ABSTRACT

It is well accepted that ‘hot Jupiters’ and other short-period planets did not form in situ, as the temperature in the protoplanetary disc at the radius at which they now orbit would have been too high for planet formation to have occurred. These planets, instead, form at larger radii and then move into the region in which they now orbit. The exact process that leads to the formation of these close-in planets is, however, unclear and it seems that there may be more than one mechanism that can produce these short-period systems. Dynamical interactions in multiple-planet systems can scatter planets into highly eccentric orbits which, if the pericentre is sufficiently close to the parent star, can be tidally circularized by tidal interactions between the planet and star. Furthermore, systems with distant planetary or stellar companions can undergo Kozai cycles which can result in a planet orbiting very close to its parent star. However, the most developed model for the origin of short period planets is one in which the planet exchanges angular momentum with the surrounding protoplanetary disc and spirals in towards the central star. In the case of ‘hot Jupiters’, the planet is expected to open a gap in the disc and migrate in what is known as, the Type II regime. If this is the dominant mechanism for producing ‘hot Jupiters’ then we would expect the correct properties of observed close-in giant planets to be consistent with an initial population resulting from Type II migration followed by evolution due to tidal interactions with the central star. We consider initial distributions that are consistent with Type II migration and find that after tidal evolution, the final distributions can be consistent with that observed. Our results suggest that a modest initial pile-up at $a \sim 0.05$ au is required and that the initial eccentricity distribution must peak at $e \sim 0$. We also suggest that if higher mass close-in exoplanets preferentially have higher eccentricities than lower mass exoplanets, this difference is primordial and is not due to subsequent evolution.

Key words: planets and satellites: dynamical evolution and stability – planets and satellites: formation – planet-disc interactions – planet-star interactions – planetary systems.
These different mechanisms will result in different distributions of close-in planets. Early simulations of planet–planet scattering (Ford, Havlizzyka & Rasio 2001) found that the likelihood of a massive planet being scattered into an orbit that brought it very close to the central star was less than 1 per cent, suggesting that this was unlikely to be the primary mechanism for producing ‘hot Jupiters’. Longer term integration (Marzari & Weidenschilling 2002) increases this to 1–10 per cent, while Nagasawa, Ida & Bessho (2008) suggest that ‘Kozai migration’ (Wu & Murray 2003) can further increase this to 30–50 per cent. Recent observations (Winn et al. 2009; Collier Cameron et al. 2010; Queloz et al. 2010) of exoplanets with orbits that are retrograde with respect to the rotation of the their central star suggests that ‘Kozai migration’ must play a role in the formation of some close-in planets (Triaud et al. 2010).

Theoretical modelling (Fabrycky & Tremaine 2007) and analysis of the current sample of close-in exoplanets (Morton &Johnson 2011) suggest that Kozai cycles together with tidal friction can produce some, but not most, of the close-in planets. In this work, we intend to investigate if initial conditions consistent with Type II migration of giant planets can lead – after tidal evolution – to distributions that are consistent with those observed. Although it is fairly clear that Type II migration is not the only mechanism capable of producing close-in giant planets, if it is the dominant mechanism we would at least expect the current distribution to be consistent with one having evolved from an initial distribution resulting from Type II migration. The paper is structured as follows. Section 2 describes the basic model and assumptions. Section 3 presents the results and in Section 4 we discuss the results and draw conclusions.

2 BASIC MODEL

In this paper, we use knowledge of Type II migration to choose initial semi-major and eccentricity distributions of gas giant planets around Sun-like stars and evolve these systems by integrating the equations that describe the tidal evolution of these systems.

2.1 Tidal evolution

We base our model of tidal evolution on that described in Dobbs-Dixon, Lin & Mardling (2004). We reproduce the relevant parts here. If the spin axes of a star and planet are aligned, tidal energy will be dissipated at rates defined by the tidal quality factors \( Q'_s \) and \( Q'_p \). Consider a planet of mass \( M_p \) and radius \( R_p \) orbiting a star of mass \( M_\star \) and radius \( R_\star \). If the system has a semimajor axis \( a \) and angular frequency \( \dot{\omega} \), the rate of change of the eccentricity of the orbit is given by (Eggleton et al. 1998; Mardling & Lin 2002; Dobbs-Dixon et al. 2004)

\[
\dot{e} = g_p + g_\star, \tag{1}
\]

where

\[
g_p = \left( \frac{81}{2} \frac{n e}{Q'_p} \right) \frac{M_p}{M_\star} \left( \frac{R_p}{a} \right)^5 \left[ -f_1(e) + \frac{11}{18} f_2(e) \right], \tag{2}
\]

\[
f_1(e) = \left( 1 + 15 \frac{e^2}{4} + 15 \frac{e^4}{8} + \frac{5}{64} e^6 \right) \left( 1 - e^2 \right)^{13/2}, \tag{3}
\]

\[
f_2(e) = \left( 1 + \frac{3}{2} e^2 + \frac{1}{8} e^4 \right) \left( 1 - e^2 \right)^5. \tag{4}
\]

