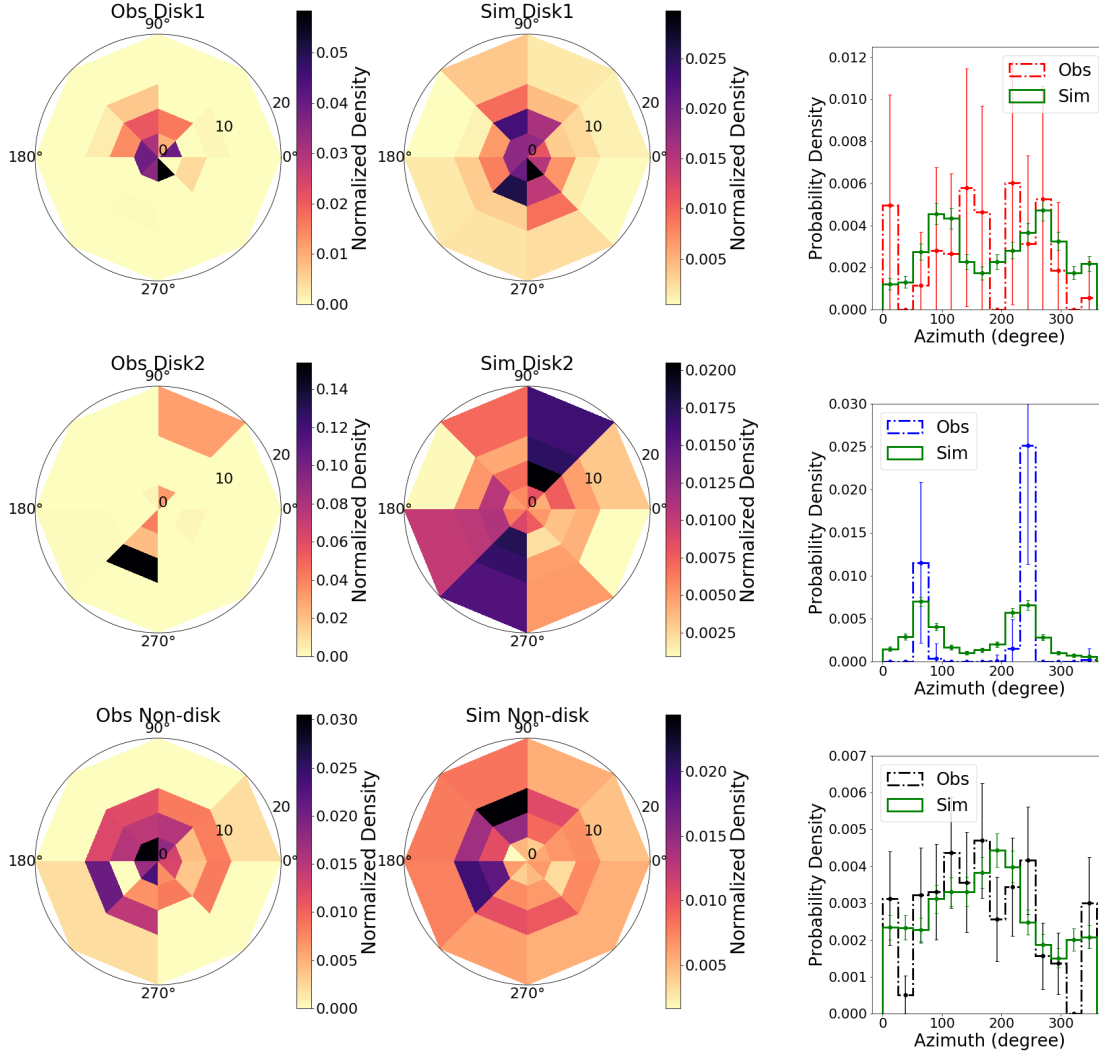


Figure 3.10: Stellar density map comparison between observation and simulation for disk1 (top), disk2 (middle) and non-disk (bottom) subgroups. The first two columns are the stellar density map in polar system for observation (left) and simulation (middle), in which North is 0° and East is 90° . We choose this coordinate system to be consistent with how Ω is defined. The right column is the 1D histogram of azimuth. For disk1 and disk2 group, the over-dense region are mostly from the orientation of the disk, while for non-disk group, the asymmetry is less significant, resulting from the differential reddening map. Within uncertainties, the 1D azimuth distribution is consistent between the observation and simulation for disk1 and non-disk group. However for disk2, the lack of stars in northeast direction might in observation cannot be explained by the simulation itself.

3.5 Discussion

CW Disk1 Compared to Previous work

While the CW disk has been verified many times in previous work [25, 103, 16, 27, 39], other kinematic structures in the young nuclear cluster have been controversial. A counter-clockwise (CCW) disk was originally reported in [103] and confirmed in [16]. Later [18] found that the CCW disk was highly extended, anisotropic, and showed evidence for a warped disk on large scales. However, this CCW disk was not detected by [27] and [39] in an independent analysis with different observations for RVs and proper motions. In our work, we confirmed the existence of CW disk, whose position is consistent with previous analysis, and we did not detect the CCW disk either. Instead, we detect a new edge-on disk called disk2 in our paper, and we will discuss the possibility of disk2 in Section 3.5 in more detail. The comparison of the CW disk, CCW disk and disk2 is presented in Figure 3.11.

The stars within $0.8''$ are more randomized by dynamic effects like vector resonant relaxation [104, 21, 105], so the inner radius cut for the CW disk is roughly at $0.8''$ as reported in [106, 8, 84]. This agrees with our analysis. In our sample, we have 14 stars within $0.8''$ and all of them have zero probability of being on disk1.

The stars in disk1 are found to have non-circular orbit. [18] combined their result with [84] and reported an eccentricity of $\langle e \rangle = 0.36 \pm 0.06$. [39] divided stars into accelerating sources and non-accelerating sources, with $\langle e \rangle = 0.27 \pm 0.09$ and $\langle e \rangle = 0.43 \pm 0.24$ respectively. Our eccentricity distribution is shown in Figure 3.8, with a peak at 0.2 and a much larger uncertainty of 0.2. We did a similar analysis as in [39], dividing stars into accelerating and non-accelerating moving ones. The accelerating sources are defined in [94], including 5 stars in our sample: S1-3, S1-8, S1-4, S2-6 and S4-169. The comparison of e distribution between accelerating and non-accelerating sources in disk1 are plotted in Figure 3.12, which agrees with the conclusion from [39] - accelerating sources have a well constrained e peaking at 0.2, while the e for non-accelerating sources are a lot more uncertain. The reason could be that the accelerating sources have better constrained orbital parameters with well measured a_R (see Figure 3.4).

