Planet migration

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A planetary system may undergo significant radial rearrangement during the early part of its lifetime. Planet migration can come about through interaction with the surrounding planetesimal disk and the gas disk—while the latter is still present—as well as through planet-planet interactions. We review the major proposed migration mechanisms in the context of the planet formation process, in our Solar System as well as in others.

1. Introduction

The word planet is derived from the Greek word “planetes”, meaning wandering star. Geocentric views of the Universe held sway until the Middle Ages, when Copernicus and Kepler developed a better phenomenological explanation of planetary wanderings, which with small modifications has withstood the test of time. Kepler’s first law of planetary motion states that planets travel along elliptical paths with one focus at the Sun. Thus, although planets wander about the sky, in this model their orbits remain fixed and they do not migrate. In his physical model of the Solar System, Newton theorized that planets gradually altered one another’s orbits, and he felt compelled to hypothesize occasional divine intervention to keep planetary trajectories well-behaved over long periods of time. In the early 1800s, Poisson pointed out that planetary-type perturbations cannot produce secular changes in orbital elements to second order in the mass ratio of the planets to the Sun, but Poincaré’s work towards the end of the 19th century suggests that the Solar System may be chaotic. The stability of mature planetary systems is a fascinating topic, but we shall be concerned herein with the potentially much more rapid migration of planets during and immediately following the epoch of their formation. Modern research into this topic began when Goldreich and Tremaine (1980) showed that density wave torques could have led to significant orbital evolution of Jupiter within the protoplanetary disk on a timescale of a few thousand years, and research accelerated when giant planets were found much closer to their stars (Mayor and Queloz 1995) than predicted by models of their formation (Lin et al. 1996, Bodenheimer et al. 2000).

In Section 2, we discuss models for the migration of giant planets within our own Solar System which may have occurred as the results of interactions of the planets with one another and with small solid bodies. Section 3 summarizes models of the potentially substantial planetary migration that results from (primarily gravitational) interactions between a planet and a gaseous protoplanetary disk. The predictions of this model are compared to observations of Saturn’s rings and moons in Section 4. In Section 5, the models are applied to extrasolar planetary systems.

2. Migration of Jupiter, Saturn, Uranus, Neptune and Pluto

In studying the accretion of the giant planets, Fernandez and Ip (1984) first noted that proto-Uranus and -Neptune could undergo significant orbital migration due to angular momentum exchange with a (sufficiently massive) planetesimal disk. The mechanism operates as follows: Gravitational stirring by Uranus and Neptune imparts high eccentricities on the surrounding planetesimals. Those which acquire sufficiently small perihelia
can be “handed off” to the next-innermost planet, with a resultant gain in angular momentum for the first planet. In this way, planetesimals get passed inward from Neptune to Uranus to Saturn and finally to Jupiter, which is massive enough to readily eject them from the Solar System. (The other giant planets are also massive enough to eject planetesimals from the Solar System; however, the characteristic timescales for direct ejection are longer than for passing the planetesimals inwards to the control of Jupiter.)

It is generally accepted that planetesimal scattering by the giant planets, principally Jupiter, is what formed the Oort cloud, the quasi-spherical distribution of comets orbiting at $\sim 10^3$–$10^5$ AU from the Sun (Duncan et al. 1987). Since Jupiter does the lion’s share of the work, it undergoes a net loss of angular momentum and its orbit shrinks; Saturn, Uranus and Neptune mainly contribute by passing planetesimals down, so they gain angular momentum. For a planet on a circular orbit, change in angular momentum $L$ is related to change in semimajor axis $a$ according to

$$\Delta L = \frac{1}{2} M \sqrt{\frac{GM_\odot}{a}} \Delta a,$$

where $M$ is the planet’s mass, so Jupiter, being the innermost and most massive, migrates the shortest distance, while Neptune migrates farthest. Hahn and Malhotra (1999) investigated this migration mechanism in detail, using it to try and reproduce the eccentricities of Pluto and the other Kuiper belt objects sharing the Neptune exterior 3:2 resonance (Plutinos). They found that with a planetesimal disk of about 50 Earth masses ($M_\oplus$), Neptune migrates the right distance to pump the eccentricities of objects carried along in its 3:2 resonance to the observed values, $e \sim 0.3$. Larger/smaller disk masses produced too much/not enough migration. Migration takes place over a timescale of tens of millions of years, after which time it stalls due to depletion of the planetesimals.

