Planetary Migration: What Does It Mean for Planet Formation?

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Key Words

extrasolar planet, protoplanetary disk, accretion, hot Jupiter, resonance

Abstract

Gravitational interactions between a planet and its protoplanetary disk change the planet's orbit, causing the planet to migrate toward or away from its star. Migration rates are poorly constrained for low-mass bodies but reasonably well understood for giant planets. In both cases, significant migration will affect the details and efficiency of planet formation. If the disk is turbulent, density fluctuations will excite orbital eccentricities and cause orbits to undergo a random walk. Both processes are probably detrimental to planet formation. Planets that form early in the lifetime of a disk are likely to be lost, whereas late-forming planets will survive and may undergo little migration. Migration can explain the observed orbits and masses of extrasolar planets if giants form at different times and over a range of distances. Migration can also explain the existence of planets orbiting close to their star and of resonant pairs of planets.

1. INTRODUCTION

AU: astronomical unit

MMR: mean-motion resonance

A planet embedded in a disk of gas or small particles modifies the distribution of material in its vicinity, causing the planet's orbit to change. Gravitational interaction with a disk can excite or damp a planet's orbital eccentricity and can also alter the size of the planet's orbit, causing the planet to migrate toward or away from its star. These changes can occur rapidly on timescales much shorter than the time required to form a planet. The phenomenon of planetary migration has been known for decades (Goldreich & Tremaine 1980, Lin & Papaloizou 1986), yet for a long time migration received little attention from the planetary science community, which was intent on understanding the formation of the Solar System, for which there seems little need to invoke migration.

The discovery of extrasolar planets has changed all this. At the time of writing, several hundred planets are known, providing a large enough sample to test theories of planet formation statistically. Several aspects of the mass and orbital distribution of extrasolar planets (see **Figure 1**) are difficult to explain by using the standard model of planet formation developed for the Sun's planets. The existence of hot Jupiters orbiting <0.1 AU (astronomical units) from their star (Mayor & Queloz 1995), pairs of planets locked in mean-motion resonances (MMRs) (Laughlin & Chambers 2001), the gradual increase in the frequency and masses of giant planets with orbital distance (Cumming et al. 2008), and the sheer variety of extrasolar systems are all more readily explained when migration is accounted for.

In this article, I describe how the standard model of planet formation is being modified to include planetary migration—a transformation that is by no means complete at present. This review focuses on migration within a gaseous protoplanetary disk. Planetary orbits may also change once the gas has dispersed, owing to interactions with a debris disk of solid particles (Tsiganis et al.

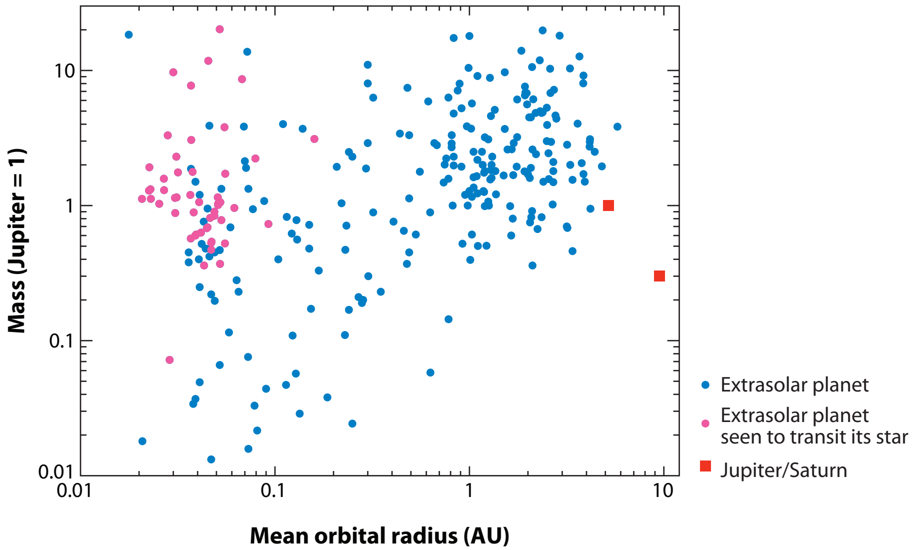

Figure 1

The masses and mean orbital distances of known extrasolar planets. The lack of low-mass planets beyond 1 AU (astronomical unit) is an observational selection effect, but the increased frequency of planets at large distances is real (Cumming et al. 2008). When the biased transiting population is removed, there is a deficit of massive planets orbiting closer than 1 AU (Cumming et al. 2008). Data were taken from the Extrasolar Planets Encyclopedia at **http://exoplanet.eu.**

2005) or to close encounters between pairs of planets (Ford et al. 2005). These topics are not considered here. In the following section, I briefly review the standard model of planet formation in the absence of migration. Section 3 describes our current understanding of how a planet interacts with a gaseous disk and how these interactions alter the planet's orbit. In Section 4, I discuss how the formation of terrestrial and giant planets is likely to change when migration is a factor.

Planetesimal: one of the first large bodies to form from dust grains in a protoplanetary disk, roughly 1–100 km in radius

$M_\odot$: solar mass

Myr: million years

MRI: magnetorotational instability

Dead zone: region of the disk that does not couple to the local magnetic field because the ionization level is too low

2. THE STANDARD MODEL OF PLANET FORMATION

In this section, I briefly describe the standard model for planet formation in the Solar System, based on the planetesimal hypothesis (Safronov 1969). The standard model is traditionally broken down into a series of chronological steps. The evolution in a given region of the disk progresses through these stages in order, but different stages may take place concurrently in different regions of the disk. For more detailed reviews of planet formation, see Armitage 2007a, Chambers 2004, Lissauer & Stevenson 2007, and Nagasawa et al. 2007.

2.1. Protoplanetary Disks

Most stars less than a few million years old are surrounded by disks of gas and dust. These disks are widely believed to be the sites of planet formation because they contain all the raw materials needed to build planets and their shape reflects the roughly coplanar arrangement of material in the sun's planetary system. A typical disk has a mass of 0.01–0.1$M_\odot$ and a radius of a few hundred astronomical units (Andrews & Williams 2007). Strong optical and UV emission lines in the spectra of young stars suggest that they are accreting gas from their disks at rates of 10^{-9} to $10^{-7} M_\odot$ year^{-1} for solar-mass stars (Gullbring et al. 1998). This accretion and the absence of disks around stars older than ~10 Myr (million years) (Haisch et al. 2001, Pascucci et al. 2006) suggest that disks evolve viscously over time, with most of the material accreting onto the star. Toward the end of their lives, disks disappear rapidly (Hartmann et al. 2005), probably owing to the presence of a giant planet (Varnière et al. 2006) or to a combination of viscous accretion and photoevaporation by UV radiation from the star (Alexander et al. 2006).

Viscous accretion is generally assumed to involve turbulence, although the origin of this turbulence is uncertain. A leading candidate is magnetorotational instability (MRI), in which the disk's magnetic field couples to electrons and ions in the differentially rotating disk (Hawley & Balbus 1991). MRI requires a minimum level of ionization to operate. This level is probably exceeded close to a star, owing to thermal ionization, and in the outer disk where X-rays and cosmic rays can penetrate to the disk midplane. At intermediate distances, a magnetic dead zone may exist, reducing turbulence and viscous accretion near the midplane (Matsumura & Pudritz 2006). Ionization levels remain high away from the midplane, forming a layered accretion disk (Gammie 1996). The radial extent of any dead zone is currently unclear (Matsumura & Pudritz 2006, Turner et al. 2007). Recent work suggests that MRI turbulence may not operate at the midplane in much of the region where planets form, but interaction with the magnetic field can continue to drive accretion in a laminar region of the disk, dubbed an undead zone (Turner & Sano 2008).

