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GIANT PLANET FORMATION

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Giant planet formation is closely interrelated with star formation, protoplanetary disks, the growth of dust and solid planets in those nebula disks and finally nebula dispersal. Models of the interiors and evolution of giant planets in our Solar System point to a bulk enrichment of heavy elements more than a factor of two above solar composition and imply heavy element cores ranging from greater than one Earth mass to a considerable fraction of the total mass. Detailed models of giant planet formation explain the diversity of Solar System and extrasolar giant planets by variations in the core growth rates caused by a coupling of the dynamics of planetesimals and the contraction of the massive envelopes they dive into, as well as by changes in the hydrodynamical accretion behavior of the envelopes resulting from differences in nebula density, temperature and orbital distance.

I. INTRODUCTION

Our four giant planets contain 99.5% of the angular momentum of the Solar System, but only 0.13% of its mass. On the other hand, more than 99.5% of the mass of the planetary system is in those four largest bodies. The angular momentum distribution can be understood on the basis of the 'nebula hypothesis' (Kant 1755). The nebula hypothesis assumes concurrent formation of a planetary system and a star from a centrifugally-supported flattened disk of gas and dust with a pressure supported, central condensation (Laplace 1796, Safronov 1969, Lissauer 1993). Theoretical models of the collapse of slowly rotating molecular cloud cores have demonstrated that such preplanetary nebulae are the consequence of the observed cloud core conditions and the hydrodynamics of radiating flows, provided there is a macroscopic angular momentum transfer process (chapter by Stone et al., Cassen and Moosman 1981, Morfill et al. 1985, Laughlin and Bodenheimer 1994, Podosek and Cassen 1994). Assuming turbulent viscosity to be that process, dynamical models have shown how mass and angular momentum separate by accretion through a viscous disk onto a growing central protostar (Tscharnuter 1987, Tscharnuter and Boss 1993; see chapter by Stone et al. for viscosity mechanisms). Those calculations, however, do not yet reach to the evolutionary state of the nebula where planet formation is expected. Observationally-inferred disk sizes and masses are overlapping theoretical expectations and fortify the nebula hypothesis. High resolution observations at millimeter-wavelengths are now sensitive to disk conditions at orbital distances $> 50\,\mathrm{AU}$ (see, e.g., chapter by Wilner and Lay, Dutrey et al. 1998, Guilloteau and Dutrey 1998). However, observations thus far provide little information about the physical conditions in the respective nebulae on scales of 1 to 40 AU, where planet formation is expected to occur.

Planet formation studies therefore obtain plausible values for disk conditions from nebulae that are reconstructed from the present planetary system and disk physics. The so obtained 'minimum reconstituted nebula masses' *defined as the total mass of solar composition material needed to provide the observed planetary/satellite masses and compositions by condensation and accumulation*, are a few percent of the central body for the solar nebula *and* the circumplanetary protosatellite nebulae (Kusaka et al. 1970, Hayashi 1980, Stevenson 1982a). The total angular momenta of the satellite systems, however, are only about 1% of the spin angular momentum of the respective giant planet (Podolak et al. 1993), in strong contrast with the planetary system/Sun ratio. Assembling planets from a nebula disk and advecting the angular momentum due to keplerian shear until the present giant planet masses are reached results in total angular momenta overestimating the present spin angular momenta of the giant planets only by small factors (Götz 1993). Even if giant planets would have kept this angular momentum, they still would not rotate critically! Giant planets, unlike stars, therefore do not have an angular momentum problem. This may justify why most studies of proto-giant planets neglect rotation or treat it as a small perturbation.

We thoroughly discuss new results on interior models in Section II. We review recent work on planetesimal formation and growth of solid planets in Section III. The 'nucleated instability hypothesis' is the only model for the formation of Uranus and Neptune at the moment, while other models also exist for Jupiter and Saturn; in Section IV, we review these various models. We put emphasis on envelope evolution and gas accumulation using the 'nucleated instability' model in Section V. We apply the formation theories to extrasolar planets in Section VI.