Disk1 stars have significant intrinsic thickness as been estimated in [16, 27] and [39]. For example, [16] reported the disk thickness $\Delta_\theta = 14^\circ \pm 4^\circ$ and [27] reported the disk thickness $\Delta_\theta = 7^\circ \pm 2^\circ$. Our disk thickness of $\Delta_\theta = 4^\circ \pm 1^\circ$ is slightly smaller compared to previous work.

The surface density profile is predicted to be $\Sigma(r) \propto r^{-2}$ for *in-situ* formation scenario [107, 108] and has been verified in many observations [16, 18, 27, 39]. In our paper, the semi-major axis a also follows a power law function with a slope of -2.

Is Disk2 Related to IRS 13 Group?

IRS 13 is a group of comoving massive young stars, located $3.5''$ away from the SMBH [109, 16, 110]. It has been proposed to lie within the previously claimed CCW disk [109, 111],

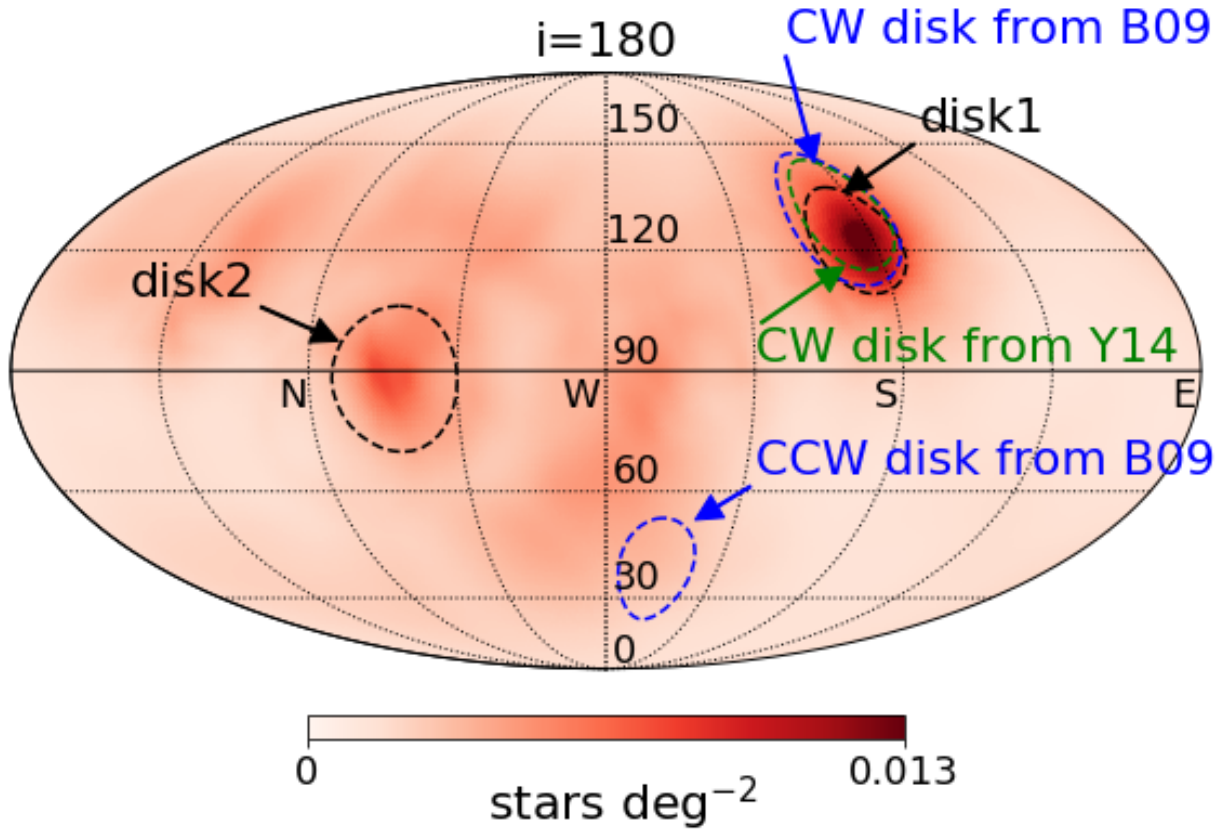


Figure 3.11: Proposed stellar disk from previous work and our work. Here we plot the CW disk and CCW disk from [18] (shortten as B09) in blue and the CW disk from [39] in green. Stellar disk reported in our work in shown in black. We can see that the CW disk is consistent between different analysis.