Disk accretion is often parameterized using an α model. Here it is assumed that the largest turbulent eddies will not exceed the vertical scale height $H_{\rm gas}$ of the gas, and the eddies will not circulate faster than the local sound speed c_s. In this case, the viscosity ν can be expressed as

$$\nu = \alpha c_s H_{\rm gas}, \tag{1}$$

where $\alpha \leq 1$ (Shakura & Syunyaev 1973). Observed stellar mass accretion rates suggest that typically $\alpha \sim 0.01$–0.001 (Hartmann et al. 1998, Hueso & Guillot 2005).

Disks are heated by radiation from the central star and the release of gravitational energy as material flows inward through the disk. Both effects decline with distance from the star, so temperature decreases radially outward in the disk. Gas pressure and density also generally decline with distance.

MMSN: minimum-mass solar nebula

Micrometer-sized dust grains make up ~1% of the initial mass in a disk with a composition similar to the Sun, whereas the remaining 99% is gas. In the hot inner disk, the dust is composed of refractory materials: silicates, oxides, sulfides, and metal grains. Further from the star, beyond the snow line, solids are augmented by water ice, with more volatile ices present in the coolest regions. The snow line moves inward over time as viscous accretion declines and the disk cools. Shortly after the disk forms, the snow line lies at 4–6 AU for a solar-mass star (Boss 1996, Kennedy & Kenyon 2008), moving inward to 1–2 AU just before the disk disperses (Garaud & Lin 2007, Kennedy & Kenyon 2008, Lecar et al. 2006). The increase in the amount of solid material at the snow line is somewhat uncertain. Conventional models for the sun's protoplanetary disk (the solar nebula) assume an increase by a factor of approximately four (Hayashi 1981). The increase is only a factor of two if highly volatile ices like methane and ammonia are neglected (Lodders 2003). However, the increase may be >4 if turbulent diffusion transports water vapor from the inner disk out across the snow line, allowing the water vapor to condense (Ciesla & Cuzzi 2006, Stevenson & Lunine 1988).

Many models of planet formation make use of an idealized disk called the minimum-mass solar nebula (MMSN), which is based on attempts to determine the original distribution of mass in the solar system. The MMSN has a mass of 0.01–0.02$M_\odot$ and a surface density that varies as $1/a^{3/2}$, where a is heliocentric distance (Hayashi 1981, Weidenschilling 1977b).

2.2. Formation of Planetesimals

The first stage of planet formation involves the aggregation of micrometer-sized dust grains into 1–100-km-sized bodies referred to as planetesimals. Several planetesimal formation mechanisms have been proposed, although it is unclear which occurs in practice.

Dust grains can collide and stick together electrostatically to form aggregates (Krause & Blum 2004), with further growth occurring as small aggregates embed themselves in larger ones (Wurm et al. 2005). Numerical simulations suggest that 10-km bodies can grow this way in $<10^4$ orbital periods (Weidenschilling 1997), but only in the absence of strong global turbulence (Weidenschilling 1984). These short growth times are at odds with meteoritical data that indicate that many meteorite parent bodies required several million years to form (Scott 2007).

In the absence of turbulence, small particles tend to sediment to a thin layer at the midplane of the disk. This layer will become gravitationally unstable and fragment to form bound clumps if the height of the layer $d \lesssim \Sigma_{gas} a^3/2M_*$, where Σ_{gas} is the gas surface density, and the clump is a distance a from a star of mass M_* (Goldreich & Ward 1973, Safronov 1969, Wetherill 1980). Sedimentation of particles to the center of a clump will form a planetesimal ~10 km in radius at 1 AU, with larger bodies forming further out in the disk. However, planetesimals can form this way only in a laminar disk. An unstable layer is unlikely to form if turbulence ($\alpha \gtrsim 10^{-8}$) is present (Cuzzi & Weidenschilling 2006).

The concentration of meter-sized objects at local pressure maxima via gas drag (Johansen et al. 2007), and the concentration of millimeter-sized particles between the smallest turbulent eddies (Cuzzi et al. 2008), provide two ways to generate gravitationally bound clumps in a turbulent disk. The latter mechanism is particularly promising because its low efficiency explains why meteorite parent bodies took so long to form and why they are largely composed of millimeter-sized chondrules (Cuzzi et al. 2001, 2008).

2.3. Runaway and Oligarchic Growth

A population of planetesimals evolves as a result of numerous close encounters that change the orbital distribution and as a result of occasional collisions that lead to growth. A planetesimal of radius R and mass M sweeps up smaller objects at a rate

$$\frac{dM}{dt} \simeq \frac{\pi R^2 v_{\rm rel} \Sigma_{\rm solid}}{2 H_{\rm solid}} F_{\rm grav}, \tag{2}$$

where the planetesimals have a surface density $\Sigma_{\rm solid}$, have a velocity dispersion $v_{\rm rel}$, and move in a disk with a scale height $H_{\rm solid} \propto v_{\rm rel}$ (Lissauer 1987). The collision rate is increased by a factor $F_{\rm grav} \simeq 1 + (v_{\rm esc}/v_{\rm rel})^2$, as gravity focuses the trajectories of passing planetesimals toward one another, where $v_{\rm esc}$ is the mutual escape velocity. Encounters tend to increase relative velocities, whereas gas drag reduces $v_{\rm rel}$. Gas drag also causes planetesimals to drift slowly toward the star at rates that are inversely proportional to their radius (Weidenschilling 1977a).

$M_\oplus$: Earth mass

Oligarchic growth: growth stage in which each region of a disk contains a single planetary embryo and many planetesimals

Planetary embryo: a body formed by oligarchic growth that dominates the size distribution in a region of the disk

Over the course of many encounters, large planetesimals acquire small $v_{\rm rel}$, and vice versa, an effect commonly referred to as dynamical friction. The combination of dynamical friction and gravitational focusing naturally leads to runaway growth, in which the largest objects grow the most rapidly, forming bodies >100 km in radius in $\sim 10^4$ orbital periods after planetesimals appear in large numbers (Wetherill & Stewart 1993).

During runaway growth, $v_{\rm rel}$ is determined by the mean mass of the planetesimals, most of which remain small. Once the largest bodies reach 10^{-6}–$10^{-5} M_\oplus$, growth enters a slower regime, called oligarchic growth, in which $v_{\rm rel}$ is determined by gravitational interactions with the largest objects (Ida & Makino 1993, Kokubo & Ida 1998, Thommes et al. 2003). Numerical simulations show that much of the solid mass at each location in the disk is swept up by a single planetary embryo (Kokubo & Ida 1998), which grows by accumulating planetesimals in an annular region around its orbit called a feeding zone.

In the absence of significant radial transport of solid material, embryos cannot grow beyond their isolation mass $M_{\rm iso}$, given by

$$M_{\rm iso} = \left(\frac{8\pi^3 b^3 \Sigma_{\rm solid}^3 a^6}{3 M_*} \right)^{1/2}, \tag{3}$$

where a is the distance from the star, and simulations suggest that $b \sim 10$ (Kokubo & Ida 1998). At this point, an embryo has accumulated all the material in its feeding zone. Radial drifting of planetesimals can alter $M_{\rm iso}$, especially in massive disks or when planetesimals are small (Chambers 2008, Thommes et al. 2003).

Simulations find that oligarchic growth ends when $M \sim M_{\rm iso}/2$ (Kenyon & Bromley 2006). At this point, dynamical friction weakens, the orbits of the embryos become inclined and eccentric, and growth rates slow substantially (Chambers 2001). In the inner disk, $M_{\rm iso} \sim 0.1 M_\oplus$, and postoligarchic growth is required to form Earth-mass planets. Oligarchic growth generates much larger bodies in the outer disk, because of the increasing amount of solid material beyond the snow line and because feeding zone widths increase with distance from the star.