A preceding phase of migration for Neptune, Uranus and Saturn, shorter but more violent, was proposed in the model of Thommes et al. (1999, 2002a). In trying to account for the formation of the “ice giants”, Uranus and Neptune (14.6 and 17.2 $M_\oplus$ respectively, $\sim 90\%$ of which are condensibles) at their present orbital radii (19 and 30 AU respectively), one runs into serious timescale problems (Lissauer et al. 1995, Levison and Stewart 2001, Thommes et al. 2002b). At such large heliocentric distances, the accretion timescale of such massive bodies far exceeds the lifetime of the nebular gas (about $10^7$ years or less, e.g., Strom et al. 1990); once planetesimal random velocities are no longer damped by aerodynamic gas drag, it may not be possible to produce an ice giant-sized body on any timescale. Indeed, at 20 AU, the mass at which the surface escape velocity from a (Uranus/Neptune-density) body equals the escape velocity from the Sun, is only a bit over an Earth mass. In the model of Thommes et al., the Jupiter-Saturn region serves as the birthplace of all the giant planets, thus alleviating the timescale problem. Assuming that gas giant planets form by core accretion (e.g., Pollack et al. 1996), one of the protoplanets—likely proto-Jupiter—eventually reaches runaway gas accretion and acquires a massive gas envelope, abruptly (over $\sim 10^5$ years) expanding its gravitational reach and destabilizing its neighbours’ orbits. As a result the other giant protoplanets are scattered, predominantly outwards, ending up with aphelia in the still accretionally un-evolved outer planetesimal disk. Dynamical friction with the planetesimals then reduces the eccentricities of these scattered giant protoplanets, decoupling them from Jupiter and from each other on a timescale of a few million years. Numerical simulations show that this sequence of events commonly results in an outer planetary system similar to our own, with the scattered protoplanets eventually settling down to nearly circular orbits of radii comparable to those of Saturn, Uranus and Neptune. An example is shown in Fig. 1. A frequent side effect in the (stochastic) simulations is strong gravitational stirring of
Figure 1. A numerical simulation showing how Uranus and Neptune could have originated among Jupiter and Saturn. Semimajor axis, perihelion and aphelion distance as a function of time are plotted for each body. Initially, four $10 \, M_{\oplus}$ bodies are placed on circular orbits in the Jupiter-Saturn region. The innermost one’s mass is increased over a $10^5$ year timescale to that of Jupiter, to simulate the accretion of a massive gas envelope. As a result, the remaining bodies are scattered outward; their initially high eccentricities are damped by interaction with the planetesimal disk which exists beyond $\sim 12$ AU. After 5 Myrs of evolution, the result is a giant planet system which looks qualitatively similar to our own, with the scattered and recircularized bodies standing in for Saturn, Uranus and Neptune. See Thommes et al. (2002a) for details.

In situ accretion of Uranus and Neptune could still have occurred if the growing ice giants had somehow continued to be supplied with planetesimals of relatively low velocity dispersion. There are several ways in which this could have happened. One possibility is that differential migration between the growing protoplanets and the planetesimals caused the former to sweep, predator-like, through heretofore dynamically cold parts of the planetesimal disk (Ward and Hahn 1995, Tanaka and Ida 1999, Bryden et al. 2000). Accretion would stall when the protoplanet mass becomes sufficiently large that it no longer ploughs through the planetesimals, but instead captures the planetesimals in its
mean-motion resonances and pushes them ahead of it (Kary et al. 1993). Insofar as the differential migration is to come about from interaction with the gas disk (aerodynamic gas drag on the planetesimals, gravitational interaction of the protoplanets with the disk (see below), or some combination thereof), this mechanism will only operate for $\sim 10^7$ years. After that, migration will become much slower, since migration-inducing interactions would only involve the much less massive solids component of the protoplanetary disk. Another possibility is that collisions among the planetesimals acted to damp their velocities sufficiently for significant further accretion to take place after dissipation of the gas. Thommes et al. (2002b) make a simple estimate of post-gas growth timescales in the limit of “maximally effective” collisional damping (meaning simply that the random velocity damping timescale is taken to be equal to the inter-planetesimal collisional timescale), and show that planetesimals of order ten kilometers in size or smaller—collision rate is inversely proportional to planetesimal radius—could conceivably permit the accretion of $\sim 10 M_{\oplus}$ bodies in the 20 - 30 AU region in of order $10^8 - 10^9$ years.