II. INTERIORS OF THE GIANT PLANETS

Our knowledge of the mechanisms that led to the formation of the giant planets is essentially based on numerical models and on the constraints provided by studies of the internal structure and composition of Jupiter, Saturn, Uranus and Neptune. This involves the calculation of interior models matching the observed gravitational field. Each of the four giant planets of our Solar System are thus believed to consist of a central, dense core, and a surrounding envelope composed of hydrogen, helium and small amounts of heavy elements. The cores of Jupiter and Saturn are very small compared to the total masses of the planets, whereas Uranus and Neptune are mostly 'core' and possess small (i.e., low mass) envelopes.

The giant planets, with the exception of Uranus, emit significantly more energy than received from the Sun, a consequence of their progressive cooling and contraction. Two important consequences can be drawn from this: (i) they have inner temperatures of a few thousand Kelvins or more, and therefore, their hydrogen-helium envelopes are *fluid*; (ii) they are mostly convective (see Hubbard 1968, Stevenson and Salpeter 1977). The convective hypothesis has been challenged (Guillot et al. 1994a), but the regions where convection could be suppressed due to radiative transport are limited to a small fraction of the envelope, at temperatures between 1500 and 2000 K, or in low-temperatures regions where the abundance of water is small. These two conclusions are also expected to hold for Uranus for essentially two reasons: first, it is highly unlikely that its interior has cooled much more than that of Neptune (thus, one can expect that its intrinsic heat flux is small but larger than zero); second, it possesses a magnetic field of similar strength to the other giant planets, a sign of convective activity in its interior. It seems, therefore, logical to assume that the envelopes of all four giant planets are homogeneously mixed. Some caveats are necessary, however: (i) condensation and chemical reactions alter chemical composition (these should be confined to the external regions); (ii) a first order phase transition (such as one between molecular and metallic hydrogen) imposes an abundance discontinuity across itself; (iii) hydrogen-helium phase separation might occur and lead to a variation of the abundance of helium in the planet; (iv) the envelopes of Uranus and Neptune are small and enriched in heavy elements; it is thus conceivable that molecular weight gradients inhibit convection and yield non-homogeneous envelopes.

On this basis, a 3-layer structure is generally adopted for the four giant planets (Figure 1). In the case of the less massive Uranus and Neptune, in which hydrogen is believed to remain in molecular phase, the planets are divided into a 'rock' core (a mixture of the most refractory elements including silicates and iron), an 'ice' layer (consisting of H_2O, CH_4, NH_3) and a hydrogen-helium envelope. The latter is substantially enriched in heavier elements, as demonstrated by the ~ 30 times solar enrichment in carbon, in the form of CH_4, spectroscopically measured in the tropospheres of Uranus and Neptune (e.g., Fegley et al. 1991; Gautier et al. 1995). Other elements, in particular oxygen (mostly in the form of H_2O), are also believed to be substantially enriched compared to a solar composition mixture, but are hidden deep in the atmosphere because of condensation.

Although the two planets share many similarities (mass, magnetic field, atmospheric structure), several factors point towards some differences in their internal structure: Uranus emits scarcely more energy than received from the Sun (whereas Neptune possesses a very significant intrinsic heat flux), and Uranus's gravitational field indicates that it is more centrally condensed (see Hubbard et al. 1995). Furthermore, 3-layer models of these planets that assume homogeneity of each layer and adiabatic temperature profiles succeed in reproducing Neptune's gravitational field but not that of Uranus (Podolak et al. 1995). The difficulty is circumvented by using slightly reduced (by $\sim 10\%$) densities in the ice layer, which is interpreted either as hydrogen mixed to the ice, or as higher temperatures (superadiabatic temperature gradients). Both explanations imply that substantial parts of the planetary interior are not homogeneously mixed. The existence of such compositional gradients could also explain the fact that Uranus's heat flux is so small: its heat would not be allowed to escape to space by convection, but through a much slower diffusive process in the regions of high molecular weight gradient. Such regions would also be present in Neptune, but much deeper, thus allowing more heat to be transported outward (Podolak et al. 1991). This could also explain the fact that the magnetic fields of these two planets possess a very significant quadrupolar component, by allowing a hydromagnetic dynamo to form only in a relatively thin shell rather than in a sphere (Ruzmaikin and Starchenko 1991; Hubbard et al. 1995).