2.4. Formation of Gas-Giant Planets

Planetary embryos larger than Mars acquire substantial atmospheres captured from the surrounding gas disk (Inaba & Ikoma 2003). The mass and density of an atmosphere increase rapidly with increasing embryo mass. The atmospheric structure also depends on the opacity κ and the luminosity L (Ikoma et al. 2000), where L depends mainly on the energy released by impacting planetesimals. The presence of a dense atmosphere increases an embryo's collision cross section because passing planetesimals are slowed by gas drag, increasing the chance of capture. The capture radius can increase by a factor of $\sim$100 for embryos a few times more massive than Earth, greatly increasing the growth rate (Inaba & Ikoma 2003).

Above a critical embryo mass, $M_{\rm cnit}$, the pressure gradient within the atmosphere is insufficient to offset the object's gravity, and gas begins to flow onto the embryo. $M_{\rm cnit}$ depends on L and κ but is typically $\sim 10M_{\oplus}$ (Ikoma et al. 2000, Inaba et al. 2003). The growth of this gaseous envelope is slow at first but becomes increasingly rapid when the total mass reaches $\sim 20M_{\oplus}$ (Pollack et al. 1996). Gas continues to accrete onto the object until the supply is shut off, either because the planet clears a gap in the disk or because the disk dissipates. The end result is a gas-giant planet, with a massive gaseous envelope surrounding a solid core.

The time required for a critical-mass core to grow into a Jupiter-mass planet depends on κ (Hubickyj et al. 2005), which is poorly constrained at present. Early calculations adopted interstellar dust opacities and found growth times of up to 10 Myr, longer than the lifetime of most protoplanetary disks (Pollack et al. 1996). However, the growth of grains in the lower envelope can substantially reduce κ (Movshovitz & Podolak 2008). Models in which κ is 50 times lower than that for interstellar dust typically find shorter growth times of 1–3 Myr (Hubickyj et al. 2005).

Gas-giant planets may also form by the gravitational collapse of portions of a protoplanetary disk. The stability of a disk is characterized by the Toomre criterion

$$Q = \frac{M_* c_s}{\Sigma_{\rm gas} \pi a^2 v_{\rm kep}}, \tag{4}$$

where $\Sigma_{\rm gas}$ and c_s are the disk surface density and sound speed at a distance a from the star, and $v_{\rm kep}$ is the orbital velocity. In the cool, outer regions of relatively massive disks, $Q \lesssim 2$, and hydrodynamical simulations show that dense clumps with masses comparable to that of Jupiter can form on timescales of a few orbital periods (Boss 2000). For gravitationally bound clumps to form and survive to become planets, they must cool on a timescale comparable to the orbital period. Whether this happens is currently the subject of much debate (Boss 2008, Cai et al. 2006, Mayer et al. 2007). If a bound clump does form, refractory grains entrained in the gas will subsequently sediment to the center to form a solid core (Helled et al. 2008).

2.5. Postoligarchic Growth

Once planetary embryos contain most of the solid mass in a region of the disk, dynamical friction from planetesimals becomes weak. If the gas disk has dispersed, tidal damping is also absent (see Section 3.1), and the eccentricities and inclinations of embryos increase owing to mutual interactions. Gravitational focusing is reduced, and the collision rate slows by several orders of magnitude. In the outer disk, little further growth is likely on billion-year timescales (Levison & Stewart 2001), unless the remaining planetesimals are $\lesssim$1 m in size (Goldreich et al. 2004).

In the inner disk, growth continues, but at a slower rate than during the oligarchic phase, as embryos collide with one another and sweep up residual planetesimals. Numerical simulations find that Earth-mass planets form in 10^7–10^8 years (Chambers 2001, O'Brien et al. 2006, Raymond et al. 2006b), which is broadly consistent with cosmochemical estimates for the time required for Earth's formation (Halliday 2004, Touboul et al. 2007). The final stage of accretion is highly chaotic, so many outcomes are possible for similar initial conditions (Chambers 2001).

The number, orbits, masses, and compositions of terrestrial planets are strongly influenced by the presence of giant planets in the system (Chambers & Wetherill 2001, Chambers & Cassen 2002, Levison & Agnor 2003, O'Brien et al. 2006). Planetary embryos tend to be removed from regions of the disk containing strong secular and MMRs associated with giant planets (Chambers & Wetherill

2001), leading to the formation of an asteroid belt (Petit et al. 2001). Secular perturbations from giant planets also excite planetary eccentricities in stable regions, determining the number and spacing of surviving terrestrial planets (Laskar 1997, Levison & Agnor 2003).

Type-I migration: regime in which a low-mass planet induces a linear perturbation in the surrounding disk. Typically causes fast inward migration

Type-II migration: migration regime in which a planet is massive enough to open a gap in the disk so that the disk and planet evolve together

Lindblad resonance: location where the orbital periods of the disk and a planet have a ratio of $(m \pm 1)/m$, where m is an integer

3. PLANETARY MIGRATION

The presence of a planet in a protoplanetary disk modifies the distribution of gas in the planet's vicinity. Gravitational interactions between the planet and the nonuniform arrangement of gas generate torques that alter the planet's orbit. This causes a planet to migrate toward or away from the star and also damps or excites the orbital eccentricity and inclination.

The direction and rate of migration vary depending on the mass of the planet and the local properties of the gas disk. Two limiting cases have been studied extensively. Type-I migration affects Earth-mass planets, where the planet-disk interaction can be treated using linear approximations. The resulting migration rate is proportional to the mass of the planet and the surface density of the disk (Ward 1997). Type-II migration affects Jupiter-mass planets. Here, the planet is massive enough to clear an annular gap in the disk around its orbit, and the planet's motion becomes tied to the viscous evolution of the disk (Ward 1997).

3.1. Classical Type-I Migration

When a planet has a low mass, its gravitational perturbations on the surrounding gas disk are small. However, the interaction can be significant in the vicinity of Lindblad resonances (**Figure 2**), where the angular velocity of the gas Ω and that of the planet Ω_p are related by

$$\Omega = \Omega_p \pm \kappa_e / m. \tag{5}$$

Here, m is an integer, and $\kappa_e \simeq \Omega$ is the epicyclic frequency of the gas (Papaloizou et al. 2007). These resonances are equivalent to first-order MMRs, in which the period P of the gas is $P \simeq P_p(m \pm 1)/m$, where P_p is the period of the planet.

Spiral density waves are launched at Lindblad resonances, carrying angular momentum away from the planet to other regions of the disk (**Figure 3**). The disk outside the planet's orbit exerts a

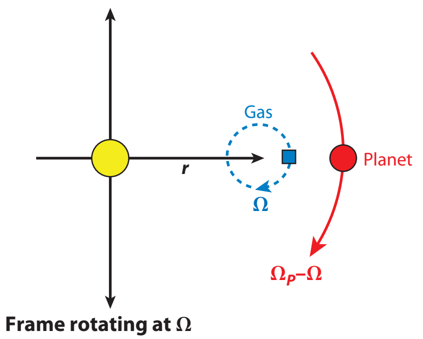

Figure 2

Schematic diagram showing the motion of a parcel of gas and a planet in a reference frame rotating at the average orbital velocity of the gas, Ω. The gas (*blue*) undergoes small epicycles in the rotating frame owing to the eccentricity of its orbit. The planet (*red*) moves at an angular velocity of $\Omega_P - \Omega$ in the rotating frame. A corotation resonance occurs when $\Omega = \Omega_P$. A Lindblad resonance occurs when the gas undergoes an integer number of epicycles for every passage of the planet.