3. Disk - Planet Interactions

A disk with nonzero viscosity will transport angular momentum outward, resulting in the spreading of the disk; less and less of the disk material ends up with more and more of the angular momentum and as a result, there is a net inward transport of mass (Lynden-Bell and Pringle 1974). In modeling disk viscosity $\nu$, the prescription of Shakura and Sunyaev (1973) is commonly used:

$$\nu = \alpha h^2 \Omega,$$

(3.1)

where $\alpha$ is a dimensionless parameter, $h$ is the disk half-thickness, and $\Omega$ is the disk angular velocity. Estimates for $\alpha$ in the protosolar nebula range from $10^{-4}$ to $10^{-2}$ (e.g., Cabot et al. 1987, Dubrulle 1993). The source of viscosity is unknown; it may be turbulent motion resulting from convective instability in the disk (Lin and Papaloizou 1980), damping of density waves launched by embedded planetary bodies (Larson 1989, Goodman and Rafikov 2001), or magnetohydrodynamic turbulence (Balbus and Hawley 1991). The latter mechanism requires that the disk be sufficiently ionized to be strongly coupled to the magnetic field. At a radius of less than about 0.1 AU from the star, collisional ionization ought to be able to accomplish this; further out, ionization by cosmic rays near the disk surface will dominate. Thus it is possible that beyond $\sim 0.1$ AU, a protoplanetary disk only transports angular momentum within relatively thin layers on the outer surfaces of the disk (Gammie 1996).

Individual gas molecules move along paths which are nearly Keplerian ellipses. However, the picture becomes more complex once protoplanets form and perturb the disk. This problem was first investigated in the context of satellite interactions with planetary rings by Goldreich and Tremaine (1980), and extended by Ward (1986, 1997) and Artymowicz (1993) to planetary interactions with a protoplanetary disk. A review of the theory of planet-disk interactions is given by Lin et al. (2000), as well as by Goldreich and Sari (2002), who investigate the resulting evolution of planet eccentricities. We provide an abbreviated summary below.

It is convenient to begin by expressing the gravitational potential of the planet as a Fourier series:

$$\psi = \sum_l \sum_{m=0}^{\infty} \psi_{l,m}(r) \cos(m(\phi - \Omega_{l,m} t)),$$

(3.2)

where $\phi$ is the azimuthal angle, $\psi_{l,m}(r)$ is an amplitude which depends on radius, and $\Omega_{l,m} = \Omega_p + (l-m)\kappa_p/m$ is the pattern speed, with $\Omega_p$ being the angular frequency...
of a planet on a circular orbit, and $\kappa_p$ being the planet’s epicyclic (radial oscillation) frequency. Both $l$ and $m$ are integers; for each value of the azimuthal mode number $m$, there are—to lowest order in eccentricity—three components that contribute to the pattern speed: $\Omega_{l=m} = \Omega_p$, and $\Omega_{m\pm1,m} = \Omega_p \pm \kappa_p/m$. Each component, in turn, has three associated resonances: a corotation resonance, at which the pattern speed matches the disk angular velocity, $\Omega_{l,m} = \Omega_d$, and an inner and outer Lindblad resonance, at which the difference between the disk angular velocity and the pattern speed is a harmonic of the epicyclic frequency, $\Omega_d - \Omega_{l,m} = \pm \kappa_d/m$. For Keplerian orbits, $\kappa_p = \Omega_p$ and $\kappa_d = \Omega_d$.