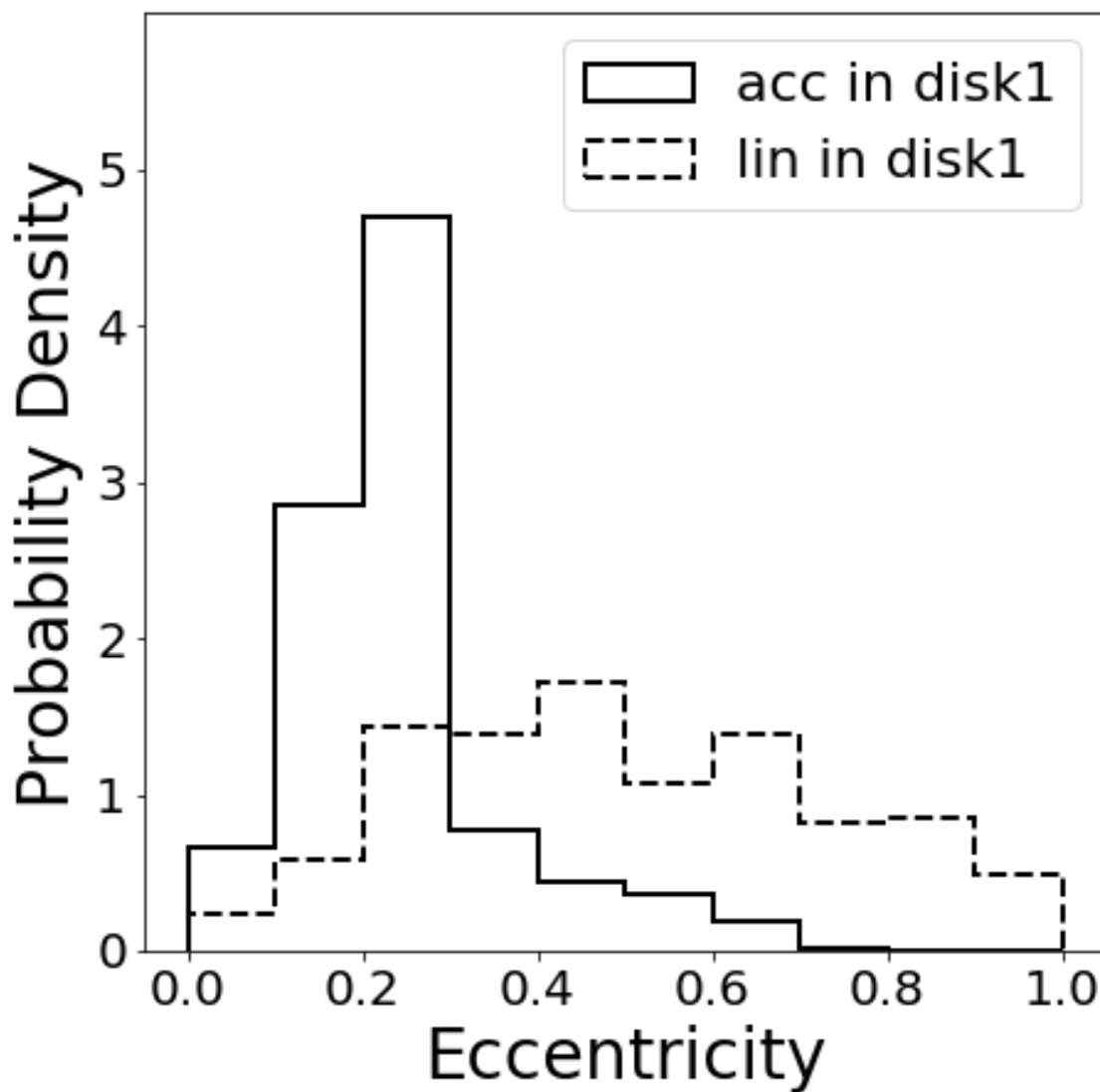


Figure 3.12: The eccentricity distribution for accelerating sources and non-accelerating sources on disk1.

but later this CCW disk is not detected in [27, 39] or this work. Interestingly, one of IRS 13 group star, IRS 13E1 is a potential disk2 star, with a 50% probability being on disk2. The reason that we only included IRS 13E1 in our analysis is because that a known completeness curve is required for a star to be selected in our sample, but other literatures that reported RV for IRS 13 sources did not report their completeness corrections.

Since the IRS 13 cluster is moving in the similar direction, here we run the same orbital and disk membership analysis for other three IRS 13 stars: IRS 13E2, IRS 13E3b, IRS 13E4, as for IRS 13E1. We use the RV reported in [112] for these three IRS 13 sources. It is the most recent RV measurements for IRS 13 group and we note that using RV measurements from [16] and [18] will generate similar results. We found that IRS 13E3b is also a potential disk2 star, with a disk membership probability $P_{\text{disk}} = 0.3$, but IRS 13E2 and IRS 13E4 are not on disk2 with $P_{\text{disk}} = 0$. The proper motions for IRS 13 stars are plotted in Figure 3.13. We can see that IRS 13E1 and IRS 13E3b are moving in almost the same direction, implying that both of them are likely to be on disk2. IRS 13E2 and IRS 13E4 are moving in a slightly different direction. The normal vector $\hat{L}$ density map for IRS 13 group stars are plotted in Figure 3.14.

We limit our analysis to stars with both RVs and completeness corrections so that we could compare simulation and observation. In order to include as many disk2 stars as possible, here we report all potential disk2 stars as long as they have reported RVs and proper motions, resulting in 146 stars in Table 3.2. The potential disk2 stars and their disk membership is reported in Table 3.8. A quiver plot showing the proper motions and RVs for those potential disk2 stars is shown in Figure 3.15.

Table 3.8: Potential Disk2 stars

Name	P_{disk2}	SA
S0-16	0.41	0.024
S10-185	0.74	0.132
S10-238	0.78	0.002
S12-5	1.00	0.003
S13-3	0.27	0.133
S5-106	0.34	0.029
S5-236	0.57	0.024
S5-34	0.37	0.034
S9-114	0.52	0.063
S9-221	0.76	0.044
IRS 13E1	0.50	0.002
IRS 13E3b	0.29	0.265