The stellar and planetary spins can also play an important role in the tidal evolution of close-in planets. If the rotation axes are aligned and if \( M_\star \gg M_p \), the rate of change of the stellar spin frequency is (Mardling & Lin 2002; Dobbs-Dixon et al. 2004)

\[
\dot{\Omega}_\star = \frac{9}{2} \frac{n^2 \epsilon^2}{\epsilon_p Q'_p} \frac{M_p}{M_\star} \left( \frac{R_p}{a} \right)^3 \left[ f_1(e) - f_3(e) \left( \frac{\Omega_p}{n} \right) + \dot{\omega}_s \right], \tag{5}
\]

where

\[
f_3(e) = \left( 1 + \frac{15}{2} e^2 + \frac{45}{8} e^4 + \frac{5}{16} e^6 \right) \left( 1 - e^2 \right)^6. \tag{6}
\]

and \( \dot{\omega}_s \) is the change of stellar rotation due to angular momentum loss through a stellar wind, which we will discuss in more detail in a later section. The quantity \( \epsilon_p \) is the fraction of the mass of the star participating in tidal exchange, and \( \epsilon_\star \) is the star’s moment of inertia coefficient. We use \( \epsilon_p = 0.1 \), appropriate for G stars which have shallow convection zones, and \( \epsilon_\star = 0.1 \).

The rate of change of the planet’s spin is

\[
\dot{\omega}_p = \frac{9}{2} \frac{n^2 \epsilon_p \epsilon_\star Q'_p}{\epsilon_p Q'_p} \frac{M_p}{M_\star} \left( \frac{R_p}{a} \right)^3 \left[ f_1(e) - f_3(e) \left( \frac{\Omega_p}{n} \right) \right]. \tag{8}
\]

The quantities \( \epsilon_p \) and \( \epsilon_\star \) are the fraction of the planet’s mass involved in tidal exchange, and the planet’s moment of inertia coefficient. Since gas giant planets can be fully convective, we use \( \epsilon_p = 1 \) and use \( \epsilon_\star = 0.2 \).

The total angular momentum perpendicular to the orbit is

\[
J_{\omega} = J_\omega(a, e) + J_p(\Omega_p) + J_\star(\Omega_\star). \tag{9}
\]

The orbital angular momentum is

\[
J_\omega(a, e) = M_p \alpha_e M_\star \left( 1 - e^2 \right)^{1/2} (M_p + M_\star), \tag{10}
\]

while the angular momenta of the star and planet are

\[
J_\star = \alpha_\star \epsilon_\star M_\star R_\star^2 \Omega_\star, \tag{11}
\]

\[
J_p = \epsilon_p M_p R_p^2 \Omega_p, \tag{12}
\]

where it is assumed that only a fraction \( \epsilon_\star \) of the star, but all of the planet, is involved in angular momentum exchange. The rate of change of the total angular momentum is then equal to the rate at which the system loses angular momentum through the stellar wind, \( J_{\omega} \). Differentiating equation (9) and using \( J_{\omega} = \alpha \epsilon e M_\star R_\star^2 \dot{\omega}_s \)

\[
\dot{\alpha} \epsilon e M_\star R_\star^2 \dot{\omega}_s = J_\omega(a, e) \left( \frac{\alpha}{2\alpha} \frac{\epsilon e}{1 - e^2} \right) + \alpha \epsilon M_p R_p^2 \dot{\omega}_p + \alpha \epsilon \epsilon_\star M_\star R_\star^2 \dot{\omega}_s. \tag{11}
\]

The rate of change of the semi-major axis of the orbit can then be determined by simply rearranging equation (11).

2.2 Stellar wind model

We take the stellar wind model from Collier Cameron & Jianke (1994). In the unsaturated dynamo regime with purely thermal driving, the rate of change of a star’s angular frequency is

\[
\dot{\omega}_s = -\kappa \Omega^2, \tag{12}
\]

where \( \kappa \) is the rate of energy dissipation by the stellar wind.
where

\[
\kappa = \frac{2}{3} B_\odot^2 \left( \frac{\tau_\odot}{\tau_\bullet \Omega_\bullet} \right)^2 \left( \frac{\beta m_p}{2 \kappa_\bullet T_\bullet} \right)^{1/2} \frac{R_\odot^5}{k^2 M_\odot}. \tag{13}
\]

For the solar values, we assume a magnetic flux density of \(B_\odot = 3\) G, a convective turnover time of \(\tau_\odot = 8.9 \times 10^5\) s and an angular velocity of \(\Omega_\odot = 4.0 \times 10^6\) rad s\(^{-1}\). To determine \(\tau_\bullet\), we assume a linear fit to the values in table 1 of Collier Cameron & Jianke (1994). The quantities \(m_p\) and \(k_\bullet\) are the proton mass and Boltzmann’s constant and from Mestel & Spruit (1987) we assume \(\beta = 0.16\). The final quantity \(k\) is the effective radius of gyration of the radiative interior and convective envelope combined, which we take to be \(k^2 = 0.1\) (Collier Cameron & Jianke 1994).

In the case of stars with saturated dynamos,

\[
\dot{\Omega}_\bullet = -k \tilde{\Omega}^2 \Omega_\bullet, \tag{14}
\]

where \(\tilde{\Omega}\) is the saturation limit which we get by assuming a linear fit to the values in table 5 of Collier Cameron & Jianke (1994).

### 2.3 Basic setup

We carry out Monte Carlo simulations of a large sample of close-in planets (\(a < 0.3\) au) around solar-like stars. We consider a range of different initial radial distributions and eccentricity distributions that we will discuss in more detail in a later section. There are, however, some common elements to all of the simulations that we discuss here.