At each of the resonance sites, angular momentum and energy are exchanged between the disk and the planet, and significant epicyclic motion is excited in the gas. We can view the situation as a resonating disk particle always being at the same phase in its radial oscillation when it experiences a particular phase of a Fourier component of the satellite’s forcing. This enables continued coherent kicks from the satellite to build up the particle’s radial motion, and significant forced oscillations may thus result. In this way, spiral density waves—horizontal density oscillations which result from the bunching of streamlines of gas molecules on eccentric orbits—are launched. If the perturbing planet is inclined relative to the disk, resonances involving the vertical frequencies will also arise, and these will launch spiral bending waves, which are vertical corrugations of the disk plane resulting from the induced coherent inclinations of gas molecule orbits. We do not consider vertical resonances further, since they do not directly affect planet migration; the topic is investigated in detail by, e.g., Lubow and Ogilvie (2001).

4. Saturn’s Rings - Observational Tests of Disk-Satellite Interactions

Several of Saturn’s numerous moons orbit near or within the planet’s spectacular main ring system. These moons excite spiral density waves at resonant locations within the rings of Saturn. Gaps (or ring edges) are produced at strong resonances and close to moons where resonances overlap. Pan, a small moon within the A ring, has cleared a gap around its orbit, and Pandora and Prometheus, somewhat larger moons orbiting just exterior Saturn’s main ring system, confine particles to the narrow, irregular F ring. Spiral density waves and sharp ring boundaries were detected in Saturn’s rings by four instruments on the Voyager spacecraft. These data provide for valuable tests of models of interactions between secondaries and astrophysical disks.

4.1. Wave characteristics and ring properties

Several dozen density waves in Saturn’s rings have thus far been identified with exciting resonances and analyzed to determine the local surface mass density of the rings (e.g., Esposito et al. 1984, Rosen et al. 1991). The observed and predicted resonance locations agree within observational uncertainties, which are generally less than 1 part in $10^4$. The surface density, $\sigma$, at most wave locations in the optically thick A and B rings is of order 50 g/cm$^2$. Measured values in the optically thin C ring are $\sigma \sim 1$ g/cm$^2$; an intermediate value of $\sigma \sim 10$ g/cm$^2$ has been estimated for Cassini’s Division.

The outer edges of the B and A rings are maintained by the Mimas 2:1 and Janus 7:6 resonances, which are the strongest resonances within the ring system (Smith et al. 1981, Holberg et al. 1982, Porco et al. 1984a, Lissauer and Cuzzi 1982, Borderies et al. 1982). Nearly empty gaps with embedded optically thick ringlets have been observed at strong resonances located in optically thin regions of the rings (Holberg et al. 1982). These features are probably caused by a resonance-related process; however, no explanation for the embedded ringlets currently exists. Nearly empty gaps with embedded ringlets have also been observed at nonresonant locations (Porco et al. 1984b).
Although resonance features in Saturn’s rings are well understood in many respects, there remain several major outstanding issues. Some of these problems are related to angular momentum transport, a key factor in planetary migration.

The damping behavior of most observed spiral waves differs significantly from that predicted by the simple fluid approximation in a constant viscosity disk (Goldreich and Tremaine 1978). This has led to severe problems with attempts to estimate the viscosity of the rings using observations of wave damping. Some of the factors which can influence the damping behavior of waves are variations in the background surface density and velocity dispersion, interference with other waves and wave nonlinearities (Lissauer et al. 1984). The damping of nonlinear density waves has been studied in detail by Shu et al. (1985), who find that damping rates can be very sensitive to particle collision properties and to optical depth. Thus, the anomalous damping behavior of many spiral waves presents both a challenge and an opportunity for researchers attempting to deduce ring properties other than surface mass density from the study of nonlinear waves (Longaretti and Borderies 1986). Spiral waves carry (positive or negative) angular momentum, and deposit it in regions of the disk in which they damp, so the poor correspondence between theory and observation is of concern to planet migration modelers.

The cause of the observed enhancement of material in regions where strong waves propagate is poorly understood. Density waves excited at inner Lindblad resonances carry negative angular momentum, i.e., the angular momentum of the ring particles is temporarily reduced by the passage of these waves. When such waves damp, particles drift inwards. Resonant removal of angular momentum causes sharp outer ring edges and gaps to be produced at the strongest resonances within the ring system (Borderies et al. 1982). Waves are observed to be excited at the strongest resonances which do not produce gaps. However, waves do not appear to deplete material from the regions in which they propagate; on the contrary, a surface density enhancement is often observed in such regions (Holberg et al. 1982, Longaretti and Borderies 1986, Rosen et al. 1991).