The existence of these non-homogeneous regions is further confirmed by the fact that if hydrogen is supposed to be confined solely to the hydrogen-helium envelope, models predict ice/rock ratios of the order of 10 or more, much larger than the protosolar value of ~ 2.5. On the other hand, if we impose the constraint that the ice/rock ratio is protosolar, the overall composition of both Uranus and Neptune is, by mass, about 25% rock, $60-70\%$ ices, and $5-15\%$ hydrogen and helium (Hubbard and Marley 1989; Podolak et al. 1991, 1995; Hubbard et al. 1995). The formation of these non-homogeneous regions is certainly contemporaneous with the accretion of the planets (Hubbard et al. 1995). The importance of stochastic processes during that epoch is shown by the 98° obliquity of Uranus, a strong sign that giant impacts shaped the actual structure of these ice giants (Lissauer and Safronov 1991, cf., however, Tremaine 1991 for an alternative explanation of giant planet obliquities).

The structure of the much more massive Jupiter and Saturn, which are mostly formed from hydrogen and helium, is comparatively simpler. Most interior models (Hubbard and Marley 1989, Chabrier et al. 1992, Guillot et al. 1994b) of these planets assume a three-layer structure: a core, an inner envelope where hydrogen is in metallic phase, and an outer one where hydrogen is mostly in the form of H_2. More complex models (e.g., Zharkov and Gudkova 1991) can be calculated, but these further divisions into multiple layers do not qualitatively affect the main results.

Each layer is assumed to be globally homogeneous (i.e., neglecting condensation and chemical reactions), a consequence of efficient mixing by convection. Because less helium is observed in the external layers of Jupiter and Saturn than was present in the protosolar nebula (von Zahn and Hunten 1996; Gautier and Owen 1989), it is believed that the metallic regions of these planets contain more helium than the molecular ones. The difference is thought to be due either to a first order molecular/metallic phase transition of hydrogen, or to a hydrogen-helium phase separation, or both (e.g., Stevenson and Salpeter 1977; Hubbard and Marley 1989). Both phenomena are expected to occur in similar regions (e.g., Stevenson 1982b), and therefore we do not differentiate one from the other. We emphasize, however, that a first order transition, such as that suggested by Saumon and Chabrier (1989), would lead to a discontinuity of abundance of all chemical elements, whereas a phase *separation* would principally affect helium and other minor species, such as neon, that tend to be dissolved into helium droplets (unless, e.g., water is present in large enough abundances and can also separate from hydrogen). The lack of neon measured by the Galileo probe (Niemann et al. 1996) suggests that, in Jupiter, helium phase separation has begun (Roulston and Stevenson 1995), and that, consequently, it also occurs in Saturn, which is colder.

With these hypotheses, Guillot et al. (1997) and Guillot (1999) calculate the ensemble of interior models of Jupiter and Saturn that match the gravitational moments within the error bars of the measurements. Using the inferred mass mixing ratio of helium in the protosolar nebula and the presently observed one in the atmospheres of Jupiter and Saturn, they retrieve the possible abundances of heavy elements in the metallic and molecular regions. Their calculations include uncertainties in the hydrogen-helium and heavy elements equations of state, in the inner temperature profile (convective/radiative) and regarding the internal rotation (solid/differential).

The resulting constraints on the core mass and total mass of heavy elements in Jupiter, Saturn, Uranus and Neptune are summarized on Figure 2. (The cases of Uranus and Neptune are relatively trivial; these planets contain little hydrogen and helium.) A first result is that the gravitational fields of Jupiter and Saturn do not necessarily imply that these planets have ice/rock cores. In the case of Jupiter, models without a core are obtained only in the case of the less favored interpolated equation of state of hydrogen, whose calculation is not completely thermodynamically consistent (see Saumon et al. 1995). In the case of Saturn, it is difficult to distinguish between heavy elements in the core from those in the metallic region, hence yielding an even larger uncertainty in the core mass. As a result, Jupiter has a core whose mass lies between 0 and 12 $M_\oplus$, and Saturn's core is between 0 and 15 $M_\oplus$. We stress, however, that larger core masses are possible if gravitational layering occurs and the cores, possibly eroded by convective mixing, extend into the metallic envelope.

