Galactic Dynamics

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ABSTRACT
I thought I would try out the MNRAS LATEX style.

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1 EVENT RATES

1.1 The density function

We wish to calculate the probability that there is an encounter between a compact object on an orbital trajectory described by eccentricity e and periapse radius r_p and the supermassive black hole at the galactic centre. We begin by following the work of Bahcall & Wolf (1976, 1977) and assuming that the density function within the galactic core is just a function of the orbital energy; the number of stars is given by

$$N = \int \mathrm{d}^3 r \int \mathrm{d}^3 v f(\mathcal{E}). \tag{1}$$

We will define the energy per unit mass of the orbit as

$$\mathcal{E} = \frac{v^2}{2} - \frac{GM_\bullet}{r} \tag{2}$$

where $M_\bullet$ is the mass of the supermassive black hole. Close to the centre of the galactic core dynamics will be dominated by the influence of the black hole as it is significantly more massive than the surrounding stars. We will define the radius of influence for the black hole as

$$r_c = \frac{GM_\bullet}{\sigma_\star^2} \tag{3}$$

where σ^2 is the line-of-sight velocity dispersion. We will assume that the mass of stars enclosed within the radius is greater than the black hole mass, which is much greater than the mass of a typical star $M_\star$ (Bahcall & Wolf 1976). We will define a reference number density from the enclosed mass as

$$m_\star(r_c) = \frac{4\pi r_c^3}{3} n_\star M_\star. \tag{4}$$

Within the the core, the density function can be calculated using the approximation of Fokker-Planck formalism. The population of bound ($\mathcal{E} < 0$) stars is evolved numerically until a steady state is reached: the unbound ($\mathcal{E} > 0$) stars

form a reservoir with an assumed Maxwellian distribution. Denoting a species of star by its mass M:

$$f_M(\mathcal{E}) = \frac{C_M n_\star}{(2\pi\sigma_M^2)^{3/2}} \exp\left(-\frac{\mathcal{E}}{\sigma_M^2}\right) \tag{5}$$

where C_M is a normalisation constant.[1] If different stellar species are in equipartition (as was assumed by Bahcall & Wolf (1976, 1977)) then we expect

$$M\sigma_M^2 = M_\star \sigma_\star^2. \tag{6}$$

However, if the unbound stellar population has reached equilibrium by violent relaxation (Lynden-Bell, D. 1967), then all mass groups are expected to have similar velocity dispersions: this has been used by Alexander & Hopman (2009); O'Leary et al. (2009) and will be assumed here. The steady-state density function is largely insensitive to this choice of boundary condition (Bahcall & Wolf 1977; Alexander & Hopman 2009).

For bound orbits ($\mathcal{E} < 0$) the density function can be approximated as a power law

$$f_M(\mathcal{E}) = \frac{k_M n_\star}{(2\pi\sigma_M^2)^{3/2}} \left(-\frac{\mathcal{E}}{\sigma_M^2}\right)^{p_M}. \tag{7}$$

The exponent p_M varies depending upon the mass of the object, yielding mass segregation. For a system with a single mass component, Bahcall & Wolf (1976) find that $p = 1/4$. The normalisation constant k_M reflects the relative abundances of the different species.[2]

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[1] C_M determines the population ratio of species M far from the black hole Alexander & Hopman (2009).
[2] For a single mass population ($p = 1/4$) $k = 2C$ gives a fit correct to within a factor a two (Bahcall & Wolf 1976; Keshet et al. 2009), we will assume this holds for the dominant species of stars as although it will vary slightly with p this change is small compared to errors introduced by fitting a simple power law (Hopman & Alexander 2006a; Alexander & Hopman 2009).

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Table 1. Stellar model parameters for the galactic centre using the results of Alexander & Hopman (2009) We use the main sequence star as our reference. The number fractions for unbound stars are estimates corresponding to a model of continuous star formation (Alexander 2005).

Star	$M/M_\odot$	$C_M/C_\star$	p_M	$k_M/k_\star$ [a]
MS	1.0	1	−0.1	1
WD	0.6	0.1	−0.1	0.09
NS	1.4	0.01	0	0.01
BH	10	0.001	0.5	0.008

[a] Toonen et al. (2009)

1.2 Model parameters

We will be using the Fokker-Planck model of Hopman & Alexander (2006a,b); Alexander & Hopman (2009). This includes four stellar species: main sequence stars (MS), white dwarfs (WD), neutron stars (NS), and black holes (BH). Their properties are summarised in table 1. The steeper power law for black holes means that they segregate about the massive black hole: they dominate in place of main sequence stars for radii $r < 10^{-4} r_c$.