The problem of “dredging” of ring material due to wave damping has never been solved in a self-consistent manner, in which the evolution of a region due to wave propagation, wave damping and general ring viscosity is following until a quasi-steady state is attained. (The most detailed study of this problem thus far attempted is presented by Borderies et al. 1986.) Damping of the wave in the outer portion of the region in which it propagates could bring material towards resonance, but what stops (or at least slows) this material from further inward drift? The answer to this question may be relevant to two of the major questions facing ring theorists today: What maintains inner edges of the major rings of Saturn? and Do the short timescales for ring evolution due to density wave torques imply Saturn’s rings are much younger than the planet itself?

4.2. Migration of moons

Before Voyager arrived at Saturn, Goldreich and Tremaine (1978) predicted that torques due to density waves excited by the moon Mimas at its 2:1 resonance had removed sufficient angular momentum from ring material to have cleared out Cassini’s Division, a 4000 km wide region of depressed surface mass density located between the broad high density A and B rings. Although the hypothesis of density waves clearing Cassini’s Division remains unverified, Voyager found a multitude of density waves excited by small newly-discovered satellites orbiting near the rings. Although the torque at these individual resonances is less than that at Mimas’ 2:1 resonance, the sum of their torques is much greater. Moreover, the waves are observed and amplitudes agree with theory to within a factor of order unity.

Goldreich and Tremaine (1982) pointed out that the back torque the rings exert on
the inner moons causes them to recede on a timescale short compared to the age of the Solar System; estimates suggest that all of the small moons orbiting inside the orbit of Mimas should have been at the outer edge of the A ring within the past $10^8$ years, with Prometheus’ journey outward occurring on a timescale of only a few millions years; more recent, lower estimates for the masses of these moons increase these timescales by a factor of a few, but do not solve the problem. Resonance locking to outer more massive moons could slow the outward recession of the small inner moons; however, angular momentum removed from the ring particles should force the entire A ring into the B ring in $\sim 10^9$ yr (Lissauer et al. (1984) and Borderies et al. (1984) quote somewhat shorter times based on larger masses of the moons).

If the calculations of torques are correct, and if no currently unknown force counter-balances them, then small inner moons and/or the rings must be new, i.e., much younger than the age of the Solar System. Both of these possibilities appear to be a priori highly unlikely. Rings could be remnants of Saturn’s protosatellite disk which never accreted into moon-sized bodies due to the strong tidal forces of Saturn inside Roche’s limit, in which case they would be $\sim 4.5 \times 10^9$ yr old. Alternatively, Saturn’s ring particles could be part of the debris from a moon that was collisionally disrupted, in which case they would most likely date from the first $\sim 10^9$ years of the Solar System, when much more debris large enough to cause such a disruption was available than is today (Lissauer et al. 1988), or from a tidally disrupted giant comet, although this possibility is also a priori unlikely (Dones 1991). Recent accretion of the inner moons within the ring system may be possible (Borderies et al. 1984), but why did such accretion only occur during the past $\sim 10^8$ yr? For these reasons, the issue of short timescales due to density wave torques is a major outstanding problem in the field of planetary rings.

### 5. Migration of Extrasolar Planets

#### 5.1. Type I migration

Goldreich and Tremaine (1980) pointed out that interactions with the gas disk could move planets large distances during the lifetime of the gas. Relative to the location of the planet, the inward/outward-propagating density waves carry a net negative/positive flux of angular momentum. If this angular momentum is deposited in the disk (as opposed to reflecting at the disk boundaries and returning to the planet), there will be a positive/negative torque on the planet from the interior/exterior parts of the disk. Variations in disk properties ought, in general, to produce a mismatch between the net interior and exterior torques, and a resultant migration rate for the planet of order

$$v_1 = k_1 \frac{M}{M_*} r \Omega \frac{\Sigma_d r^2}{c} \left( \frac{r \Omega}{c} \right)^3,$$

(5.1)

where $k_1$ is a measure of the torque asymmetry, $M$ is the mass of the planet, $M_*$ is the mass of the primary, $r$ is the distance from the primary, $\Sigma_d$ is the disk surface density, and $c$ is the gas sound speed. From dimensional arguments, $k_1$ should scale with the disk aspect ratio $h/r$ (Goldreich and Tremaine 1980). This type of orbital drift is commonly referred to as Type I migration. Ward (1997) showed that the outer torque ought to dominate in general, resulting in orbital decay for the planet.