Based on the properties of the known exoplanets, we distribute the planet masses as \(M_p \propto M^{-1.15}\) (Marcy et al. 2008) with a minimum planet mass of \(M_p = 0.1\, M_{\text{Jup}}\) and a maximum planet mass of \(M_p = 15\, M_{\text{Jup}}\). The minimum and maximum masses are chosen such as to represent the range of planet masses that could undergo Type II migration (D’Angelo, Kley & Henning 2003). The planet radii are computed using \(R \propto M^{3/4}\) where \(m\) is varied from \(m = 0\) for low-mass planets to \(m = 1.5\) for the most massive gas giants, to give a mass–radius relation that matches the expected theoretically (Chabrier et al. 2009) and that is normalized to give \(R_p = 1\, R_{\text{Jupiter}}\) when \(M_p = 1\, M_{\text{Jupiter}}\). The stars are assumed to have masses evenly distributed between 0.75 \(M_\odot\) and 1.25 \(M_\odot\) and have rotation periods distributed as a Gaussian with a peak at 8 d and with a half-width of 1.5 d. The planets are all assumed to start with rotation rates 10 times greater than their orbital periods (i.e. \(\Omega_p = 10\omega\)).

In each simulation we consider 1800 planets. The initial semi-major axis, eccentricity, planet mass, stellar mass and rotational period are all chosen randomly, but in such a way that the final distributions match those described above (we will discuss the chosen semi-major axis and eccentricity distributions below). The star and planet radii are then determined as above and the star mass and rotation periods determine – using equations (12) and (13) – the rate at which the star loses angular momentum through a stellar wind. For the star and planet, we use the tidal quality factors of \(Q_* = 10^6\) and \(Q_p = 5 \times 10^9\). These are similar to those used by Jackson, Barnes & Greenberg (2009) and Dobbs-Dixon et al. (2004) and consistent with those determined by Brown et al. (2011). We did try using the tidal quality factors estimated by Hansen (2010), but these resulted in circularization time-scale for planets with 4 and 5 d periods that were much longer than expected.

Equations (1), (5), (8) and (11) – with equation (11) rewritten in terms of \(\dot{a}\) – are then integrated using a fourth-order Runge Kutta integrator for a time chosen randomly between \(2 \times 10^6\) and \(6 \times 10^9\) yr. Any planet that reaches its Roche limit, given by \(R_\text{Roche} = (R_p/0.462)(M_\ast/M_p)^{1/3}\) (Faber, Rasio & Willems 2005), is assumed to be tidally disrupted and destroyed (Jackson et al. 2009).

### 3 RESULTS

We assume here that close-in giant planets are unable to form in situ, but rather form beyond the snowline (\(a \geq 2.7\) au for a Solar-mass star) and migrate inwards to their current locations (Lin, Bodenheimer & Richardson 1996; Trilling et al. 1998). Although there are a number of mechanisms that could result in planet migration, in the case of giant planets the most likely mechanism is Type II migration in which the planet migrates within a gap in the protoplanetary disk (Goldreich & Tremaine 1980; Lin & Papaloizou 1986).

Using theoretical models of Type II migration, Armitage (2007) (following on from Armitage et al. 2002) has shown that the resulting radial distribution of extrasolar planets is consistent with that observed, the best fit occurring if migration is assumed to be slightly suppressed at small radii. Although this ignores some effects, in particular planet–planet scattering which can also change the orbital elements of surviving planets (Ford et al. 2001), it is at least a reasonable starting point.

Inside 1 au, the best-fitting model of Armitage (2007) has a radial distribution of \(dN/d\log a \propto a^{-0.4}\). We therefore use this, shown in Fig. 1, as the base distribution for all of our simulations. We consider a few different initial eccentricity distributions, but since the final radial distributions do not depend strongly on the initial eccentricity distribution (Jackson et al. 2009), most of our simulations use the initial eccentricity distribution shown in Fig. 2.

#### 3.1 Type II migration only

Our first set of simulations consider the case of Type II migration with slight suppression at small radii (Armitage 2007). We therefore use the initial distribution shown in Fig. 1 and integrate each system for a randomly chosen time of between 2 and 6 Gyr. The final radial distribution is shown in Fig. 3. This illustrates that the tidal interaction causes most of the planets inside 0.04 au to migrate inwards ultimately over flowing their Roche lobes and being destroyed (Jackson et al. 2009). Outside 0.04 au, however, the distribution is not significantly different to the initial distribution.

![Figure 1. Initial radial distribution determined by assuming Type II inward migration with migration slightly suppressed at small radii (Armitage 2007). Also shown is the curve dN/d log a ∝ a^{-0.4} for comparison.](http://example.com/figure1.png)
likely to undergo Type II migration, we only include exoplanets for close-in exoplanets. Since we are only interested in those that induce a stellar radial velocity of less than $2 \, \text{m s}^{-1}$, we have also excluded any planets in our simulated sample that would have also been excluded using the Doppler radial velocity technique (Mayor & Queloz 2007), and will truncate the disc close to the corotation radius. Planets will continue to migrate inside this cavity, but this is expected to slow and become very inefficient once planets pass inside the 2:1 resonance with the inner edge of the disc (Lin et al. 1996; Kuchner & Lecar 2002; Rice & Armitage 2008). The distribution of stellar rotation periods (Herbst et al. 2007), with a peak at $\sim 8 \, \text{d}$, would therefore produce a pile-up of planets with orbital periods $\sim 4 \, \text{d}$, corresponding to an orbital radius of $\sim 0.05 \, \text{au}$ around a solar-like star.