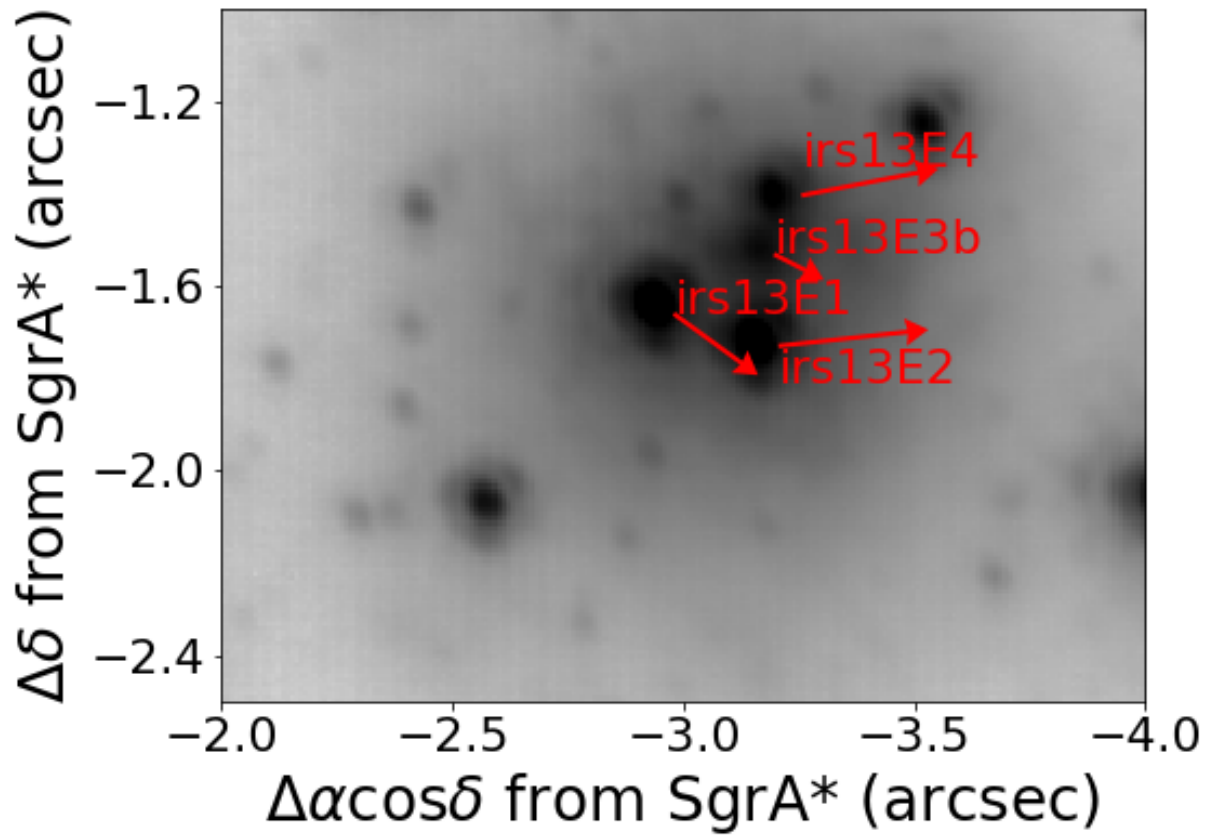


Figure 3.13: The proper motion for IRS 13 sources, where IRS 13E1 and IRS 13E3b are likely to be on disk2 while IRS 13E2 and IRS 13E4 are not.

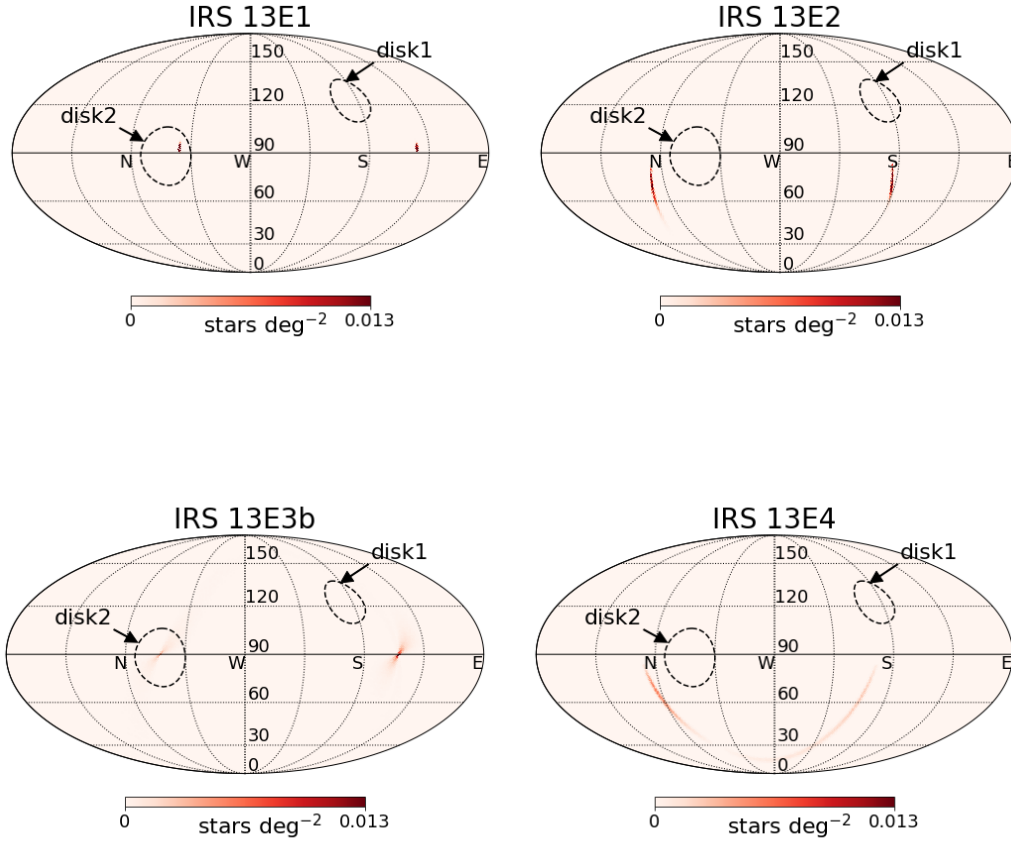


Figure 3.14: The (i, Ω) density map for IRS 13E1, IRS 13E2, IRS 13E3b, IRS 13E4.

Star Formation History

Besides the widely accepted *in-situ* formation theory, people have also used inspiraling star cluster to explain the formation of young stars in our GC, where a massive young cluster migrate towards the center of the galaxy under dynamical friction [113]. However, this scenario will deposit stars with a profile of $r^{-0.75}$, which is inconsistent with the observed density profile. So the fact that the semi-major axis profile from Section 3.4 follows $P(a) \propto a^{-2}$ supports the *in-situ* formation scenario.