A second result concerns the total mass of heavy elements. The models of Saturn show that the planet is significantly enriched in heavy elements, by a factor of 10 to 15 compared to the solar value (corresponding to 20 to 30 $M_\oplus$, including the core), and by at least a factor of 5 when considering only the envelope. The constraints are much weaker in the case of Jupiter, due to the larger metallic region, where the equation of state is considerably more uncertain. Figure 2 shows that Jupiter contains 10 to 42 $M_\oplus$ of heavy elements, implying that it is moderately to significantly enriched in heavy elements compared to the protosolar nebula.

Recent interior models calculated with different assumptions (Hubbard and Marley 1989; Zharkov and Gudkova 1991; Chabrier et al. 1992) generally predict core masses and heavy elements abundances that fall within the ranges given in Figure 2. Larger core masses (10 – 30 $M_\oplus$) were found in previous calculations (see Stevenson 1982b for a review), but the largest core masses also yielded helium mass fractions well below the protosolar value and therefore are unrealistic. The main reason for the discrepancy with today's values is, however, that the calculation of core masses, especially in the case of Jupiter, is very sensitive to changes in the equation of state. At present, we can only *hope* that advances in our understanding of the behavior of hydrogen and helium at high pressures have led us in the right direction, but progress in compression experiments on liquid deuterium (Weir et al. 1996; Collins et al. 1998) should allow us to check that assertion in the near future.

In the case of Jupiter and Saturn, further constraints on today's internal structure can be sought from evolution models accounting for the progressive sedimentation of helium (Hubbard et al. 1999; Guillot 1999). Models with small cores tend to require a more pronounced helium differentiation and hence yield longer cooling times. The time to cool to the present temperature is constrained by the age of the Solar System (4.56 Gyr). Some static solutions are thus ruled out. Figure 2 shows that an upper limit of 10 $M_\oplus$ is obtained for Jupiter's core, and that Saturn's core mass lies between 6 and 15 $M_\oplus$.

Finally, observations of the atmospheric composition give two important clues: First, the C/H ratio is steadily increasing from Jupiter to Neptune, a fact that has to be explained by formation models (we refer the reader to the review by Podolak et al. 1993 for further details). Second, the D/H isotopic ratios recently measured in Jupiter by the Galileo probe (Mahaffy et al., 1998), and in Saturn by ISO (Griffin et al., 1996) are, within the error bars, consistent with the protosolar value derived from $^3He/^4He$ in the solar wind, namely $2.1 \pm 0.5 \times 10^{-5}$ (Geiss and Gloecker, 1998), whereas they are about three times larger in Uranus and Neptune (Feuchtgruber et al. 1999). This has to be compared to the D/H values measured in comets Halley, Hyakutake and Hale-Bopp, which are all about ten times larger than the protosolar value (Bockelée-Morvan et al. 1998). If Uranus and Neptune contained mixtures of comet-like ices and protosolar H_2 which were isotopically homogenized within these planets, their large ice fractions would have produced a more deuterium-rich atmospheric composition than that observed. Thus, either a significant isotopic exchange between vaporized ices and hydrogen took place in an early hot turbulent solar nebula (Drouart et al., 1999), or Uranus and Neptune formed from high D/H, comet-like, ices that have never been fully mixed with hydrogen in their interiors. The precise determination of D/H in the giant planets is thus an important tool for constraining their formation. We leave a more thorough discussion of this problem to the chapter by Lunine et al.

III. FORMATION OF PLANETESIMALS AND GROWTH OF SOLID PLANETS

A. Formation of Planetesimals

Even a very slowly rotating molecular cloud core has far too much rotational angular momentum to collapse down to an object of stellar dimensions, so a significant fraction of the material in a collapsing core falls onto a rotationally-supported disk in orbit about the pressure-supported star. Such a disk has the same elemental composition as the growing star, that is primarily H and He, with $\sim 1-2\%$ heavier elements. Sufficiently far from the central star, it is cool enough for some of this material to be in solid form, either remnant interstellar grains or condensates formed within the disk. Dust agglomerates via inelastic collisions and gradually settles towards the disk midplane as it grows large enough to be able to drift relative to the surrounding gas (Weidenschilling and Cuzzi 1993).