As discussed above, Armitage (2007) suggests that Type II migration would lead to a radial distribution, within 1 au, approximated by $dN/d \log a \propto a^{\alpha}$. This distribution would suggest that $\sim 55$ per cent of the planets between $r = 0.01$ and $0.3 \, \text{au}$ would be located inside $0.1 \, \text{au}$. We therefore consider three different pile-up scenarios, illustrated in Fig. 5. The solid line is a case in which the number of planets between $r = 0.01$ and $0.1 \, \text{au}$ is the same as would be expected from the radial distribution derived by Armitage (2007). The dash–dotted line increases the fraction to 65 per cent, and the dash–dot–dot line is a case in which the fraction inside $0.1 \, \text{au}$ is enhanced to $\sim 88$ per cent.

As before, we evolve each system for a randomly chosen time of between 2 and 6 Gyr. The final radial distributions are shown in Fig. 6. The thick dashed line is again the radial distribution of the known exoplanets and again we have only included those discovered using the Doppler radial velocity technique and that orbit stars with masses between 0.75 and $1.25 \, \text{M}_\odot$. We have also removed any simulated planet that would induce a radial velocity of less than $2 \, \text{m s}^{-1}$. The line styles in Fig. 6 correspond to those shown in Fig. 5. In all three cases, the pile-up remains but the best fit with the observed distribution is somewhere between the solid and dash–dotted line. The cumulative distributions are shown in Fig. 7. The line styles again correspond to those in Figs 5 and 6. Again this illustrates that the best fit to the observed distribution would be somewhere

The dashed line in Fig. 3 shows the radial distribution of known close-in exoplanets. Since we are only interested in those that are likely to undergo Type II migration, we only include exoplanets for which $M \geq 0.1 \, \text{M}_\text{Jup}$. We have included only those planets discovered using the Doppler radial velocity technique (Mayor & Queloz 1995; Udry, Fischer & Queloz 2007). We considered also including planets detected via transits, but were concerned that selection effects may somewhat enhance any pile-up at small radii. So that our simulated sample and the observed sample can be compared, we have also excluded any planets in our simulated sample that would induce a stellar radial velocity of less than $2 \, \text{m s}^{-1}$, and only included those observed exoplanets whose host star masses are between 0.75 and $1.25 \, \text{M}_\odot$. Fig. 3 suggests that these two distributions are not the same, with some evidence for a pile-up of observed exoplanets at around $0.05 \, \text{au}$, that is not seen in the simulated distribution. Fig. 4 shows the cumulative distribution for the two samples and again illustrates that a larger fraction of the observed exoplanets are located inside $0.1 \, \text{au}$, compared to the simulated systems. Formally a Kolmogorov–Smirnov test shows that the probability that the two distributions are drawn from the same parent distribution is $P_{\text{KS}} = 0.36$. 

### Figure 2
Initial eccentricity distribution used in the simulations presented. The final radial distributions do not, however, depend particularly strongly on the initial eccentricity distribution.

### Figure 3
The final semi-major axis distribution for the simulated planets (solid line) compared to the semi-major axis distribution for observed exoplanets (dashed line) in systems with properties comparable to those considered in the simulations.

### Figure 4
Cumulative semi-major axis distribution for the simulated exoplanets (solid line) and for the observed exoplanets (dashed line), showing that the radial distribution of the observed exoplanets does not appear to be consistent with the simulated systems that are taken to have an initial radial distribution resulting from Type II migration alone.

#### 3.2 Type II migration with stopping mechanism

It has been suggested that rather than migrating all the way into the central star, planets may be prevented from doing so by some kind of stopping mechanism that halts planets with periods of $\sim 4 \, \text{d}$, corresponding to a semimajor axis of $\sim 0.05 \, \text{au}$. One possibility is that the magnetic fields of T Tauri stars may be strong enough to disrupt the inner edge of protoplanetary discs (Königl 1991; Bouvier et al. 2007), and will truncate the disc close to the corotation radius. Planets will continue to migrate inside this cavity, but this is expected to slow and become very inefficient once planets pass inside the 2:1 resonance with the inner edge of the disc (Lin et al. 1996; Kuchner & Lecar 2002; Rice & Armitage 2008). The distribution of stellar rotation periods (Herbst et al. 2007), with a peak at $\sim 8 \, \text{d}$, would therefore produce a pile-up of planets with orbital periods $\sim 4 \, \text{d}$, corresponding to an orbital radius of $\sim 0.05 \, \text{au}$ around a solar-like star.

As discussed above, Armitage (2007) suggests that Type II migration would lead to a radial distribution, within 1 au, approximated by $dN/d \log a \propto a^{\alpha}$. This distribution would suggest that $\sim 55$ per cent of the planets between $r = 0.01$ and $0.3 \, \text{au}$ would be located inside $0.1 \, \text{au}$. We therefore consider three different pile-up scenarios, illustrated in Fig. 5. The solid line is a case in which the number of planets between $r = 0.01$ and $0.1 \, \text{au}$ is the same as would be expected from the radial distribution derived by Armitage (2007). The dash–dotted line increases the fraction to 65 per cent, and the dash–dot–dot line is a case in which the fraction inside $0.1 \, \text{au}$ is enhanced to $\sim 88$ per cent.