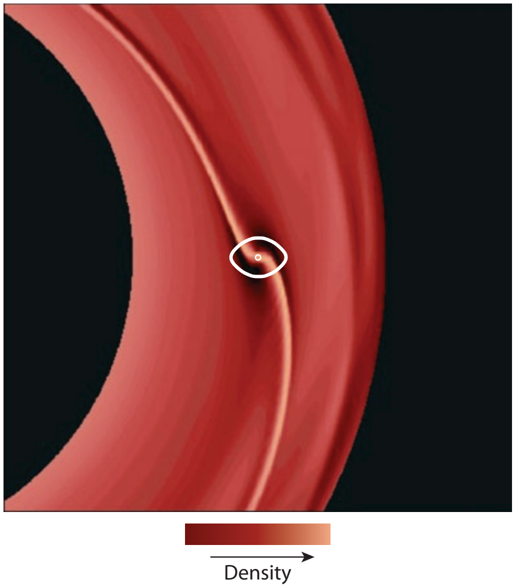

Figure 3

Hydrodynamic simulation showing the response of a protoplanetary disk to an embedded Earth-mass planet. Different shades of red indicate different gas surface densities, and brighter shades indicate higher densities. The inner and outer black regions were not included in the calculation. The planet, located at the center of the plot, launches spiral waves in the gas that apply torques to the planet's orbit, causing migration. The planet is too small to alter the azimuthally averaged surface density, and no gap forms. Figure kindly supplied by Pawel Artymowicz.

negative torque on the planet, whereas the disk inside the orbit exerts a positive torque. In general, these torques are not equal. The outer Lindblad resonances lie closer to the planet than do the inner ones, especially when pressure gradients in the disk are taken into account (Papaloizou et al. 2007). As a result, the outer torque is stronger, and the planet migrates toward the star.

Interactions can also be important at the corotation resonance $\Omega = \Omega_P$, where the gas and planet orbit at the same rate. For low-mass planets, these torques are typically too small to offset the differential Lindblad torque except in regions where there is a strong surface density gradient (Papaloizou et al. 2007).

Corotation resonance: location where the disk orbits the star at the same rate as does a planet

In a vertically isothermal disk, a planet of mass M on a circular orbit a distance a from a star of mass $M_\odot$ migrates inward at a rate

$$\left.\frac{da}{dt}\right)_{\mathrm{I}} \simeq -(2.7-1.1x)\left(\frac{M}{M_*}\right)\left(\frac{\Sigma_{\mathrm{gas}}a^2}{M_*}\right)\left(\frac{1}{h^2}\right)v_{\mathrm{kep}}, \qquad (6)$$

where the unperturbed gas surface density is $\Sigma_{\mathrm{gas}} \propto a^x$, and $h = H_{\mathrm{gas}}/a$, where H_{gas} is the disk scale height (Tanaka et al. 2002). This result, derived analytically, has been confirmed by numerical simulations for planets of up to a few Earth masses (D'Angelo & Lubow 2008, D'Angelo et al. 2003). Type-I migration is extremely rapid: An Earth-mass planet at 1 AU and a 10 $M_\oplus$ core at 5 AU both have lifetimes $\sim 10^5$ years in the MMSN (see **Figure 4**).

Torques from the disk also damp a planet's orbital eccentricity e and inclination i on a timescale that is shorter than the type-I migration time by a factor of $\sim h^2$ (Tanaka & Ward 2004). This damping acts in addition to dynamical friction from planetesimals and will ensure that embryos have low e and i during the oligarchic growth phase. Modifications to the migration and damping

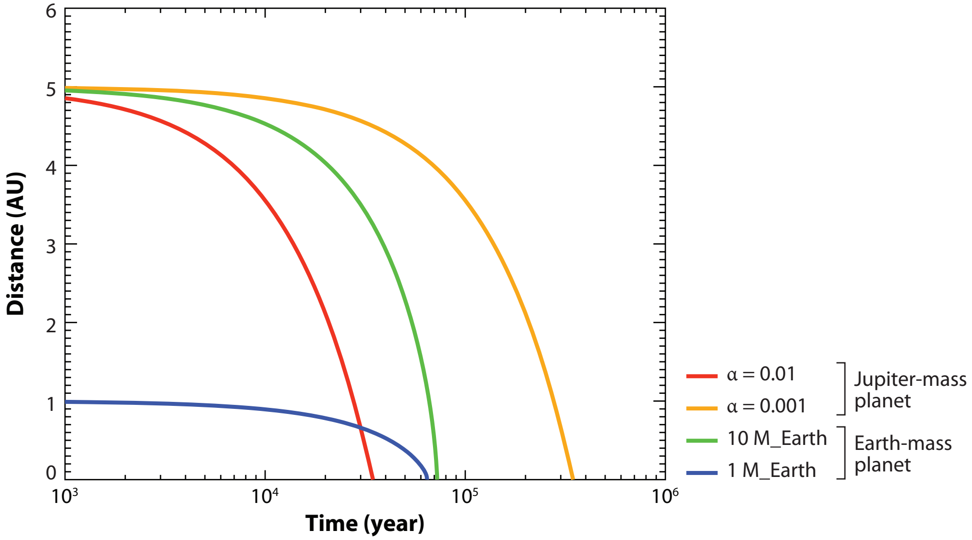

Figure 4

Blue and green curves denote orbital evolution of an Earth-mass planet at 1 AU (astronomical unit) and a 10-Earth-mass core at 5 AU, undergoing type-I migration in the minimum-mass solar nebula according to Equation 6. Red and orange curves denote orbital evolution of a Jupiter-mass planet, beginning at 5 AU, undergoing type-II migration in disks with $\alpha = 10^{-2}$ and 10^{-3}, respectively, according to Equation 8.

formulae have been derived for planets moving on eccentric or inclined orbits (Cresswell & Nelson 2008, Papaloizou & Larwood 2000). Migration is slowed for planets moving on eccentric orbits and reverses direction when $e \gtrsim h$ (Cresswell & Nelson 2008, Papaloizou & Larwood 2000).

3.2. Modified Type-I Migration

The very rapid inward migration predicted by Equation 6 poses severe problems for the formation and survival of planets in general and of giant-planet cores in particular. Much effort has been devoted to searching for factors that modify classical type-I migration. Here, I briefly discuss a few of these.

Most studies of type-I migration have considered a vertically isothermal disk. Recently, it has become apparent that migration changes substantially when radiative transfer within the disk is taken into account (Kley & Crida 2008, Paardekooper & Mellema 2006). If the opacity of the disk is similar to that of interstellar dust, planets with masses $<50M_\oplus$ are likely to migrate outward owing to increased corotation torques (Kley & Crida 2008). If the opacity is lower, owing to dust coagulation, migration will be inward in the outer disk and outward in the inner disk (Paardekooper & Mellema 2006), possibly leading to a favored location for planets to form and survive.

Corotation torques and differential Lindblad torques are both altered in regions where the surface density of the disk changes rapidly with distance from the star. A steep surface density gradient will be present at the inner edge of the disk and also at the inner edge of a dead zone. Outward migration is likely at each of these locations, even if the overall migration trend is inward (Masset et al. 2006b, Morbidelli et al. 2008). Low-mass planets may be preferentially trapped just outside these regions. Inward migration is also slowed at locations where the opacity changes rapidly owing to condensation fronts in the disk (Menou & Goodman 2004). The presence of a toroidal magnetic field, perhaps like that found in an undead zone (Turner & Sano 2008), will also slow or reverse migration (Fromang et al. 2005).

Once a planet grows to $\sim 10 M_{\oplus}$, corotation torques become increasingly important, and type-I migration is no longer linear. In this regime, migration is slowed or reversed depending on the disk viscosity and surface density gradient (D'Angelo et al. 2003, Masset et al. 2006a).