Although to a first approximation the gas in the disk is centrifugally supported in balance with the star's gravity, negative radial pressure gradients provide a small outwardly-directed force which acts to reduce the effective gravity, so the gas rotates at slightly less than the keplerian velocity. Small solid bodies (dust grains) rotate with the gas. Large solid bodies orbit at the keplerian velocity, and medium-sized particles move at a rate intermediate between the gas velocity and the keplerian velocity; thus macroscopic solid bodies are subjected to a headwind from the gas (Adachi et al. 1976). This headwind removes angular momentum from the particles, causing them to spiral inwards towards the central star. This inward drift can be very rapid, especially for particles whose coupling time to the gas is similar to their orbital period; smaller particles drift less rapidly because the headwind they face is not as strong, whereas large particles drift less because they have a greater mass-to-surface-area ratio. Orbital decay times for meter-sized particles at 1 AU from the Sun have been estimated to be only $\sim$ 100 years (Weidenschilling 1977). The large radial velocities of bodies in this size range relative to both larger and smaller particles implies frequent collisions, so it is possible that most solid bodies grow through the critical size range quickly without substantial radial drift. However, it is also possible that a large amount of solid planetary material is lost from the disk in this manner.

Solid bodies larger than $\sim$ 1 km in size face a headwind only slightly faster than meter-size objects (for parameters thought to be representative of the planetary region of the solar nebula), and because of their much greater mass-to-surface-area ratio they suffer far less orbital decay from interactions with the gas in their path. The growth of solid bodies from the meter-sized 'danger zone' to the kilometer-sized 'safe zone' could occur by collective gravitational instabilities in a thin solids subdisk (Safronov 1960, Goldreich and Ward 1973) in regions of protoplanetary disks that aren't too turbulent, or (more likely) via continued binary accretion (Weidenschilling and Cuzzi 1993). Kilometer-sized planetesimals appear to be reasonably safe from loss (unless they are ground down to smaller sizes via disruptive collisions) until some of these planetesimals grow into planetary-sized bodies.

B. Growth of Solid Planets

The primary perturbations on the keplerian orbits of kilometer-sized and larger bodies in protoplanetary disks are mutual gravitational interactions and physical collisions (Safronov 1969). These interactions lead to accretion (and in some cases erosion and fragmentation) of planetesimals. Gravitational encounters are able to stir planetesimal random velocities up to the escape speed from the largest common planetesimals in the swarm (Safronov 1969). The most massive planets have the largest gravitationally-enhanced collision cross-sections, and accrete almost everything with which they collide. If the random velocities of most planetesimals remain much smaller than the escape speed from the largest bodies, then these large 'planetary embryos' grow extremely rapidly (Safronov 1969, cf. Greenzweig and Lissauer 1990, 1992 for 3-body growth rates). The size distribution of solid bodies becomes quite skewed, with a few large bodies growing much faster than the rest of the swarm in a process known as *runaway accretion* (Greenberg et al. 1978, Wetherill and Stewart 1989, Kokubo and Ida 1996). Eventually, planetary embryos accrete most of the (slowly moving) solids within their gravitational reach, and the runaway growth phase ends. Planetary embryos can continue to accumulate solids rapidly beyond this limit if they migrate radially relative to planetesimals as a result of interactions with the gaseous component of the disk (Tanaka and Ida 1999).

The eccentricities of planetary embryos in the inner Solar System were subsequently pumped up by long-range mutual gravitational perturbations; collisions between these embryos eventually formed the terrestrial planets (Wetherill 1990, Chambers and Wetherill 1998). However, time scales for this type of growth in the outer Solar System (at least 10^8 years, Safronov 1969) are longer than the lifetime of the gaseous disk (cf., Lissauer et al. 1995). Moreover, unless the eccentricities of the growing embryos are damped substantially, embryos will eject one another from the star's orbit (Levison et al. 1998). Thus, runaway growth, possibly aided by migration (Tanaka and Ida 1999), appears to be the way by which solid planets can become sufficiently massive to accumulate substantial amounts of gas while the gaseous component of the protoplanetary disk is still present (Lissauer 1987).

Most models of the accumulation of giant planet atmospheres have assumed a constant accretion rate for planetesimals. The models of Pollack et al. (1996) calculate the planetesimal accretion rate together with that of gas; however, these models neglect growth of competing planetary cores as well as radial migration. Models of giant planet growth will improve once atmospheric accumulation models are coupled to sophisticated models of solid planet growth such as the multi-zoned numerical accretion code of Weidenschilling et al. (1997), and when radial migration of planetesimals and planets is better understood and included in the models.