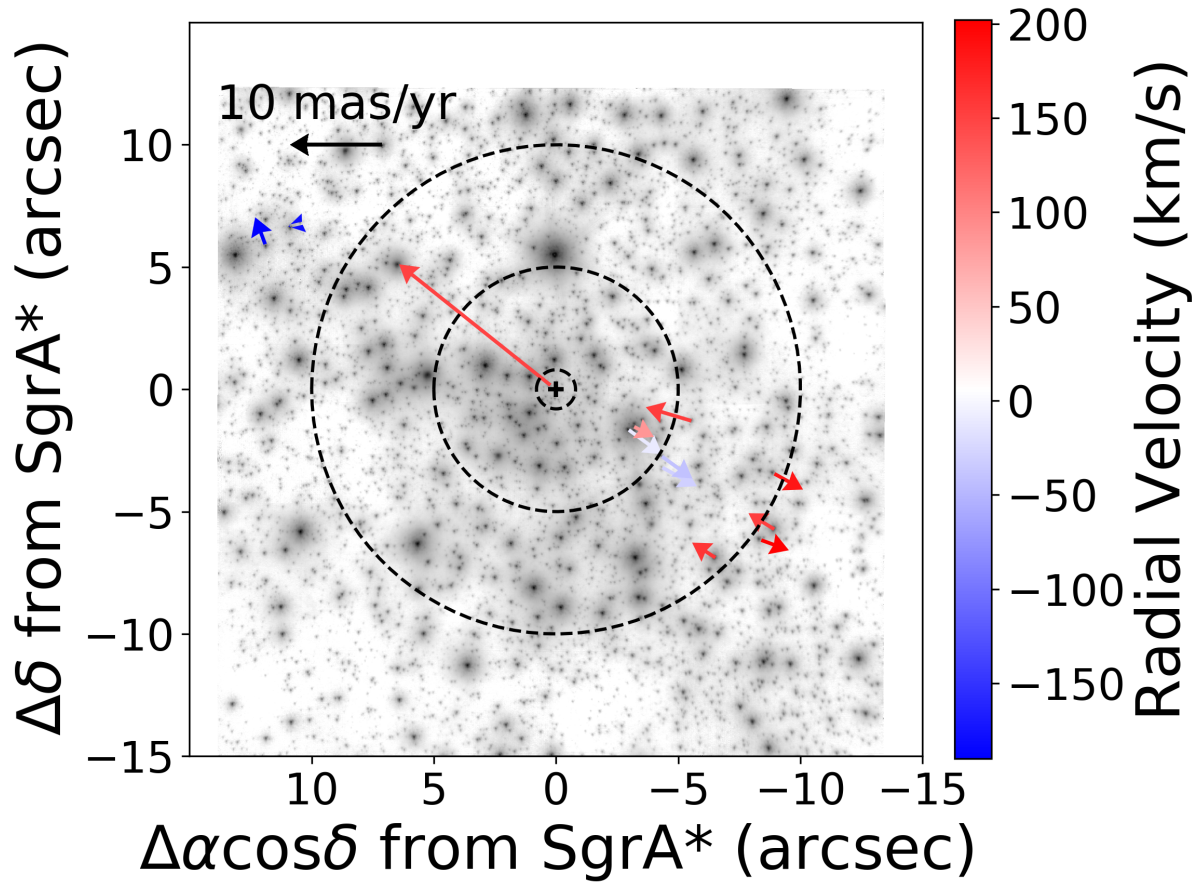


Figure 3.15: The quiver plot of all potential disk2 stars, with color showing the RV.

3.6 Summary

In summary, we analysed the dynamical structures for 88 young stars in our Galactic center with projected radii from 0'' to 15'' with well measured proper motion, radial velocity and spectroscopic completeness correction. The proper motion comes from a long time baseline observation with 10m Keck telescope since 1995 in the inner region and HST WFC-IR observation in the outer region. The radial velocity comes from our GCOWS survey using diffraction-limited Keck OSIRIS spectra, combined with RVs published in [20]. This is the first time a sample of GC young stars are published with proper motion, radial velocity and completeness correction, enabling us to simulate star cluster and redden it to directly compare with what is actually observed on sky. We use this sample and detected the well known clockwise disk (disk1 in this paper), which is consistent with previously published results. Another almost edge-on counter-clockwise disk is found for the first time (disk2 in this paper). Disk2 is highly asymmetric with stars concentrated in the northwest direction, and two IRS 13 stars are found to potentially reside on this disk. Disk1 has an eccentricity peaking at 0.2, consistent with previous analysis, but disk2 is even more eccentric than disk1, with $P(e) \propto e$. By simulating disk1, disk2 and non-disk stars using the observed disk property, we are able to reproduce density map for disk1 and non-disk stars, but the highly asymmetric structure in disk2 can not be explained by extinction alone. In future studies, we will use this sample with known completeness correction to constrain IMF for different dynamical subgroups, which will greatly help us understand their formation history.

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Appendix A

Analysis details

A.1 Chapter 2

Central Arcsecond Astrometry Correction

We refine our method for local distortion correction in the central arcsecond in order to include many additional stars with motions on the sky that exceed a linear model. This refined correction is applied only to the observed positions of S0-2 (Do et al. in prep.). The correction procedure incorporates orbital fits to seven stars close to Sgr A* as well as linear and acceleration model fits to all other stars in the central arcsecond with $r < 1''$ across all epochs of observations (1995 – 2018).

We fit linear, acceleration, and orbital models to the stars located in the central arcsecond. We first generated orbital models for seven stars (S0-1, S0-3, S0-5, S0-16, S0-19, S0-20, S0-38). These seven stars are those for which we have radial velocity measurements and whose orbital motions include a turning point in astrometry. S0-2, the star of interest, is excluded from the sample of orbital stars used to calculate the correction. For the remaining stars within the central arcsecond around Sgr A* ($r < 1''$), we fit both linear and accelerating models to their measured astrometric positions. We calculated the χ^2_{red} statistic for the linear and acceleration fits of each of these stars. We selected a linear model for those stars that had a lower χ^2_{red} under the linear model compared to the acceleration model. An acceleration model was used otherwise. We next removed those stars that had $\chi^2_{\text{red}} > 10$ in all cases since their motions were not well fit by either the linear or acceleration models. We also removed stars detected in fewer than 28 epochs since these stars may not have enough astrometric detections for well-constrained fits. Ultimately, we obtained one star in the central arcsecond with astrometric measurements well fit by a linear model and 19 stars with astrometric measurements well fit by an acceleration model. Along with the seven orbital model stars, we obtained a total of 27 stars for calculating the central arcsecond astrometry correction.

We next calculated the correction to S0-2's astrometric measurement using the stars detected in the central arcsecond in each epoch. For each star in the central arcsecond as-

trometry correction sample, we calculated the residual of each measured astrometric positions from the star’s respective proper motion model (i.e.: measured position – model position). Uncertainties on the astrometry residual included uncertainty in the astrometric position (including both positional and alignment errors) and uncertainty in each star’s respective proper motion model. For each epoch, we then calculated the weighted mean of the astrometry residuals of all central arcsecond astrometry correction stars detected in that epoch, $\bar{x}, \bar{y}$, defined in the following way:

$$\bar{x} = \frac{\sum w_{x,i} x_i}{\sum w_{x,i}}$$

Here, $w_{x,i}$ represents the weight on each star’s measured position x_i . The weights, $w_{x,i}$ and $w_{y,i}$, were calculated from the uncertainty on the astrometric differences, $\sigma_{x,i}$ and $\sigma_{y,i}$: $w_{x,i} = 1/\sigma_{x,i}^2$. In each epoch, the weighted mean of the residuals was subtracted from S0-2’s astrometry measurements to derive the distortion corrected position for S0-2.

Uncertainties on the astrometric residuals were calculated by bootstraps. In each epoch, we constructed 1000 bootstrap trials. Each bootstrap trial had a full sample, drawn randomly with replacement, for that epoch’s stellar astrometry differences and associated uncertainties. The weighted mean was then calculated on each bootstrap trial. The uncertainty on the weighted mean astrometric difference for each epoch was next calculated as half of the median-centered 1σ (i.e. 68%) range of the bootstrap trial means for the epoch.