3.3. Type-II Migration

The torques exerted by a massive planet can be sufficiently strong to clear an annular gap in the region of the disk around the planet's orbit (Ward 1997) (see **Figure 5**). To maintain a gap, the torque exerted by the planet on the disk must overcome pressure and viscous forces within the gas. Calculations show that a planet of mass M clears >90% of the gas around its orbit when

$$\frac{3H}{4R_H} + \frac{50\nu}{a v_{\mathrm{kep}}}\left(\frac{M_*}{M}\right) \lesssim 1, \tag{7}$$

where a is the orbital radius, ν is the viscosity, and $R_H = a(M/3M_*)^{1/3}$ is the Hill radius (Crida et al. 2006).

Once a gap forms, a planet's orbital evolution becomes tied to the disk, and the planet migrates at the same rate that gas moves viscously through the disk. If the planet approaches either edge of its gap, the resulting torque imbalance acts to return the planet toward the middle of the gap.

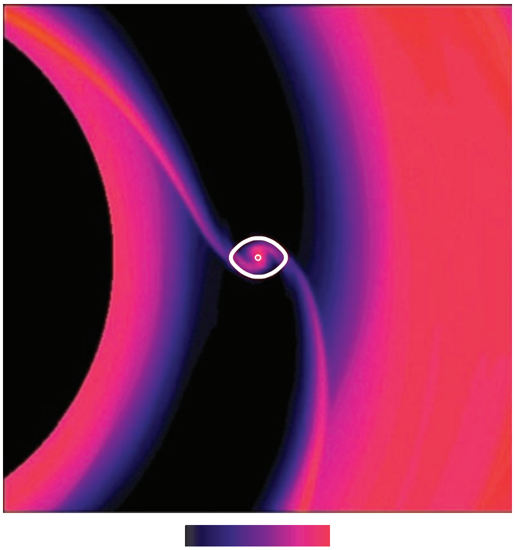

Figure 5

Hydrodynamic simulation showing the response of a protoplanetary disk to an embedded Jupiter-mass planet. Different colors indicate different gas surface densities: pink is high density, whereas blue and black are low density. The innermost black region was not included in the calculation. The planet, located at the center of the plot, is massive enough to clear an annular gap in the disk, and its motion becomes tied to that of the gas. Some gas continues to approach the planet through narrow spiral tendrils extending from the inner and outer portions of the disk. Figure kindly supplied by Pawel Artymowicz.

Away from the inner and outer edges of the disk, a planet migrates inward at a rate

$$\left.\frac{da}{dt}\right)_{\mathrm{II}} = -\frac{3\nu}{2a}. \tag{8}$$

Type-II migration rates are typically slower than type-I rates for large planets, although they are still short enough to pose problems for planetary survival in disks with large viscosities (see **Figure 4**).

For subjovian-mass planets, the surface density in the gap can be a significant fraction of that in the surrounding disk. This is referred to as a partial gap. These planets undergo type-II migration at a reduced rate, both because of corotation torques and because the planet does not feel the full strength of the torque from the outer disk (Crida & Morbidelli 2007).

Very massive planets also migrate more slowly than predicted by Equation 8 because there is a maximum torque that can be exerted by the disk as a result of its viscous evolution. When the planet's mass $M \gtrsim \pi a^2 \Sigma_{\mathrm{gas}}$, the migration rate depends on $\Sigma_{\mathrm{gas}} a^2/M$ rather than on just the viscosity (Edgar 2007, Ida & Lin 2004, Syer & Clarke 1995).

3.4. Type-III Migration

Planets with masses comparable to that of Saturn can undergo a third kind of migration in massive disks. In this situation, the planet clears a partial gap in the disk. The remaining gas close to the planet exerts a corotation torque that grows in proportion to the migration speed. This can generate a positive feedback, leading to very rapid inward or outward migration (Masset & Papaloizou 2003). Currently, it is unclear whether the conditions needed to generate type-III migration are likely to occur in practice (D'Angelo & Lubow 2008).

3.5. Stochastic Migration

If a protoplanetary disk is turbulent, perhaps as a result of MRI, planets will experience torques caused by turbulent density fluctuations. These torques can lead to migration and excite the orbital eccentricity e and inclination i (Nelson 2005). Turbulent eddies vary on timescales comparable to the orbital period (Laughlin et al. 2004, Nelson 2005), so the resulting torques are stochastic rather than smooth, causing a planet's orbit to undergo a random walk. Stochastic migration may act in addition to or replace smooth migration, depending on the level of turbulence (Laughlin et al. 2004, Nelson 2005).

Unlike classical type-I migration, stochastic migration rates are independent of planetary mass M because the disk surface density variations are not generated by the planet. In addition, the random-walk nature of stochastic migration means it is progressively more important than smooth migration at early times when the disk surface density is high (Chambers 2008).

Turbulent torques tend to excite e and i, whereas smooth torques damp these quantities; the latter effect becomes stronger with increasing M (Nelson 2005). Simulations suggest that embryos with $M < M_\oplus$ may reach $e \sim 0.1$ in a disk three times more massive than the MMSN, whereas $10M_\oplus$ cores have $e \sim 0.02$ (Nelson 2005). Planetesimals will also be affected, and damping due to gas drag is unlikely to prevent e excitation for objects $\gtrsim$1 km in size (Britsch et al. 2008). Turbulence levels are much lower in magnetic dead zones, and the corresponding stochastic torques decrease by two orders of magnitude (Oishi et al. 2007).

4. HOW MIGRATION AFFECTS PLANET FORMATION

In this section, we examine how planetary migration is likely to modify the standard model of planet formation for terrestrial and giant planets.

4.1. Giant-Planet Formation

If giant planets form via core accretion, their cores will be susceptible to type-I migration until they are massive enough to open a gap in the disk. In the absence of migration, cores large enough to initiate gas accretion ($M \geq M_{\mathrm{cnit}}$) should form in $\sim 10^6$ years in disks 3–10 times more massive than the MMSN (Chambers 2006a, Lissauer 1987, Thommes et al. 2003). However, the lifetime of a core at 5 AU with respect to classical type-I migration is $<10^5$ years, so migration poses a severe problem for the core-accretion model if cores really migrate this rapidly.

As we saw in Section 3.2, there is considerable uncertainty about the rate and even the direction of type-I migration. However, the rapid pace of recent developments has left models of planet formation struggling to keep up. As a result, most studies of planet formation that include type-I migration tend to adopt a migration rate similar to that of Equation 6, modified by a migration efficiency factor $f_I \leq 1$.

Some studies have concluded that giant-planet cores will inevitably be lost before they reach M_{cnit} if $f_I \gtrsim 0.1$ (Alibert et al. 2005, Ida & Lin 2008a). This conclusion may be overly pessimistic. The presence of small fragments generated by collisions between planetesimals, and the existence of atmospheres around embryos larger than Mars, can combine to allow cores to form very rapidly (Inaba et al. 2003, Rafikov 2004). When these effects are taken into account, critical-mass cores may form and survive with $f_I = 1$ in the later stages of a disk's evolution, when Σ_{gas} has declined from its initial value and migration rates are lower (Chambers 2006b, Thommes et al. 2007). If cores form too late, however, there will not be enough gas left to grow into a gas-giant planet. This suggests there is a window of opportunity for giant-planet formation, which will naturally be longer in disks with a high dust-to-gas ratio (see **Figure 6**). In this scenario, cores form at a late stage either because planetesimal formation is delayed (Cuzzi et al. 2008) or because these cores represent the last generation of objects to form, with earlier generations having been lost.