As before, we evolve each system for a randomly chosen time of between 2 and 6 Gyr. The final radial distributions are shown in Fig. 6. The thick dashed line is again the radial distribution of the known exoplanets and again we have only included those discovered using the Doppler radial velocity technique and that orbit stars with masses between 0.75 and $1.25 \, \text{M}_\odot$. We have also removed any simulated planet that would induce a radial velocity of less than $2 \, \text{m s}^{-1}$. The line styles in Fig. 6 correspond to those shown in Fig. 5. In all three cases, the pile-up remains but the best fit with the observed distribution is somewhere between the solid and dash–dotted line. The cumulative distributions are shown in Fig. 7. The line styles again correspond to those in Figs 5 and 6. Again this illustrates that the best fit to the observed distribution would be somewhere
Tidal evolution of close-in planets

Figure 5. The initial radial distribution of planets assuming Type II inward migration followed by a pile-up of planets inside 0.1 au with the peak of the pile-up occurring where the orbital period is half the stellar spin period (\(a \sim 0.05\) au), and a width corresponding to a variation in the orbital period of \(\pm 2\) d. The three cases shown are one in which the number of planets inside 0.1 is the same as in Fig. 1 (solid line), an enhancement of 20 per cent over that in Fig. 1 (dash–dotted line) and a further enhancement of 30 per cent (dotted line).

Figure 6. The final radial distribution of the simulated systems with initial radial distributions as shown in Fig. 5 (solid, dash–dotted and dotted lines) compared to the observed radial distribution of exoplanets (thick dashed line). The above result suggests that, to explain the current distribution of close-in giant planets, there must have been a primordial pile-up of giant planets with the peak of the pile-up occurring at \(\sim 0.05\) au. The best fit also occurs if the number of planets between 0.01 and 0.1 is similar to, but slightly higher than, that expected based on the radial distribution of planets resulting from Type II migration (Armitage 2007). One has to be slightly careful with this comparison, however, as the radial profile determined by Armitage (2007) is a steady-state profile and does not indicate what fraction of planets will have migrated inside 0.1 au and then been lost. Population synthesis models (Ida & Lin 2004) suggest that a large fraction of giant planets should migrate close to their host stars. If the stopping mechanism were efficient, we would then expect a significant primordial pile-up at small radii. Figs 6 and 7 are consistent with a pile-up but are not consistent with a particularly significant pile-up. This suggests either that the stopping mechanism is not efficient and that planets are lost before tidal evolution becomes important, or that a smaller fraction of giant planets migrate close to their parent than predicted by population synthesis models.

3.3 Mass and eccentricity dependence

In Fig. 6, we consider the entire mass range from \(M_p = 0.1\) M\(_{\text{Jup}}\) to a maximum of \(M_p = 15\) M\(_{\text{Jup}}\). If, however, we divide this into a low-mass regime (\(M_p < 2\) M\(_{\text{Jup}}\)) and a high-mass regime (\(M_p > 2\) M\(_{\text{Jup}}\)), we find (Fig. 8) that for the low-mass planets the peak at \(a \sim 0.05\) au remains and is still well matched for our simulated populations that have modest pile-ups at this radius. For the high-mass planets (Fig. 9), this is no longer the case. In the observed population there are no planets with \(a \sim 0.05\) au.

Figure 7. Cumulative semi-major axis distribution for the simulated systems (solid, dash–dotted and dotted lines) with initial radial distributions shown in Fig. 5, compared to the cumulative radial distribution of the observed exoplanets (thick dashed line). This appears to illustrate that the observed distribution is consistent with a primordial pile-up at \(r \sim 0.05\) au but not if there is a significant enhancement in the fraction of planets inside 0.1 au.

Figure 8. The final radial distribution of the simulated systems with initial radial distributions as shown in Fig. 5 (solid, dash–dotted and dotted lines) and with masses \(M_p < 2\) M\(_{\text{Jup}}\), compared to the observed radial distribution of exoplanets with the same range of masses (thick dashed line).
It has been suggested (Rice & Armitage 2008) that the interaction between a planet, located inside a magnetospheric cavity, and the surrounding disc can lead to substantial eccentricity growth for planets with $M > 1 M_{\text{Jup}}$ and that the growth rate is greatest for the highest mass planets. It was therefore suggested (Rice & Armitage 2008) that this would result in higher mass, close-in planets being preferentially destroyed when compared with lower mass, close-in planets. We therefore consider a situation in which the initial radial distribution is as shown in Fig. 5, but the initial eccentricities increase with increasing mass. We assume, somewhat arbitrarily, the planets with masses below $1 M_{\text{Jup}}$ have an eccentricity peak close to $e = 0$, but have a tail that can extend to $e = 0.6$, while planets with masses above $10 M_{\text{Jup}}$ have eccentricities that lie preferential between $e = 0.4$ and $0.8$. We assume a maximum eccentricity of $e = 0.8$ so that we do not simply lose a large fraction of the massive exoplanets immediately. We should note that simulations by Benítez-Llambay, Masset & Beaugé (2011) did not find eccentricity growth for massive planets in inner disc gaps. Their simulations, however, did not have completely evacuated inner cavities and so differ from those of Rice & Armitage (2008).