#### 5.2. Gap opening and Type II migration

The migration rate continues to increase linearly with planet mass as per Eq. (5.1), until the torque saturates and the planet pushes the inner and outer parts of the disk
apart, opening a gap for itself. Having thus established a (perhaps imperfect) barrier to the passage of disk material across its orbit, it is then (more or less) locked to the viscous evolution of the disk, i.e., carried inward along with the net inward flow of the disk material—assuming, of course, that this is all happening in a part of the disk with nonzero viscosity. This type of orbital evolution is called Type II migration. The characteristic velocity for this mode of migration is

$$v_{\text{II}} = k_2 \alpha r \left( \frac{c}{r \Omega} \right)^2,$$

where $k_2$ is a constant of order unity.

It is unclear when the formation of a gap occurs; this depends on both the viscosity of the disk, and on the way in which the density waves are dissipated. Lin and Papaloizou (1993) required the density waves to be strongly nonlinear as soon as they are launched, so that they immediately shock and deposit their angular momentum in the disk. They showed that the minimum planetary mass to accomplish this, in the limit of zero disk viscosity, is such that the Hill (or Roche) radius of the planet, defined as

$$r_H = \left( \frac{M}{3M_\odot} \right)^{1/3} r,$$

is equal to the half-thickness of the disk, $h$. This criterion—often called the thermal condition—yields a gap-opening mass of

$$M_{\text{thermal}}^\text{crit} \sim M_* \left( \frac{h}{r} \right)^3 \sim 100 \left( \frac{r}{1 \text{AU}} \right)^{3/4} \text{M}_\odot,$$

where the numerical estimate is obtained using a Solar-mass primary and the standard Hayashi (1981) nebula model, which has a half-thickness of

$$h = 0.0472 (a/1 \text{AU})^{-5/4} \text{AU},$$

and a minimum surface density—obtained from smoothly spreading out the mass contained in the planets and enhancing the gas content to solar abundance, of

$$\Sigma_{\text{min}} = 1.7 \times 10^3 \left( \frac{a}{1 \text{AU}} \right)^{-3/2} \text{g/cm}^2.$$

In a disk with nonzero viscosity, an additional condition for maintaining a gap is that the rate of angular momentum transfer across the gap exceed the intrinsic viscous angular momentum transport rate of the disk (Lin and Papaloizou 1993, Bryden et al. 1999). This leads to a critical mass of

$$M_{\text{viscous}}^\text{crit} \sim 40 M_* \alpha \left( \frac{h}{r} \right)^2 \sim 300 \left( \frac{\alpha}{10^{-2}} \right) \left( \frac{r}{1 \text{AU}} \right)^{1/2} \text{M}_\odot.$$

In a viscous disk, the minimum gap opening mass is then \(\text{Min}[M_{\text{thermal}}^\text{crit}, M_{\text{viscous}}^\text{crit}]\).

On the other hand, Ward and Hourigan (1989) assumed that damping is linear and independent of the perturber’s mass. In this way, they obtained a much smaller critical mass of

$$M_{\text{inertial}}^\text{crit} \sim \frac{h^3 \Sigma}{r} \sim 0.007 \left( \frac{r}{1 \text{AU}} \right)^{5/4} \text{M}_\odot,$$

not for gap-opening, but for the planet to induce a density contrast in the disk—essentially to pile up disk mass ahead of itself—that is strong enough to stall its migration. They referred to this as the inertial mass.

The result of Rafikov (2002) is intermediate between \(M_{\text{thermal}}^\text{crit}\) and \(M_{\text{inertial}}^\text{crit}\). In this
model, density waves travel some distance in the disk before dissipating via weak non-linearity. The critical mass obtained in this way is around $2 \, M_\oplus$ at 1 AU in an inviscid minimum-mass disk.