Type-I migration can increase the growth rate of a single core by moving it into a new region of the disk when planetesimals in its original feeding zone become depleted (Alibert et al. 2005). Similarly, stochastic migration allows a lone core to grow more rapidly (Rice & Armitage 2003). The presence of multiple cores in the same region tends to offset this enhancement, however, as all the objects compete for the same resources (Chambers 2008). Stochastic migration has the additional advantage that a core will occasionally enter a region where very few planetesimals remain. In this case, M_{crit} declines substantially as the core's atmosphere adjusts to the lower collision rate, and runaway gas accretion can begin at an earlier stage (Rice & Armitage 2003).

Stochastic migration may have a negative effect on giant-planet formation for two reasons. The torques induced by turbulent density fluctuations tend to increase the orbital eccentricities e of embryos and planetesimals. Simulations show that the maximum e tends to be larger for small bodies, reaching ~ 0.1 for objects less massive than the Earth (Nelson 2005). At these eccentricities, planetesimals and embryos will encounter one another at high speeds, greatly slowing or halting runaway and oligarchic growth (Nelson 2005). Stochastic migration also tends to reduce the lifetime of cores that do form. Although a small fraction of objects can survive by migrating into the outer disk, most objects are lost more rapidly than they would be under smooth type-I migration alone (Johnson et al. 2006). For these reasons, giant-planet formation may be restricted to laminar regions of the disk.

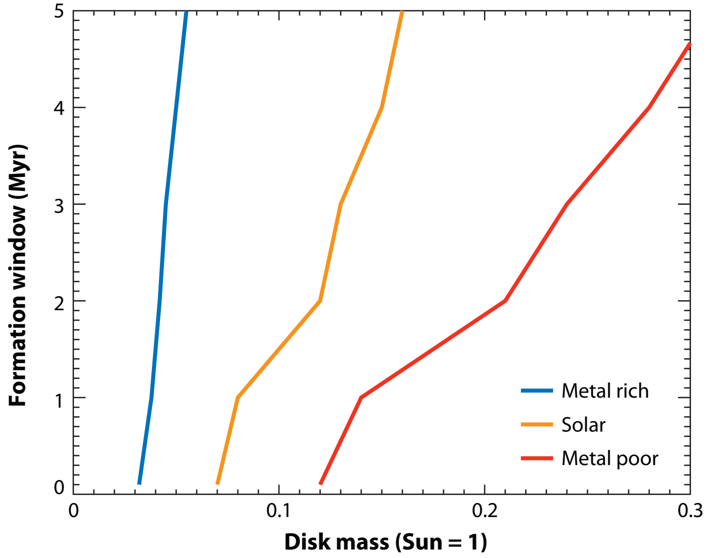

Figure 6

Window of opportunity for giant-planet formation as a function of disk mass for three protoplanetary disk models. The window of opportunity is limited by the loss of giant-planet cores via type-I migration as well as the eventual dispersal of the gas disk. The metal-poor and metal-rich disks have dust-to-gas ratios of 0.6 and 1.8 times that of the Sun's protoplanetary nebula. The disk viscosity parameter $\alpha = 0.001$ in each case. For comparison, the minimum-mass solar nebula has a mass of 0.01–0.02 $M_{\odot}$. Adapted from figure 3 of Thommes et al. (2003).

Most studies of planet formation have considered disks with surface density profiles that vary smoothly with distance from the star. However, real disks probably contain density discontinuities at the ice line (Kretke & Lin 2007) and at the outer edge of a dead zone (Matsumura et al. 2007). At these locations, the torques causing type-I migration will be reversed, generating traps for migrating embryos and preferential locations for giant-planet formation (Ida & Lin 2008b, Matsumura et al. 2007). Dead zones may provide a particularly beneficial environment for giant planets to survive because cores begin to open a gap when they are still small (Edgar et al. 2007, Rafikov 2002), and the resulting type-II migration rates are low (Matsumura et al. 2007).

Once one giant planet has formed, it may aid the growth of additional planets. Planetary embryos migrating inward from further out in the disk will be trapped at external MMRs with the planet (Thommes 2005) or will accumulate at the outer edge of the gap opened by the planet, where type-I migration reverses direction owing to strong corotation torques (Masset et al. 2006b). If sufficient material accumulates in these regions to form a critical-mass core, a second giant planet can form.

4.2. Giant-Planet Migration

Type-II migration, which affects gas-giant planets, is somewhat better constrained than the other migration regimes. It appears that giant planets will inevitably undergo type-II migration in a viscously evolving disk. When a planet's mass M is small compared with the mass of the disk, the migration rate is independent of M. In an α disk, the migration rate is also roughly independent of distance a from the star. Thus, an important factor determining how far a planet migrates is the time interval between its formation and the dissipation of the gas disk (Armitage et al. 2002, Trilling et al. 2002). Planets with $M \gtrsim \Sigma_{gas}\pi a^2$ migrate more slowly at a rate that depends on Σ_{gas}/M (Edgar 2007, Ida & Lin 2004, Syer & Clarke 1995). This means that planets that form late probably undergo little migration, both because the disk mass is lower and because there is less time in which to migrate before the disk disperses.

The rate at which a giant planet accretes gas slows substantially once it becomes massive enough to clear a gap in the disk (D'Angelo & Lubow 2008, Veras & Armitage 2004). In the absence of migration, the observed increase in planetary mass with distance (**Figure 1**) may simply reflect the likely increase in the gap-opening mass with distance from the star (Crida et al. 2006, Edgar et al. 2007). However, migration will modify this picture. Planets that form early will tend to have lower masses because both formation time and the gap-opening mass are likely to increase with a (Edgar et al. 2007). Many of these planets will migrate substantially to become hot Jupiters or fall into the star. Planets with larger a form late and have large M, so they undergo little migration. However, there may be a relatively short window of opportunity for giant-planet cores to form and become planets like Jupiter (Armitage et al. 2002). If the core forms too early, the planet will migrate inward and be lost. If the core forms too late, there will be too little gas to form a giant planet, or the disk will have dissipated. As a result, many stars may not have surviving giant planets.

This combination of factors may explain several aspects of the observed distribution of extrasolar planets, including the increasing mass and frequency of planets with increasing a (Cumming et al. 2008) and the fraction of stars that have planets. It may also explain the low frequency of planets with $0.1 < a < 1$ AU, sometimes referred to as a planetary desert (Armitage 2007b, Ida & Lin 2004). Because type-II migration is generally inward, the desert should extend inward from the innermost distance at which giant planets are able to form and avoid significant migration. In the standard core-accretion model, giant planets form beyond the snow line, traditionally located several AU from a solar-mass star. However, the outer edge of the desert appears to lie at ~1 AU (see **Figure 1**). This suggests either that giant-planet cores can form interior to the ice line in some cases (perhaps in massive disks) or that the ice line in many disks lies at ~1 AU at the time when planetesimals form.

It remains unclear whether most planets stop migrating and become hot Jupiters when their orbits approach the star. Some simulations of planet formation that incorporate type-II migration predict that ~90% of hot Jupiters must collide with their star to explain the observed distribution of extrasolar planets (Ida & Lin 2004). The observational evidence for a pileup of planets with $a < 0.1$ AU is relatively weak (Armitage et al. 2002, Rice et al. 2008), and the high frequency of hot Jupiters seen in **Figure 1** can be largely attributed to a selection bias caused by surveys for transiting planets.

Several mechanisms have been proposed to halt migration close to the star. Protoplanetary disks are probably truncated several stellar radii from the surface of the star by interactions with the star's magnetic field (Bouvier et al. 2007). This edge is located at or somewhat interior to the corotation point, where the orbital period of the gas is equal to the star's rotation period. Once a planet moves far inside the inner edge of a disk, the planet's outer Lindblad resonances no longer lie within the disk, and migration will typically cease. The presence of a giant planet may also clear away the disk interior to its orbit, forming a cavity (Varnière et al. 2006), with the same effect. However, recent simulations suggest that a planet must already lie close to the inner edge of the disk for a cavity to form (Crida & Morbidelli 2007).