Fig. 10 shows the resulting radial distribution for the lower-mass planets ($M_p < 2 M_{\text{Jup}}$ – solid line) which still agrees well with the observed radial distribution (thick dashed line). The radial distribution for the higher mass planets ($M_p > 2 M_{\text{Jup}}$ – solid line) is shown in Fig. 11, and although not a particularly good fit to the observed distribution (thick dashed line) does at least indicate that the pile-up is significantly depleted. Furthermore, if one considers the final eccentricity distribution, shown in Fig. 13, again divided into planets with masses above $2 M_{\text{Jup}}$ (solid line) and planets with masses below $2 M_{\text{Jup}}$ (dashed line), a group of high-mass exoplanets with large eccentricities remains and is at least qualitatively similar to that observed (Fig. 12). This suggests that if higher mass exoplanets do indeed preferentially have higher eccentricities than lower mass exoplanets, this could explain why the pile-up at $a \sim 0.05$ au is only observed for the lower mass exoplanets and suggests that the mass-dependent eccentricity distribution must be primordial rather than being due to a difference in their subsequent evolution (Rice & Armitage 2008).

**3.4 Planet–planet scattering**

The main goal of this work has been to show that an initial distribution consistent with that expected from Type II migration (Armitage 2007) followed by a pile-up due to truncation of the inner disc (Rice & Armitage 2008) is at least consistent – after tidal evolution – with the observed distribution of close-in exoplanets. We cannot, however, preclude the possibility that migration occurs through other mechanisms.

Wright et al. (2009) have shown that there is a significant difference in the semi-major axis distribution of single- and multiple-planet systems. The single-planet systems have a pile-up at $a \sim 3$ au and a radial distribution that increases with increasing radius. The multiple-planet systems, on the other hand, have a semi-major axis...
distribution that is only weakly dependent on radius. It has been suggested (e.g. Matsumura et al. 2010) that this difference could be due to single-planet systems being dominated by planet–planet scattering (with additional planets being ejected from the system) while multiple-planet systems are dominated by migration in a gas disc. Type II migration would, according to Armitage (2007), result in a semi-major axis distribution that increases significantly with increasing radius. Qualitatively, this is more consistent with the radial distribution of the single-planet systems than it is with the multiple-planet systems, which have a semi-major axis distribution that is only weakly dependent on radius.

It is difficult to predict what kind of semi-major axis and eccentricity distribution planet–planet scattering would produce, but we assume that if the distribution of single-planet systems is dominated by planet–planet scattering, and if the additional planets are ejected or are now beyond ~5 au, the eccentricity of the close-in planets must be large. We consider an initial eccentricity distribution that peaks at $e \sim 0.6$ and that has a half-width of 0.1.

We consider three different initial radial distributions shown in Fig. 1, and two from Fig. 5 (solid line and dotted line). We integrate these systems in the same way as before. The results are shown in Fig. 14. The initially unpeaked distribution is shown as the solid line. The initially weakly peaked distribution is the dotted line, while the initially strongly peaked distribution is the dash–dotted line. A larger fraction of the close-in planets are destroyed when compared with the simulations in which the initial eccentricity peaked at $e = 0$ and so the initially weakly peaked distribution does not show a peak after tidal evolution. That the most peaked initial distribution (dash–dotted line) has a pile-up that is similar to that observed at least suggests that there could be an initial distribution that could result in a match with the observed distribution. However, tidal evolution also broadens the peak which suggests that the initial peak would need to be narrower than we assume.

Although the above results do not preclude the possibility that the radial distribution of the observed single-planet systems which includes a peak at $\sim 3$ d is due primarily to planet–planet scattering, it does suggest that this would require that planet–planet scattering is very effective at scattering planets into orbits with $a < 0.1$ au which is not entirely consistent with theoretical expectations (Ford et al. 2001; Marzari & Weidenschilling 2002). Even Nagasawa et al. (2008), who showed that the Kozai mechanism can significantly enhance the number of close-in planets, conclude that the main channel for forming close-in giant planets is probably still through Type II migration. Furthermore, in these simulations, the planets with $a > 0.1$ au undergo very little eccentricity evolution.

The observed sample of exoplanets has an eccentricity distribution that peaks at $e = 0$ (Wright et al. 2009) suggests that the initial eccentricity distribution, at least for those with $a > 0.1$, must have also peaked at $e = 0$ which may suggest that the primary mechanism for getting these planets close to their parent stars cannot be planet–planet scattering unless something else (other than tidal evolution) then reduces their initial eccentricities.
3.5 Mass–period relation

It has been suggested that the observed mass–period relation of close-in planets may give us some information about their evolution. Ford & Rasio (2006) suggest that if planets evolve from very high initial eccentricities, their final orbital distance will be about twice the Roche distance (i.e. $a \sim 2 a_R$). Pont et al. (2011) suggest – by studying the mass–period relation of transiting planets – that tidal circularization and the stopping mechanism for close-in planets must be closely related. Our results here suggest that close-in planets are not typically circularized from highly eccentric orbits and also that any stopping mechanism occurs prior to tidal evolution playing a significant role.