The astrometric residuals used to correct the observed astrometric measurements of S0-2 when deriving its orbit are shown in Figure A.1. In speckle holography epochs (1995 – 2005), median absolute corrections are 0.162 mas and 0.173 mas in x and y , respectively, with median bootstrap uncertainties of 0.153 mas and 0.228 mas in x and y , respectively. Median alignment uncertainties for S0-2 before local distortion correction in speckle holography epochs is 0.93 mas and 0.76 mas in x and y , respectively. In 2005 AO epochs, median absolute corrections are 0.189 mas and 0.319 mas in x and y , respectively, with median bootstrap uncertainties of 0.089 mas and 0.167 mas in x and y , respectively. Median alignment uncertainties for S0-2 before local distortion correction in 2005 AO epochs is 0.21 mas and 0.32 mas in x and y , respectively. In 2006–2014 AO epochs (i.e. *06–14* setup), median absolute corrections are 0.063 mas and 0.098 mas in x and y , respectively, with median bootstrap uncertainties of 0.063 mas and 0.072 mas in x and y , respectively. Median alignment uncertainties for S0-2 before local distortion correction in 2006–2014 AO epochs is 0.14 mas and 0.15 mas in x and y , respectively. In 2015–2018 AO epochs, median absolute corrections are 0.251 mas and 0.222 mas in x and y , respectively, with median bootstrap uncertainties of 0.099 mas and 0.177 mas in x and y , respectively. Median alignment uncertainties for S0-2 before local distortion correction in 2015–2018 AO epochs is 0.26 mas and 0.26 mas in x and y , respectively.

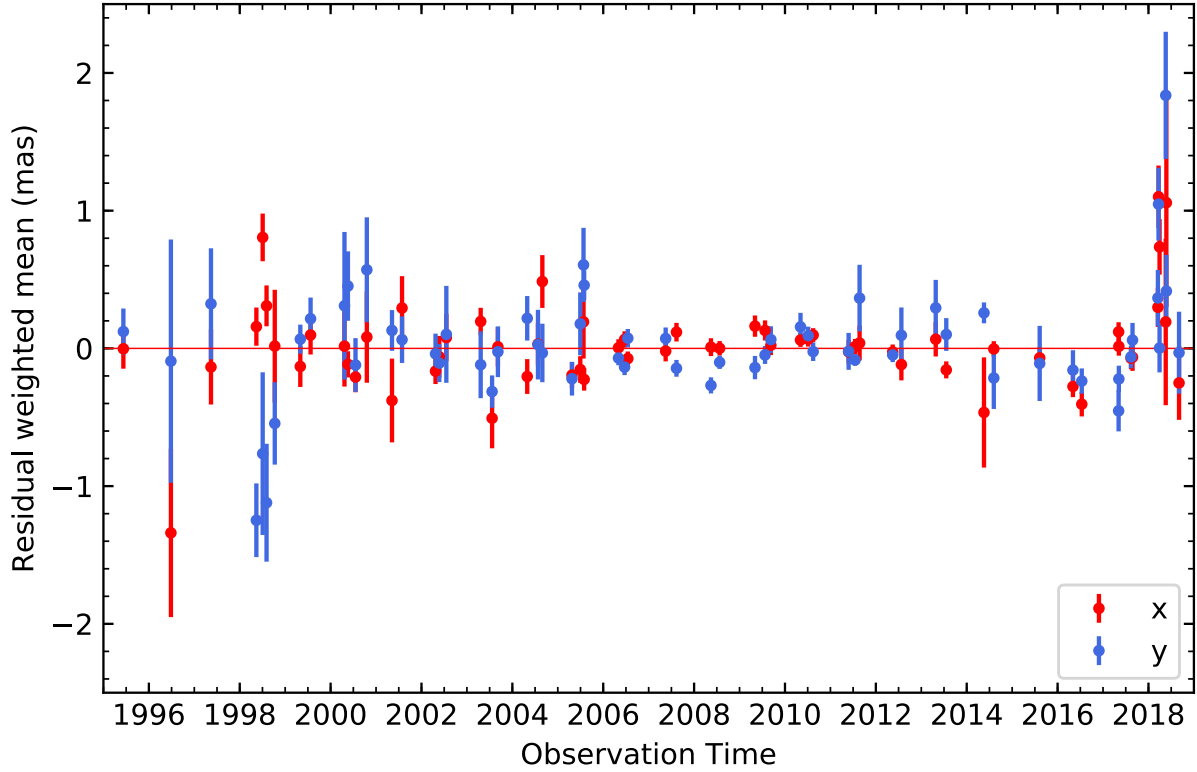
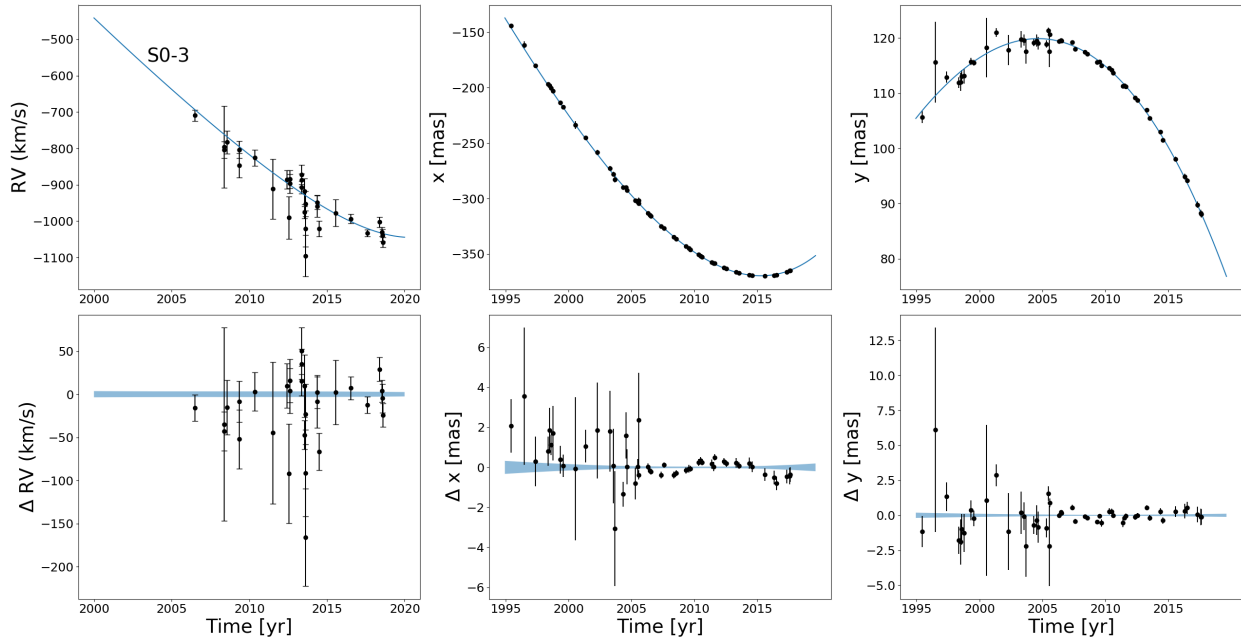
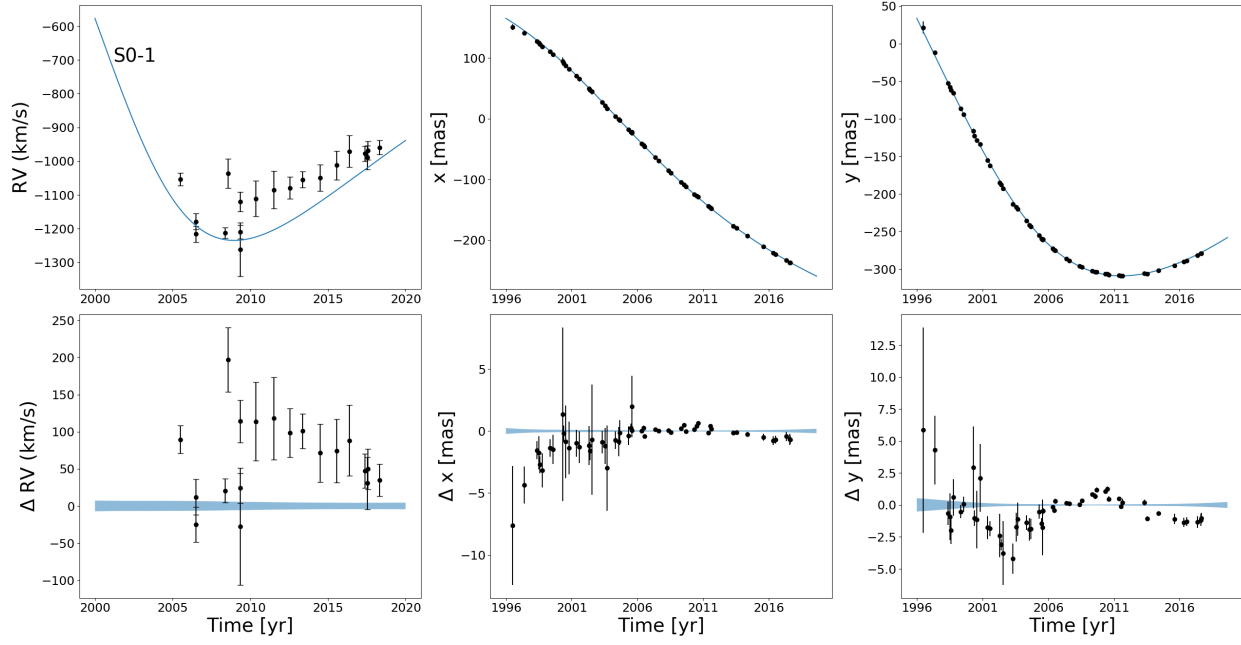