5.3. Migration versus formation?

Type I migration timescales can become as short as $\sim 10^4$ years for a planet of mass $\sim 10 \, M_\oplus$ in a sufficiently viscous disk (Ward 1997), two to three orders of magnitude less than the lifetime of the gas disk. In the core accretion model, the critical mass needed to initiate runaway gas accretion is thought to be of order $10 \, M_\oplus$; the formation timescale for a body of this size is of order a million years or more (e.g., Lissauer 1987, Lissauer and Stewart 1993, Weidenschilling 1998, Kokubo and Ida 2000, Thommes et al. 2002b). So, unless gap formation/stalling occurs at masses far below $10 \, M_\oplus$, a growing core would spiral into its parent star long before it got large enough to accrete a massive atmosphere. It has been proposed that rapid migration could actually speed up the accretion process (Ward 1986), in the same way described above for the formation of Uranus and Neptune. However, since formation timescales are shorter at smaller radii, an inward-migrating protoplanet is likely to encounter not a pristine planetesimal disk, but instead the stirred-up remains of a disk which has itself already produced large protoplanets. Furthermore, simulations have shown that even in the idealized case, accretion efficiency is low and the disk has to be enhanced by at least a factor of five relative to minimum mass in order to allow a protoplanet starting at 10 AU to grow to $\sim 10 \, M_\oplus$ before it falls into the star (Tanaka and Ida 1999).

Clearly, rapid Type I migration constitutes a major potential problem for the core accretion model of giant planet formation. However, there may be other ways out, apart from the possibility that this phase ends quickly. For one thing, the analysis thus far has been restricted to a single body interacting with the disk; it is unclear whether the coupling to the disk remains as strong when multiple protoplanets are launching density waves in close proximity to each other. Also, Tanaka et al. (2002) demonstrated that reflection of the density waves at the outer edge of the disk could permit a non-migrating steady state to be attained without the formation of a gap.

5.4. Migration and the properties of extrasolar planets

Even after locking themselves in a gap, giant planets are likely to undergo significant migration (e.g., Ward 1997). The ensemble of extrasolar planets detected by radial velocity searches (Mayor and Queloz 1995, Marcy et al. 2000) provides considerable observational support for migration playing a large role in the formation of planetary systems. The most direct clue is the large number of planets on close-in orbits. Nearly half of the known planets orbit closer than 1 AU to their parent star. Although there is certainly a strong selection effect favoring detection of short-period planets, it is clear that a non-negligible fraction of giant planets somehow end up on such orbits. In situ formation is one possibility, but there are a number of problems in trying to construct such models; in particular, a very high surface density of solids and/or substantial planetesimal migration would be required to form a giant planet core-sized body (Bodenheimer et al. 2000). Formation at larger stellocentric distances followed by inward migration seems to provide a more natural explanation for planets like those orbiting 51 Peg and $\rho$ CrB.

Migration may also have facilitated interactions between planets in some of the detected multiple-planet systems. The two planetary companions of Gliese 876 are in a 2:1 mean-motion resonance with each other (Marcy et al. 2001). Lee and Peale (2002) showed that capture into this resonance would occur if the two planets were originally farther apart in orbital period, and were induced to migrate toward each other, assuming their
eccentricities remained low throughout ($e \sim 10^{-2}$ or less). Bryden et al. (2000) and Kley (2000) performed hydrodynamic simulations of two gap-opening planets embedded in a gas disk, and demonstrated that, if the planets’ orbits are sufficiently close together, they will tend to clear the annulus of gas between them on timescales as short as thousands of years. The result is that both planets end up sharing a gap, and are pushed toward each other by the inner and outer parts of the disk. Aside from the Gliese system, the two planets discovered in HD 82943 appear to also be in a 2:1 mean motion resonance; numerical simulations suggest that, given the inferred orbital parameters, only resonant configurations are stable (Jianghui et al. 2002). Furthermore, the periods of the inner two companions of the putative three-planet system 55 Cnc are very close to a 3:1 commensurability (Marcy et al. 2002). Capture into this resonance, and other higher-order resonances, can occur if initial eccentricities of the convergently migrating planets are nonzero.