We assume a black hole mass of $M_\bullet = 4.31 \times 10^6 M_\odot$ Gillessen et al. (2009) and a velocity dispersion of $\sigma = 103 \text{ km s}^{-1}$ Tremaine et al. (2002). This gives a cusp radius of $r_c = 1.75$ pc. Using the results of Ghez et al. (2008) we would expect the total mass of stars in the core to be $m_\star(r_c) = 6.4 \times 10^6 M_\odot$ (which is within 10% of the value obtained similarly from Genzel et al. (2003)). This gives a stellar density of $n_\star = 2.8 \times 10^5 \text{ pc}^{-3}$.

1.3 Eccentricity & periapsis

We wish to make a change of variables to characterise orbits by their eccentricity e and periapse radius r_p. The latter, unlike the semimajor axis, is always well defined regardless of eccentricity. For Keplerian orbits, the energy $\mathcal{E}$ and angular momentum J per unit mass are entirely characterised by these parameters

$$\mathcal{E} = -\frac{GM_\bullet(1-e)}{2r_p}, \tag{8}$$

$$J^2 = GM_\bullet(1+e)r_p. \tag{9}$$

We start by decomposing the velocity into three orthogonal components: radial v_r, azimuthal v_ϕ and polar v_θ. We will assume that the galactic core is spherically symmetric (Genzel et al. 2003; Schödel et al. 2007), therefore we are only interested in the combination

$$v_\perp^2 = v_\phi^2 + v_\theta^2 = v^2 - v_r^2. \tag{10}$$

Under this change of variables

$$d^2v = dv_r dv_\phi dv_\theta \to 2\pi v_\perp dv_r dv_\perp. \tag{11}$$

The specific energy and angular momentum are given by

$$\mathcal{E} = \frac{v_r^2 + v_\perp^2}{2} - \frac{GM_\bullet}{r}, \tag{12}$$

$$J^2 = r^2 v_\perp^2. \tag{13}$$

If we combine these with our earlier expressions in terms of e and r_p we find

$$v_\perp^2 = \frac{GM_\bullet(1+e)r_p}{r^2}, \tag{14}$$

$$v_r^2 = GM\left[\frac{2}{r} - \frac{(1-e)}{r_p} - \frac{(1+e)r_p}{r^2}\right]. \tag{15}$$

From the later we can verify that the turning points of an orbit occur at

$$r = r_p, \quad \frac{1+e}{1-e}r_p; \tag{16}$$

the periapse is the only turning point for orbits with $e > 1$. Since we now have expressions for $\{v_r, v_\perp\}$ in terms of $\{e, r_p\}$ we can calculate the Jacobian

$$\left|\frac{\partial(v_r, v_\perp)}{\partial(e, r_p)}\right| = \frac{\partial v_r}{\partial e}\frac{\partial v_\perp}{\partial r_p} - \frac{\partial v_r}{\partial r_p}\frac{\partial v_\perp}{\partial e} \tag{17}$$

$$= \frac{1}{2v_r v_\perp}\frac{e}{r_p}\left(\frac{GM}{r}\right)^2. \tag{18}$$

Using this, we may rewrite our velocity element as

$$d^3v = dv_r dv_\phi dv_\theta \tag{19}$$

$$= 2\pi v_\perp dv_r dv_\perp \tag{20}$$

$$= \frac{\pi e}{v_r r_p}\left(\frac{GM}{r}\right)^2 de dr_p. \tag{21}$$

As a consequence of our assumed spherical symmetry, the volume element can be expressed as

$$d^3r = 4\pi r^2 dr. \tag{22}$$

Thus the phase space volume element can be expressed as

$$d^3r d^3v = \frac{4\pi^2(GM)^2 e}{v_r r_p}dr de dr_p. \tag{23}$$

The number of stars in an element $dr de dr_p$ is

$$n(r, e, r_p) = \frac{4\pi^2(GM)^2 e}{v_r r_p}f(\mathcal{E}). \tag{24}$$

From this, we can construct the expected number of stars to be on orbits with periapsis r_p and eccentricity e. We will define this locally, allowing it to vary with position. The number of stars found in a small radius range δr with given orbital properties should be given by the the total number of stars with these properties multiplied by the relative amount of time they spend in that range