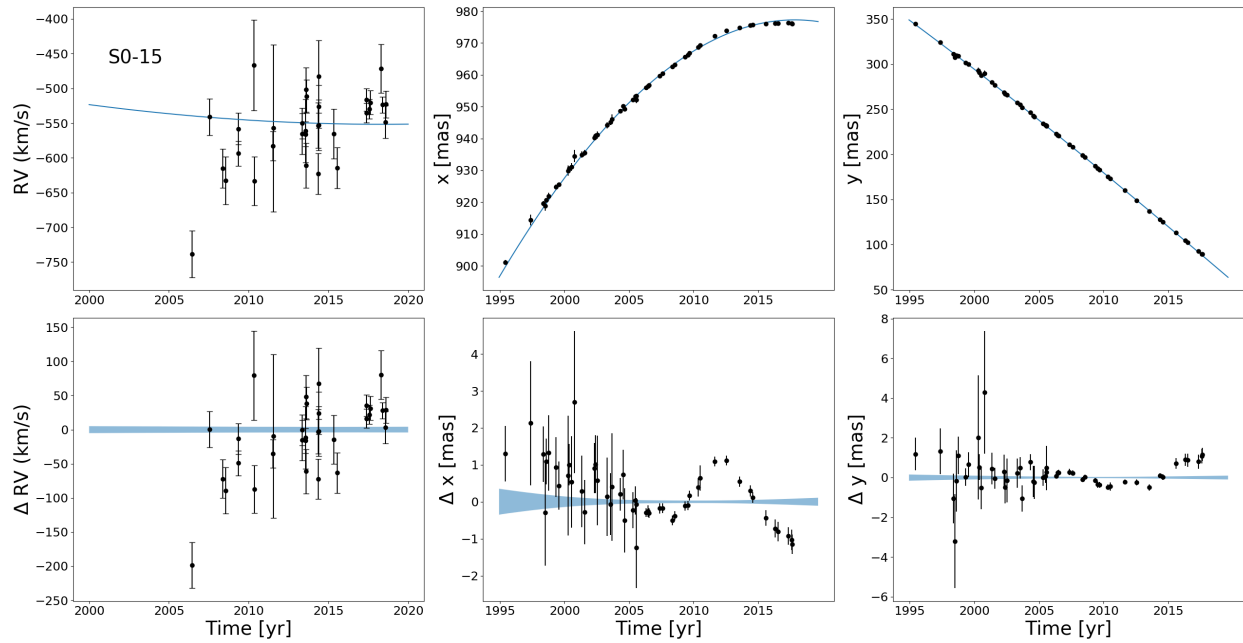
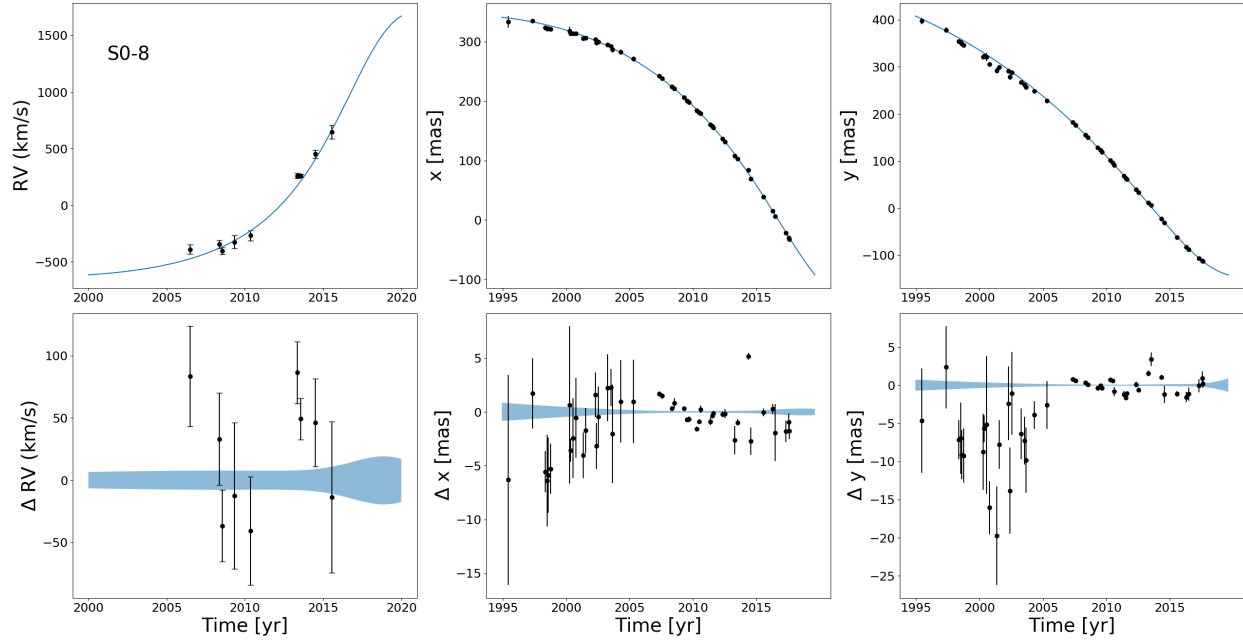
Figure A.1: Residuals in x and y derived for S0-2's position, using 27 stars in the central arcsecond fitted to linear, accelerating, or orbital motion models. These residuals were subtracted from S0-2's astrometric measurements to derive the distortion corrected positions for S0-2.