Divergent migration in multiple-planet systems is another possible mechanism for reproducing some of the observed properties of extrasolar planets. Chiang et al. (2002) show that if planets on initially circular orbits move apart in orbital period, the noncapturing resonance passages they encounter can induce significant eccentricities in the orbits of both bodies. Such a mechanism may help to account for the generally large eccentricities of those extrasolar planets with orbital radii greater than a few hundredths of an AU, out of range for tidal circularization by their parent star (e.g., Lin et al. 2000). Divergent migration may come about if there is a sufficiently steep gradient in disk viscosity, so that the inner planet migrates faster than the outer one. In particular, with reference to the model of Gammie (1996), if one planet is interior to the collisional ionization radius, where viscosity will be high, and the other is exterior to it, where viscosity is low, divergent migration may result. However, it is required that the gas annulus between the two planets persist, notwithstanding the results cited above.

5.5. Type III migration

In a low-mass disk—one in which the giant planet mass is comparable to or greater than the disk mass—the planet’s inertia will slow the evolution of the disk. Thus the planet’s migration speed will be inversely proportional to its mass:

$$v_{III} \sim \frac{M_{\text{disk}}}{M_{\text{planet}} + M_{\text{disk}}} v_{III}$$

At the same time, accretion of disk material onto the star will trail off as the part of the disk interior to the planet drains onto the star. This mode can be referred to as Type III migration. Observations hint that something like this may be happening in TW Hydra, a ten million year old T Tauri star (Calvet et al. 2002). TW Hydra’s accretion rate is very low compared to younger T Tauri stars, and at the same time, the spectral energy distribution of its infrared excess is best fit by a disk with a sharp inner edge at 4 AU, perhaps signifying truncation by a giant planet. Interestingly, however, TW Hydra’s disk mass is estimated to be quite large, $\sim 0.6 M_*$. A 10 million year old disk still containing this much mass implies a low disk viscosity, whether the disk is simply accreting slowly, or whether it is indeed being kept at bay by a planet. Calvet et al. estimate $\alpha < 10^{-3}$.

6. Conclusions

The longstanding view of planet formation as an orderly process, involving little radial migration of material, is being made to look increasingly inaccurate by both observational and theoretical findings. In our own Solar System, the high eccentricities of objects in
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exterior mean-motion resonances with Neptune imply an outward migration of several AU by the outer ice giant; modeling suggests that Uranus, and to a lesser degree Jupiter and Saturn, likewise underwent migration as they cleared the surrounding planetesimal disk. An earlier and more violent period of migration could have occurred if Uranus and Neptune originally formed among proto-Jupiter and -Saturn; such a model would alleviate the longstanding formation timescale problem of Uranus and Neptune, and could simultaneously help to account for the gravitationally stirred-up state of the Kuiper belt.

In the first ten million years or so of a planetary system’s life, the nebular gas is still present and provides a much larger sink/source of angular momentum than the planetesimal disk. Growing protoplanets exchange angular momentum by launching density waves at resonance sites in the disk. From theoretical consideration, this ought to bring about a net loss of angular momentum and rapid orbital decay—Type I migration—for bodies of order a few Earth masses. For a sufficiently massive body, the torques between it and the disk will be strong enough to open a gap, thus locking the body into the subsequent viscous evolution of the disk in what is called the Type II mode of migration. The core accretion model of giant planet formation seems to require that the Type I to II transition occur at small masses, otherwise growing cores will spiral into the star before they can acquire a massive envelope. Alternatively, it is possible that planet formation simply is an enormously wasteful process, which dumps a steady stream of growing protoplanets onto the primary, and the end result is whatever happens to be left over when the gas fades away. This is sometimes called the “last of the Mohicans” scenario. Significant post-formation migration is quite likely responsible for the large number of planets detected on close-in orbits (“giant Vulcans”, also referred to as “hot Jupiters”). In multiple-planet systems, convergent and divergent migration of planets can be invoked to explain, respectively, resonant capture and eccentricity excitation. As the nebular gas dissipates, it is likely that the tables are eventually turned; the planets, heretofore at the mercy of the gas, assert themselves and serve as anchors to slow down the viscous evolution of the last remains of the disk, so that migration ends in a Type III phase. Clearly, the present observational and theoretical “state of the art” still requires us to use a liberal amount of conjecture in attempting to sketch a coherent picture of planet migration. However, it seems equally clear that migration is intimately linked with the formation of planetary system, and a complete picture of the latter will require a full understanding of the former.

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