$$n(r, e, r_p)\delta r = N(e, r_p; r)\frac{\delta t}{P(e, r_p)} \tag{25}$$

where $N(e, r_p; r)$ is the total number of stars with orbits given by $\{e, r_p\}$ defined at r, δt is the time spent in δr and $P(e, r_p)$ is the period of the orbit. We will defer the definition of this time for unbound orbits for now. The time spent in the radius range is

$$\delta t = 2\frac{\delta r}{v_r}, \tag{26}$$

where the factor of 2 is included to account for both inwards and outwards motion. Hence

$$N(e, r_p; r) = \frac{1}{2}v_r P n(r, e, r_p) \tag{27}$$

$$= \frac{2\pi^2(GM)^2 Pe}{r_p}f(\mathcal{E}). \tag{28}$$

The right hand side of this equation is independent of position, subject to the constraint that the radius is in the allowed range for the orbit $r_p \leqslant r \leqslant (1+e)r_p/(1-e)$, and so we may define $N(e, r_p) \equiv N(e, r_p; r)$. This is a consequence of the density function being dependent only upon an orbital constant of the motion.

If a burst of radiation is emitted each time a star passes through periapse, then the event rate for burst emission from orbits with parameters $\{e, r_p\}$, is given by

$$\Gamma(e, r_p) = \frac{N(e, r_p)}{P(e, r_p)} \tag{29}$$

$$= \frac{2\pi^2 (GM)^2 e}{r_p} f(\mathcal{E}). \tag{30}$$

The orbital period drops out from the calculation, so we do not have to worry about an appropriate definition for unbound orbits.

From the event rate we may define a probability of seeing a given number of events subject to the assumption that these events are uncorrelated. The probability is simply given by the Poisson distribution; the probability of there being r events is

$$\Pr(r | \Gamma(e, r_p)) = \frac{\Gamma^r \exp(-\Gamma)}{r!}. \tag{31}$$

The probability of there being a burst from an orbit with periapse r_p and eccentricity e is hence

$$\Pr(r \neq 0 | \Gamma(e, r_p)) = 1 - \Pr(r = 0 | \Gamma(e, r_p)). \tag{32}$$

To estimate the expectation of a quantity across all orbits we use

$$\langle X \rangle = \sum_r \int_0^\infty \mathrm{d}e \int_0^\infty \mathrm{d}r_p X(r; r_p, e) \Pr(r | \Gamma(e, r_p)). \tag{33}$$

Since the probability decays rapidly for large r, we may truncate the sum to give the required level of accuracy,

To generate a representative sample for the orbital parameters e and r_p, we use $\Gamma(e, r_p)$ as an unnormalised probability distribution and draw from it appropriately.

1.4 The inner cut-off

From (30) we see that the event rate is highly sensitive to the smallest value of the periapsis. The inner cut-off for r_p could result from a number of different physical causes. Ultimately the orbits cannot encroach closer to the black hole than its last stable orbit. This depends upon the spin of the black hole, but is of the order of its Schwarzschild radius. Before we reach this point, however, there are other processes that may intervene to deplete the orbiting stars. Our treatment of these is approximate, however should hopefully produce reasonable estimates. We will consider three processes: tidal disruption, gravitational wave inspiral and collisional disruption. Tidal disruption imposes a definite definite cut-off, while the others use statistical arguments, they are therefore true for a typical star and it is unlikely that a star would be found beyond the imposed limits.

1.4.1 Tidal disruption

Tidal forces from the black hole can disrupt stars. This occurs at the tidal radius

$$r_t \simeq \left(\frac{M_\bullet}{M} \right)^{1/3} R_M \tag{34}$$

where R_M is the radius of the star (Kobayashi et al. 2004). Any star on an orbit with $r_p < r_T$ will be disrupted in the cause of its orbit. Tidal disruption is most significant for MS stars since they are least dense.

1.5 Gravitational wave inspiral

Stars orbiting about the black hole will continually radiate energy and angular momentum causing them to inspiral. Using the analysis of Peters (1964) for Keplerian binaries, for bound orbits the orbit-averaged rate of change of the periapsis and eccentricity are

$$\left\langle \frac{\mathrm{d}r_p}{\mathrm{d}t} \right\rangle = -\frac{64}{5} \frac{\Theta}{r_p^3} \frac{(1-e)^{3/2}}{(1+e)^{7/2}} \left(1 - \frac{7}{12}e + \frac{7}{8}e^2 + \frac{47}{192}e^3 \right) \tag{35}$$

$$\left\langle \frac{\mathrm{d}e}{\mathrm{d}t} \right\rangle = -\frac{304}{5} \frac{\Theta}{r_p^4} \frac{e(1-e)^{3/2}}{(1+e)^{5/2}} \left(1 + \frac{121}{304}e^2 \right), \tag{36}$$