Hot Jupiters orbiting outside the corotation point experience significant outward torques owing to stellar tides. These planets begin to lose mass to the star once they pass within the Roche limit at $a \simeq 2.2R(M_*/M)^{1/3}$, where R is the planetary radius (Faber et al. 2005). Both these processes slow inward migration and may increase the chances that a planet survives (Trilling et al. 1998).

Migration to the inner edge of the disk does not guarantee survival. If a planet lies just interior to the disk, corotation resonances that typically damp eccentricity no longer operate. Some outer Lindblad resonances can remain effective, however, and these will tend to increase a planet's eccentricity, especially if the planet is more massive than Jupiter. This opens the possibilities of continued disk interaction and migration and of eventual collision with the star (Rice et al. 2008).

A substantial fraction of the extrasolar planets in multiplanet systems are in a MMR with one another (Marcy et al. 2005). The 2:1 resonance is especially favored with at least four known examples (Laughlin & Chambers 2001, Lee et al. 2006, Sándor & Kley 2006, Sándor et al. 2006). Resonances have a low probability of occurring by chance, which suggests that these pairs of planets were captured into resonance. This is especially likely for systems like GJ 876, for which the amplitude of the resonant librations is very small and the orbital apsides are also in resonance. Capture into a MMR requires the presence of a dissipative force; the leading candidate is migration in a protoplanetary disk (Lee & Peale 2002).

When two giant planets are present in a disk, each clears a gap around its orbit. If the inner planet migrates more slowly than the outer planet, perhaps because it is not massive enough to completely clear a gap (Crida & Morbidelli 2007), the orbits will converge. Once the orbital radii are within approximately a factor of two, the gas between the planets is confined to a narrow ring that is eroded away on a timescale of $\sim$100 orbits (Bryden et al. 2000). The two planets now orbit in a common gap. The outer disk continues to push the outer planet inward, whereas the inner disk slows the inward migration of the inner planet or pushes it outward. This convergent migration can lead to capture in a MMR provided that the planets do not approach each other too quickly (Lee & Peale 2002, Quillen 2006).

Simulations show that systems like GJ 876 will form naturally as a result of convergent migration (Lee & Peale 2002), ending up deeply embedded in a MMR and an apsidal resonance. Once captured into a resonance, orbital eccentricities typically rise rapidly unless they are damped by a substantial inner disk (Crida et al. 2008, Sándor et al. 2006). Interaction with a third, smaller planet interior to the giants can subsequently break the apsidal resonance while maintaining the MMR (Sándor & Kley 2006, Sándor et al. 2006).

The Solar System appears to represent an unusual case in that its giant planets are not in resonance and show no sign of having migrated significantly within the Sun's protoplanetary disk. Unlike many of the multiplanet extrasolar systems, including those in the 2:1 resonance, the innermost giant planet in the Solar System is substantially more massive than the giant planets orbiting further from the Sun. If Saturn formed closer to Jupiter than it is today, and the two planets formed a common gap, migration of both planets would have been reduced (Morbidelli & Crida 2007). Saturn's low mass means it probably opened a partial gap in the disk while Jupiter cleared a deeper gap. As a result, the inward disk torque exerted on Saturn would have been comparable to the outward torque exerted on Jupiter, minimizing orbital migration (Morbidelli & Crida 2007).

Jupiter, lying closest to the Sun, was probably the first giant planet to form. Simulations show that a second core, migrating inward from further out in the disk, would have become trapped temporarily near the edge of the gap cleared in the disk by Jupiter, owing to strong corotation torques (Pierens & Nelson 2008). As this core, destined to become Saturn, gained mass by accreting gas, it would have moved inward, becoming trapped in the 2:1 resonance and ultimately in the 3:2 resonance with Jupiter (Masset & Snellgrove 2001, Pierens & Nelson 2008). Uranus and Neptune, forming later still, would have become trapped in additional MMRs with their interior neighbor (Morbidelli et al. 2007). It seems likely that subsequent interactions with a disk of residual planetesimals would have disrupted these MMRs, leading to the current configuration of the outer Solar System (Tsiganis et al. 2005).

4.3. Terrestrial-Planet Formation

Terrestrial planets are too small to undergo type-II or type-III migration. However, they are affected by smooth type-I migration and also by stochastic migration caused by turbulent density fluctuations. The problem posed by rapid inward type-I migration is somewhat less severe for terrestrial planets than for giant-planet cores because the former are less massive, and the final stage of terrestrial-planet formation can take place after the disk has dissipated. Nonetheless, smooth migration and stochastic migration are likely to have had a substantial impact on the formation of terrestrial planets.

Simulations of terrestrial-planet formation with type-I migration find that growth and migration initially take place in the inner disk, becoming more important at progressively larger distances over time (McNeil et al. 2005). Much of the solid mass present inside 1–1.5 AU is lost and replaced by planetary embryos migrating in from further out in the disk (Daisaka et al. 2006, McNeil et al. 2005). In situations in which a large embryo forms further out than a small one, convergent migration takes place, frequently leading to capture into a MMR (Cresswell & Nelson 2008, McNeil et al. 2005). Two or more embryos in MMRs with one another migrate inward together at a rate determined by their average mass (McNeil et al. 2005). Elsewhere, interactions between neighboring embryos increase their orbital eccentricities. However, tidal damping by the disk ensures that $e \lesssim h$, where h is the disk aspect ratio, so type-I migration remains inward (Cresswell & Nelson 2008). Other than the formation of MMRs, interactions between embryos have relatively little effect on type-I migration rates (Daisaka et al. 2006).

Type-I migration rates decline over time as $\Sigma_{\rm gas}$ falls, so embryos that form late are more likely to survive. Eventually, the migration timescale becomes longer than the disk lifetime, and migration ceases to be important. The final mass of material in the terrestrial-planet region is only weakly dependent on the initial disk mass (Ida & Lin 2008a). Massive disks generate large embryos at an early stage. These are subsequently lost via migration, and the amount of mass available to form planets is reduced (Daisaka et al. 2006). In low-mass disks, large embryos take longer to form, and the resulting migration rates are lower, so the final planetary masses more closely reflect the initial disk mass. Simulations show that Earth-mass planets form readily in disks three to four times the mass of the MMSN when the type-I migration efficiency $0.25 \leq f_I \leq 1$ (McNeil et al. 2005). A natural consequence of migration is that $\gtrsim$10% of the mass contained in these planets originates from beyond 2 AU (McNeil et al. 2005).

MRI is less likely to operate in the terrestrial-planet region than in the outer disk because the disk is more opaque and the volume density of dust grains is higher (Matsumura & Pudritz 2006). In a layered disk, in which MRI is confined to the surface layers, turbulent torque fluctuations at the midplane will be reduced by two orders of magnitude (Oishi et al. 2007), and stochastic migration will be unimportant. If MRI does operate at the disk midplane, the main effect will be to slow down the runaway and oligarchic stages of growth (Nelson 2005). In general, stochastic migration does not prevent embryos from colliding with the star if smooth type-I migration is also operating (Chambers 2008, Johnson et al. 2006, Ogihara et al. 2007). Even in the absence of smooth migration, the excitation of e by stochastic torques combined with smooth tidal damping leads to a net loss of angular momentum and slow inward migration (Ogihara et al. 2007). Toward the end of the disk's lifetime, when $\Sigma_{\rm gas}$ is small, MRI may operate throughout the disk. When $\Sigma_{\rm gas} \lesssim 1\%$ of that in the MMSN, the combination of stochastic migration and tidal damping will tend to form terrestrial planets with widely spaced, nearly circular orbits, like those in the Solar System (Ogihara et al. 2007).