A.2 Chapter 3

Full Keplerian Orbit Analysis

Here we plot the full keplerian orbit fit for the other 6 orbit stars in our sample: S0-1, S0-3, S0-8, S0-15, S0-19, S1-2, in the same format as in Figure 3.5.





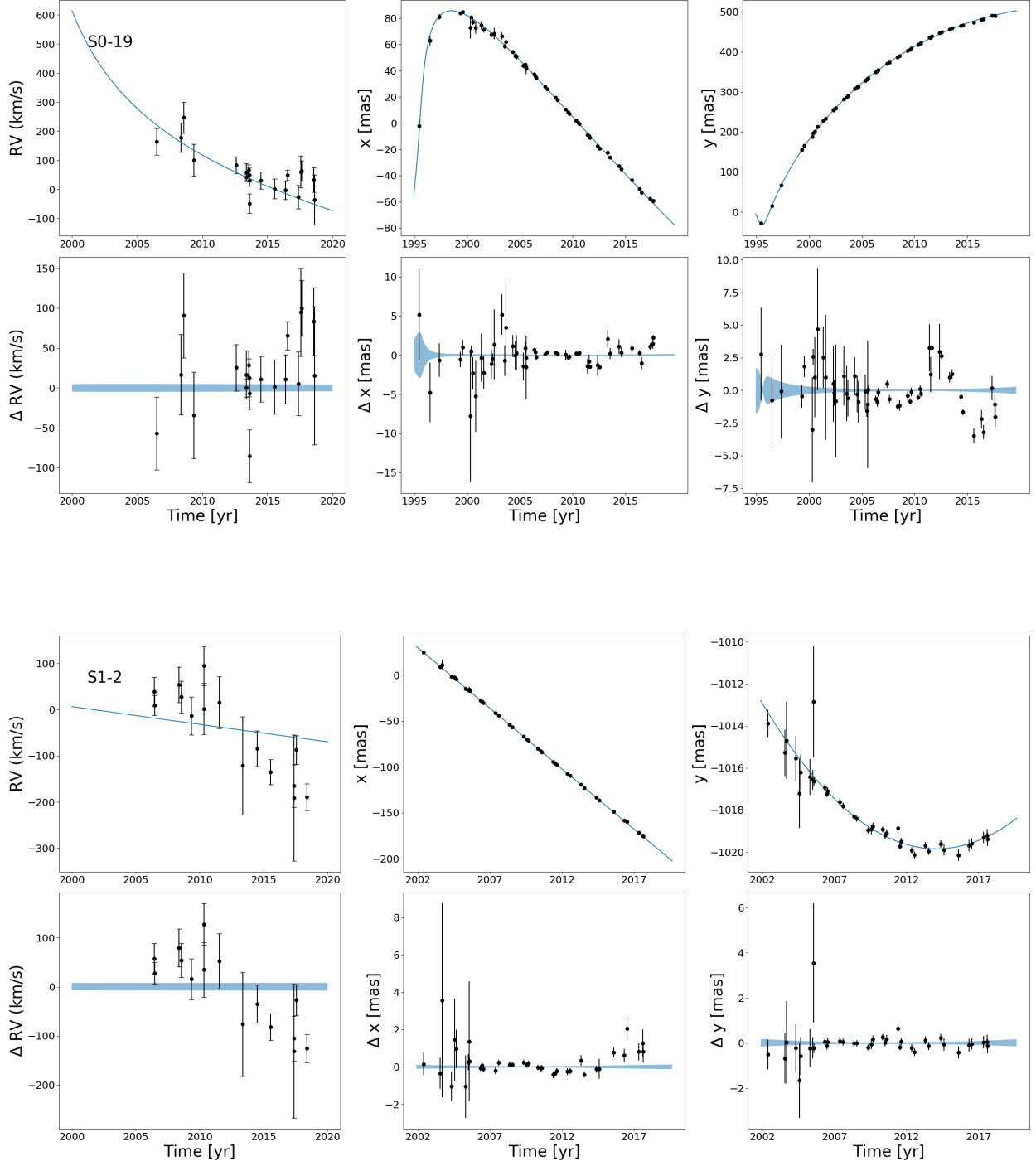


Figure A.1: Examples of full Keplerian orbit analysis for S0-1