where we have introduced

$$\Theta = \frac{G^3 M_\bullet M (M_\bullet + M)}{c^5}. \tag{37}$$

For a circular orbit the inspiral time form initial periapsis r_{p0} is

$$\tau_c(r_{p0}) = \frac{5}{256} \frac{r_{p0}^4}{\Theta}. \tag{38}$$

For an orbit of finite eccentricity ($0 < e < 1$), we can solve for the periapsis as a function of eccentricity

$$r_p(e) = A(1+e)^{-1} \left(1 + \frac{121}{304}e^2 \right)^{870/2299} e^{12/19}, \tag{39}$$

where A is a constant fixed by the initial conditions: for an orbit with initial eccentricity e_0

$$A = (1+e_0) \left(1 + \frac{121}{304}e_0^2 \right)^{-870/2299} e_0^{-12/19} r_{p0}. \tag{40}$$

The inspiral is complete when the eccentricity has decayed to zero. Consequently the inspiral time is (Peters 1964)

$$\tau(r_{p0}, e_0) = \int_0^{e_0} \frac{15}{304} \frac{A^4}{\Theta} \frac{e^{29/19}}{(1-e^2)^{3/2}} \left(1 + \frac{121}{304}e^2 \right)^{1181/2299} \mathrm{d}e. \tag{41}$$

This is best evaluated numerically, however it may be written in closed from as

$$\tau(r_{p0}, e_0) = \tau_c(r_{p0})(1+e_0)^4 \left(1 + \frac{121}{304}e_0^2 \right)^{-3480/2299}$$
$$\times F_1 \left(\frac{24}{19}; \frac{3}{2}, -\frac{1181}{2299}; \frac{43}{19}; e_0^2, -\frac{121}{304}e_0^2 \right), \tag{42}$$

using the Appell hypergeometric function of the first kind $F_1(\alpha; \beta, \beta'; \gamma; x, y)$.[3]

We will assume that an orbit is depleted of stars if the inspiral timescale is shorter than the relaxation timescale. This indicates the characteristic time for two-body collisions to change the velocity the star by order of itself (Binney & Tremaine 1987), and so indicates the time over which scattering may repopulate the orbit. Binney & Tremaine (1987)

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[3] For small eccentricities $\tau(r_{p0}, e_0) \simeq \tau_c(r_{p0})[1 + 4e_0 + (273/43)e_0^2 + \mathcal{O}(e_0^3)]$.

derive the relaxation timescale from the diffusion coefficient of the Fokker-Planck equation (section 8.4), for a system with a purely Maxwellian distribution

$$\tau_R \simeq 0.34 \frac{\sigma_\star^3}{GM_\star \rho_\star \ln \Lambda},$$ (43)

where $\rho_\star$ is the mass density and the Coulomb logarithm is $\ln \Lambda = \ln(M_\bullet / M_\star)$. Bahcall & Wolf (1977) find a similar characteristic timescale, but with numerical prefactor $3/4\sqrt{8\pi} \simeq 0.15$. We shall adopt the former to be conservative. From this we estimate the relaxation time for an orbit as

$$T_R(r_p, e) = \left(\frac{\langle v(r) \rangle_{\text{orb}}}{\sigma_\star} \right)^3 \frac{M_\star \rho_\star}{\langle M(r) \rangle_{\text{orb}} \langle \rho(r) \rangle_{\text{orb}}} \tau_R$$ (44)

$$= \left(\frac{2E(e)}{\pi} \sqrt{\frac{r_c(1-e)}{r_p}} \right)^3 \frac{M_\star \rho_\star}{\langle M(r) \rangle_{\text{orb}} \langle \rho(r) \rangle_{\text{orb}}} \tau_R$$ (45)

using the velocity $\langle v(r) \rangle_{\text{orb}}$, stellar mass of encountered stars $\langle M(r) \rangle_{\text{orb}}$ and density $\langle \rho(r) \rangle_{\text{orb}}$ averaged over the duration of the orbit as characteristic quantities. The subscript orb indicates that we average along the trajectory of the orbit specified by parameters $\{r_p, e\}$.[4] While the former may be written in closed form using the complete elliptic integral of the second kind, the latter is best evaluated numerically.