The initial semi-major axis distribution that, after tidal circularization, produced the best fit when compared with the observed distribution of close-in planets was one in which there was a modest initial pile-up at $a \sim 0.05$ au (solid line in Fig. 5). In Fig. 15, we plot the planet-to-star mass ratio against orbital period for a randomly selected sample of these planets after they have undergone tidal evolution (asterisks) together with the observed close-in planets detected via radial velocity (triangles). For the modelled systems, we have randomly selected the same number of planets as the number of observed close-in planets. The two diagonal lines are the Roche limit and twice the Roche limit. Qualitatively, the mass–period relation of the modelled systems looks very similar to that of the observed systems. Even though our modelled sample was not tidally evolved from an initially highly eccentric orbit, the inner boundary is just beyond twice the Roche limit. Admittedly, the sample is quite small and we have not analysed this in extensive details. It does, however, appear as though evolution through Type II migration (with a mechanism for stopping planets at $a \sim 0.05$ au) followed by tidal evolution can produce a mass–period relation that is consistent with that observed. It has also been suggested (Davis & Wheatley 2009) that close-in planets can also be lost through evaporation. Although evaporation may well be operating, the lack of planets close to their Roche limit is a natural consequence of inward migration followed by tidal evolution.

![Figure 15](https://academic.oup.com/mnras/article-abstract/425/4/2567/1075954/2574W.K.M.Rice,J.VeljanoskiandA.CollierCameron

4 DISCUSSION AND CONCLUSIONS

The primary goal of this paper was to determine if the current distribution of close-in, gas-giant ($M > 0.1 M_{\text{Jup}}$) exoplanets around Solar-like ($0.75 M_\odot \leq M \leq 1.25 M_\odot$) stars is consistent with an initial distribution resulting primarily from gap-opening migration in a gas disc (Type II) (e.g. Armitage et al. 2002; Armitage 2007). The chosen initial distributions were evolved using tidal evolution equations (Dobbs-Dixon et al. 2004) for a randomly chosen time between $2 \times 10^6$ and $6 \times 10^7$ yr.

The results suggest that an initial radial distribution due to Type II migration alone does not match the currently observed radial distribution which shows a pile-up of planets at $a \sim 0.05$ au. If, however, the inner regions of the disc are truncated due to interactions with the star’s magnetosphere, it has been suggested that planets should pile up when inside the 2:1 resonance with the inner disc edge (Lin et al. 1996; Rice & Armitage 2008) which would correspond to a semi-major axis of $\sim 0.05$ au for Solar-like stars with initial rotation periods of $\sim 8$ d. This study shows that the mass–period–eccentricity relation is indeed consistent with expectations from disc migration with a magnetospheric gap and tides, at least for stars with masses in the range between $0.75 M_\odot$ and $1.25 M_\odot$ and as long as the primordial pile-up is slightly (but not significantly) enhanced compared to what would be expected based on a steady-state model of Type II migration. We also find a good, qualitative, agreement between the mass–period relation of the modelled systems and that of the observed systems.

This pile-up is, however, only evident for the lower mass planets ($M_p < 2 M_{\text{Jup}}$). Simulated systems with a mass-independent initial eccentricity distribution retain an initial pile-up for all planet masses. The pile-up for higher mass planets is, however, significantly reduced if the initial eccentricity distribution is assumed to be mass dependent with the higher mass planets preferentially having higher eccentricities. This is consistent with simulations (Rice & Armitage 2008) suggesting that higher mass planets will have enhanced eccentricity growth inside a magnetospheric cavity.

We should stress that this does not prove that the primary mechanism for producing the observed close-in giant exoplanets is Type II migration but simply shows that – for reasonable assumptions about the initial distributions – it is consistent with this being the case. We also consider one set of simulations with an initial eccentricity distribution that peaks at a large eccentricity, aimed to mimic a possible planet–planet scattering scenario. Although it can result in a final distribution that matches that observed, it requires a more significant initial pile-up when compared to an initial eccentricity distribution that peaks at $e = 0$. Furthermore, the planets beyond $a \sim 0.1$ au retain their high eccentricities which is not consistent with observations.

The observation of misaligned, and in some cases retrograde, close-in planets (Winn et al. 2009; Collier Cameron et al. 2010; Triaud et al. 2010) does, however, strongly suggest that planet–planet scattering (which may undergo ‘Kozai’ cycles if a binary or distant planetary companion is present) must play a role in the formation of some close-in exoplanets. It is intriguing that misaligned systems occur at all masses, but dominate for stars with $T_{\text{eff}} > 6250$ K (Winn et al. 2010; Barnes, Linscott & Shporer 2011). In their population-synthesis studies that include disc migrations, Alibert, Mordasini & Benz (2011) find that the short disc lifetimes of high-mass stars does not leave enough time for Jupiter-mass planets to migrate into close orbits. Our results, therefore, appear to be consistent with this. For stars with masses between $0.75$ and $1.25 M_\odot$, the distribution...
of close-in giant planets can be modelled by assuming Type II migration followed by tidal evolution. There will be some systems that have been influenced by planet–planet scattering, which may have undergone ‘Kozai’ cycles, but these do not dominate. For the more massive stars ($M_\ast > 1.25 M_\odot$), where the disc lifetime is too short for Type II migration to be effective, close-in giant planets are preferentially formed through dynamical interactions in which ‘Kozai’ cycles may also operate.