The existence of hot Jupiters raises the question whether a giant planet migrating through the inner disk prevents the formation of terrestrial planets. Recent calculations suggest this is not the

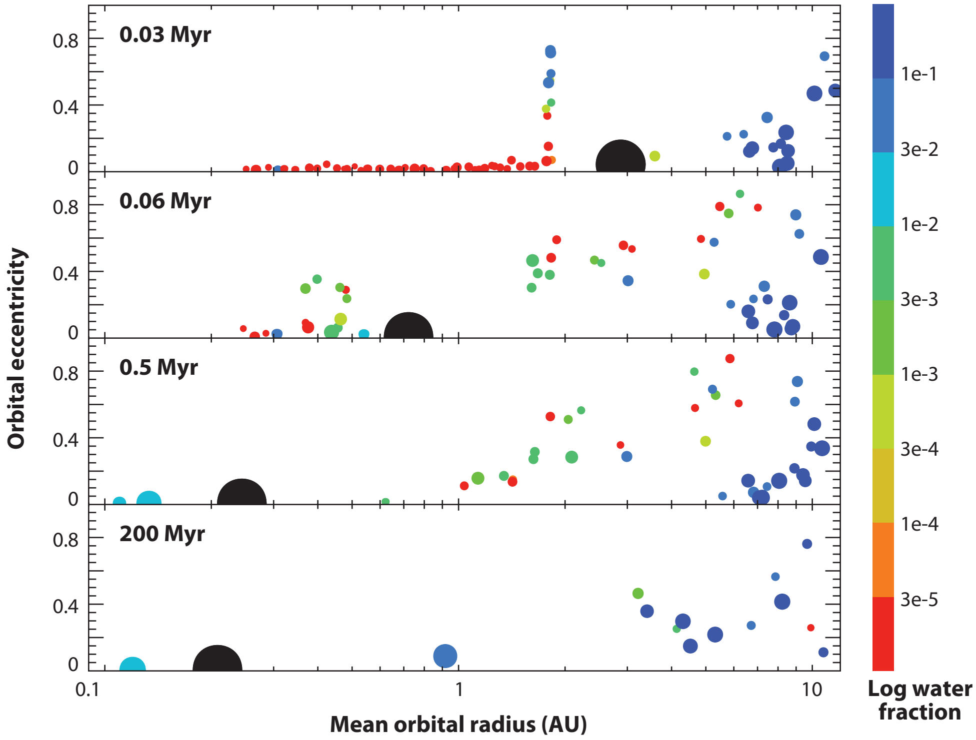

Figure 7

A simulation of terrestrial-planet formation in the presence of a migrating gas-giant planet. Each circular symbol represents a planetary embryo, with symbol radius proportional to mass to the third power, and symbol color indicates water mass fraction. A Jupiter-mass giant planet (*black circle*) begins the simulation at 5 AU (astronomical units) and migrates to 0.25 AU in 10^5 years. Planetesimals were included in the calculation but are not shown here. Adapted from figure 1 of Raymond et al. (2006a). Data were kindly supplied by Sean Raymond.

case (see **Figure 7**). In the absence of type-I migration, solid bodies interior to a migrating giant planet are either captured into MMRs or scattered into eccentric orbits that cross the giant's orbit and travel into the outer disk (Fogg & Nelson 2007a, Raymond et al. 2006a). Bodies in interior MMRs are shepherded inward by the giant planet as it migrates. As the orbits of these shepherded objects shrink, the collision rate between them rises, and they rapidly aggregate into one or more large planets (Fogg & Nelson 2007a, Raymond et al. 2006a). Continued migration while in a resonance causes e to increase, and the survival of these hot Earths depends sensitively on when the giant planet stops migrating (Fogg & Nelson 2007a).

The eccentricities of the scattered embryos are reduced by dynamical friction, and gas drag circularizes the orbits of scattered planetesimals. As a result, a new disk containing a few Earth masses of solid material forms beyond the giant planet. Postoligarchic growth within this disk forms one or more terrestrial planets on a timescale similar to that of disks without a migrating giant planet (Raymond et al. 2006a). These planets are likely to contain a substantial amount of water-rich material from the outer disk, both because a large amount of radial mixing takes place during the giant planet's migration and because, once the giant has passed, there is nothing to prevent small planetesimals from drifting inward from the outer disk owing to gas drag (Raymond et al. 2006a).

The picture outlined above remains broadly similar if the embryos in the inner disk are also undergoing migration, in this case in the type-I regime (Fogg & Nelson 2007b). Tidal damping of the embryos reduces the fraction that are scattered beyond the giant's orbit, whereas shepherding is enhanced, but the outcome is qualitatively similar to the case without type-I migration (Fogg & Nelson 2007b).

SUMMARY POINTS

1. The standard model of planet formation predicts that rocky, terrestrial planets will form in the inner regions of a protoplanetary disk ($a \lesssim 3$ AU), whereas gas-giant planets will form in the outer disk. This model was developed without considering the possibility of planetary migration.
2. The observed distribution of extrasolar planets strongly suggests that giant planets undergo orbital migration after they form. The existence of hot Jupiters close to their star, pairs of planets in resonance, and the increasing mass and frequency of planets with distance can all be understood if some planets undergo inward migration after they form.
3. Planets can migrate owing to gravitational interactions with gas in their protoplanetary disk. Two idealized cases have been studied extensively: Earth-mass planets migrate at a rate proportional to their mass and the local gas surface density (type-I migration). Jupiter-mass planets clear an annular gap in the disk and migrate with the viscous evolution of the disk (type-II migrate).
4. There is currently much uncertainty about the rate and the direction of type-I migration in real disks. Low-mass planets in nonmagnetic, isothermal disks should migrate rapidly inward. Migration may be slowed or reversed by changes in the thermal structure of the disk due to the presence of a magnetic field, if the orbit becomes eccentric or if the planet's mass $\gtrsim 10 M_\oplus$.
5. Type-II migration appears to be robust. However, migration is slowed compared with the idealized case if the planet clears only a partial gap in the disk or if the mass of the planet becomes comparable to that of the disk.
6. Giant-planet cores can form and survive under idealized type-I migration, but only if planetesimals are relatively small and the cores form after the disk has partially dissipated. Stochastic migration caused by turbulent density fluctuations is probably highly detrimental to planet formation.
7. Gas-giant planets that form early will undergo inward type-II migration to become hot Jupiters or to fall into their star. Giants that form lately probably undergo little migration. Pairs of giant planets migrating at different rates can become captured into mean-motion and apsidal resonances with one another. It is possible that all four giant planets in the Solar System were once in such resonances.
8. Terrestrial planets can form in the presence of idealized type-I migration, but the formation efficiency is reduced in more massive disks. The final mass of the terrestrial planets depends only weakly on the initial disk mass. Giant-planet migration does not prevent terrestrial planets from forming, but these planets are likely to have a much higher volatile content than Earth.

FUTURE ISSUES

1. Improved estimates of the direction and rate of migration in realistic protoplanetary disks, especially for low-mass planets, are necessary.
2. We need a better understanding of the mechanism driving viscous accretion in disks and of the level of turbulence in regions where planets form.
3. The gathering of large, unbiased samples of extrasolar planets, and extending the distribution to lower masses and larger orbital distances, should be pursued.
4. High-resolution images of protoplanetary disks, with and without planets, down to astronomical-unit scales, are needed.

DISCLOSURE STATEMENT

The author is not aware of any biases that might be perceived as affecting the objectivity of this review.

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