Unbound stars only undergo a single periapse passage and only radiate one burst of radiation, we shall therefore neglect any evolution in their orbital parameters.[5]

1.5.1   *Collisions*

As a consequence of the higher densities in the galactic core, stars may undergo a large number of close encounters with other stars. It takes 20–30 grazing collisions to disrupt a MS star (Freitag et al. 2006). We shall ignore the possibility of disruption for the other species since they have much smaller cross-sectional areas. The number of collisions a star will undergo in a time interval δt is

$$\delta K = n(r) \pi R_\star^2 v(r, e, r_p) \delta t.$$ (46)

For circular orbits we can find the radius at which collisions will lead to disruptions by setting $\delta K = 30$ and $\delta t = \tau_R$. We use the relaxation timescale for the system as this is the time over which stars are replenished from the reservoir. For non-circular orbits we must consider variation with position. Using $\delta r = v_r \delta t$, and then converting to an integral, we have for bound orbits

$$K = 2\pi R_\star^2 \frac{\tau_R}{P(r_p, e)} \int_{r_p}^{(1+e)r_p/(1-e)} n(r) \frac{v(r, e, r_p)}{v_r(r, e, r_p)} \, \mathrm{d}r,$$ (47)

[4] In calculating $\langle \rho(r) \rangle_{\text{orb}}$ and $\langle M(r) \rangle_{\text{orb}}$ we include the tidal disruption of stars but not the possible depopulation of orbits due to inspiral or collisions. This would require solving iteratively for $n(r)$, which is not justifiable at this level of approximation. The relaxation timescale will be underestimated using this simple approach. The error will be greatest for low eccentricity orbits which spend most of their time in the same region of space.

[5] Using the analysis of Turner (1977) it is possible to show that the change in eccentricity Δe, and fractional change in periapsis $\Delta r_p / r_p$ for an extreme mass-ratio binary are less than $\mathcal{O}(10\eta)$, where $\eta = M / M_\bullet$.

where P is the period of the orbit. Again we may set $K = 30$ to find the orbits for which stars will be disrupted within a relaxation timescale. For unbound stars we are only interested in stars that would become disrupted before their periapse passage, so

$$K = \pi R_\star^2 \int_{r_p}^{r_c} n(r) \frac{v(r, e, r_p)}{v_r(r, e, r_p)} \, \mathrm{d}r,$$ (48)

assuming that the stars in the reservoir external to the core are unlikely to undergo close collisions.

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APPENDIX A: EVOLUTION OF ORBITAL PARAMETERS FOR UNBOUND ORBITS DUE TO GRAVITATIONAL WAVE EMISSION

Following the approach of Turner (1977) we can calculate the evolution of the eccentricity and periapse of a Keplerian binary through the loss of energy and angular momentum through carried away by gravitational radiation. The change in eccentricity over an orbit, approximating the orbital parameters remain constant throughout the orbit, is

$$
\begin{aligned}
\Delta e \quad = \quad & -\frac{608}{15}\Sigma\left[\frac{e}{(1+e)^{5/2}}\left(1+\frac{121}{304}e^2\right)\cos^{-1}\left(-\frac{1}{e}\right)\right. \\
& \left.+\frac{(e-1)^{1/2}}{e(1+e)^2}\left(\frac{67}{456}+\frac{1069}{912}e^2+\frac{3}{38}e^4\right)\right],
\end{aligned} \qquad \text{(A1)}
$$

introducing dimensionless parameter

$$
\Sigma = \frac{G^{5/2}M_\bullet M(M_\bullet+M)}{c^5 r_p^{5/2}}. \qquad \text{(A2)}
$$

Similarly, the fractional change in periapsis is

$$
\begin{aligned}
\frac{\Delta r_p}{r_p} \quad = \quad & -\frac{128}{5}\Sigma\left[\frac{1}{(1+e)^{7/2}}\left(1-\frac{7}{12}e+\frac{7}{8}e^2+\frac{47}{192}e^3\right)\cos^{-1}\left(-\frac{1}{e}\right)\right. \\
& \left.-\frac{(e-1)^{1/2}}{e(1+e)^2}\left(\frac{67}{288}-\frac{13}{8}e+\frac{133}{576}e^2-\frac{1}{4}e^3-\frac{1}{8}e^4\right)\right].
\end{aligned} \qquad \text{(A3)}
$$

For extreme mass-ratio binaries, as is the case here, the mass-ratio is a small quantity

$$
\eta = M/M_\bullet \ll 1. \qquad \text{(A4)}
$$

The smallest possible periapsis is off order of the Schwarzschild radius of the massive black hole, such that $r_p > GM_\bullet/c^2$. These give

$$
\Sigma < \eta \ll 1. \qquad \text{(A5)}
$$

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