ACKNOWLEDGMENTS

WKMR acknowledges support from the Scottish Universities Physics Alliance (SUPA) and for support from the Science and Technology Facilities Council (STFC) through grant ST/H002380/1. The authors would also like to thank the Isaac Newton Institute for Mathematical Sciences for their hospitality during the Dynamics of Discs and Planets Programme and would like to thank Phil Armitage for many useful discussions. The Authors would like to thank the anonymous referee for helpful comments that have improved this paper.

REFERENCES

Alibert Y., Mordasini C., Benz W., 2011, A&A, 526, A63
Armitage P. J., 2007, ApJ, 665, 1381
Armitage P. J., Livio M., Lubow S. H., Pringle J. E., 2002, MNRAS, 334, 248
Barnes J. W., Linscott E., Shporer A., 2011, ApJS, 197, 10
Bell K. R., Cassen P. M., Klahr H. H., Henning Th., 1997, ApJ, 486, 372
Benítez-LLamby P., Masset F., Beaugé C., 2011, A&A, 528, A2
Bouvier J., Alencar S. H. P., Harries T. J., Johns-Krull C. M., Romanova M. M., 2007, in Reipurth B., Jewitt D., Keil K., eds, Protostars and Planets V. Univ. Arizona Press, Tucson, p. 479
Brown D. J. A., Collier Cameron A., Hall C., Hebb L., Smalley B., 2011, MNRAS, 415, 605
Chabrier G., Baraffe I., Leconte J., Gallardo J., Barman T., 2009, in Stempels E., ed., AIP Conf. Proc. Vol. 1094, Cool Stars, Stellar Systems and the Sun: 15th Cambridge Workshop on Cool Stars, Stellar Systems and the Sun. AIP, New York, p. 102
Collier Cameron A., Jianke L., 1994, MNRAS, 269, 1099
Collier Cameron A., Bruce V. A., Miller G. R. M., Triaud A. H. M. J., Queloz D., 2010, MNRAS, 403, 151
D’Angelo G., Kley W., Henning T., 2003, ApJ, 586, 540
Davis T. A., Wheatley P. J., 2009, MNRAS, 396, 1012
Dobbs-Dixon I., Lin D. N. C., Mardling R. A., 2004, ApJ, 610, 464
Eggleton P. P., Kiseleva L. G., Hut P., 1998, ApJ, 499, 853
Faber J., Rasio F., Willems B., 2005, Icarus, 175, 248
Fabrycky D., Tremaine S., 2007, ApJ, 669, 1298
Ford E. B., Rasio F. A., 2006, ApJ, 638, L45
Ford E. B., Havlickova M., Rasio F. A., 2001, Icarus, 150, 303
Goldreich P., Tremaine S., 1980, ApJ, 241, 425
Hansen B. M. S., 2010, ApJ, 723, 285
Herbst W., Eisloffel J., Mundt R., Scholz A., 2007, in Reipurth B., Jewitt D., Keil K., eds, Protostars and Planets V. Univ. Arizona Press, Tucson, p. 297
Iida S., Lin D. N. C., 2004, ApJ, 616, 567
Jackson B., Barnes R., Greenberg R., 2009, ApJ, 698, 1357
Königl A., 1991, ApJ, 370, L9
Kozai Y., 1962, AJ, 67, 591
Kuchner M. J., Lecar M., 2002, ApJ, 574, L87
Lin D. N. C., Papaloizou J., 1986, ApJ, 309, 846
Lin D. N. C., Bodenheimer P., Richardson D. C., 1996, Nat, 380, 606
Marcy G. W. et al., 2008, Phys. Scr., 130, 014001
Mardling R. A., Lin D. N. C., 2002, ApJ, 573, 829
Marzari F., Weidenschilling S. J., 2002, Icarus, 156, 570
Matsumura S., Thommes E. W., Chatterjee S., Rasio F. A., 2010, ApJ, 714, 194
Mayor M., Queloz D., 1995, Nat, 378, 355
Mestel L., Spruit H. C., 1987, MNRAS, 226, 57
Morton T. D., Johnson J. A., 2011, ApJ, 729, 138
Nagasawa M., Ida S., Bessho T., 2008, ApJ, 678, 498
Pont F., Husnoo N., Mazeh T., Fabrycky D., 2011, MNRAS, 414, 1278
Queloz D. et al., 2010, A&A, 517, L1
Rasio F., Tout C. A., Lubow S. H., Livio M., 1996, ApJ, 470, 1187
Rice W. K. M., Armitage P. J., 2008, MNRAS, 384, 1242
Triaud A. H. M. J. et al., 2010, A&A, 524, A25
Trilling D. E., Benz W., Guillot T., Lunine J., Hubbard W. B., Burrows A., 1998, ApJ, 500, 428
Udry S., Fischer D., Queloz D., 2007, in Reipurth B., Jewitt D., Keil K., eds, Protostars and Planets V. Univ. Arizona Press, Tucson, p. 685
Ward W. R., 1997, Icarus, 126, 261
Winn J. N. et al., 2009, ApJ, 703, 2091
Winn J. N., Fabrycky D., Albrecht S., Johnson J. A., 2010, ApJ, 718, L145
Wright J. T., Upadhyay S., Marcy G. W., Fischer D. A., Ford E. B., Asher Johnson J., 2009, ApJ, 693, 1084
Wu Y., Murray N., 2003, ApJ, 589, 605

This paper has been typeset from a TeX/LaTeX file prepared by the author.