IV. GAS ACCUMULATION THEORIES

The key problem in giant planet formation is that preplanetary disks are only weakly self-gravitating equilibrium structures supported by centrifugal forces augmented by gas pressure (see chapters by Stone et al., Hollenbach et al., Calvet et al., Beckwith et al.). Any isolated, orbiting object below the Roche density is pulled apart by the stellar tides. Typical nebula densities are more than two orders of magnitude below the Roche density, so compression is needed to confine a condensation of mass M inside its tidal or Hill-radius at orbital distance

a:

$$R_T = a \left(\frac{M}{3\mathrm{M}_\odot} \right)^{1/3} . \tag{1}$$

A local enhancement of self-gravity is needed to overcome the counteracting gas pressure. Giant planet formation theories may be classified by how they provide this enhancement: (i) the *nucleated instability* model relies on the extra gravity field of a sufficiently large solid core (condensed material represents a gain of ten orders of magnitude in density and therefore self-gravity compared to the nebula gas), (ii) a *disk instability* may operate on length scales between short scale pressure support and long scale tidal support, or (iii) an *external perturber* could compress an otherwise stable disk on its local dynamical time scales, e.g., by accretion of a clump onto the disk or rendezvous with a stellar companion. If the gravity enhancement is provided by a dynamical process as in the latter two cases, the resulting nebula perturbation (say of a Jupiter mass, $\mathrm{M_J}$, of material) is compressionally heated because it is optically thick under nebula conditions. Giant planet formation would then involve a transient phase of tenuous giant gaseous protoplanets, that would be essentially fully convective and contract on a timescale of $\sim 10^6$ yr (see Bodenheimer 1985). Another mechanism of forming stellar companions (iv) *fragmentation during collapse*, is plausible for binary stars and possibly brown dwarfs, but it is unlikely to form objects of planetary masses because opacity limits the process to masses above $\sim 10\,\mathrm{M_J}$ (cf. chapter by Bodenheimer et al. and Bodenheimer et al. 1993).

A. Nebula Stability

Preplanetary nebulae with minimum reconstituted mass are stable. Substantially more massive disks resulting from the collapse of cloud cores are self-stabilizing by transfer of disk mass to the stabilizing central protostar (Bodenheimer et al. 1993). Nevertheless, a moderate mass nebula disk might be found that can develop a disk instability leading to a strong density perturbation, especially when forced with a finite external perturbation. Giant gaseous protoplanets (GGPPs) might form when the instability has developed into a clump (DeCampli and Cameron 1979, Bodenheimer 1985). Boss (1997, 1998) has constructed such an unstable disk with $0.13\,\mathrm{M}_\odot$ within 10 AU and obtained maximum density enhancements (by a factor ~ 20) with $\sim 10\,\mathrm{M_J}$ above the background for a few orbital periods. (The density enhancement at the surface of a $1\,\mathrm{M}_\oplus$ core is between 10^5 to 10^7, for comparison.) These clumps, provided they are stable on a few cooling times, are candidates to become proto-giant planets via an intermediate state as tenuous GGPPs.

A key issue, as in any theory involving an instability of the disk gas, is then the *a posteriori* formation of a core. Only metals that are present initially would rain out to form a core, while material added later by impacts of small bodies after the GGPP had formed would be soluble in the envelope (Stevenson 1982a). Boss (1998) outlines how a core corresponding to the solar composition high-Z material ($6\,\mathrm{M}_\oplus$ and $2\,\mathrm{M}_\oplus$ for Jupiter- and Saturn-mass respectively) might form if the density enhancements are long lived, need no more pressure confinement and evolve into GGPPs that are *non-turbulent*. It should be noted here (see Sect. II) that although interior models of Saturn do not rule out the possibility that the planet has no core (or, equivalently a $2\,\mathrm{M}_\oplus$ core), this is not the favored solution. Also, GGPP models would probably predict that Jupiter should have a bigger core than Saturn, which is only marginally consistent with present interior models. Finally, Jupiter and certainly Saturn contain a lot of heavy elements (see Figure 2). To account for these bulk heavy element compositions, planetesimal accretion must occur anyway after the GGPPs have formed their cores.

If GGPPs need pressure confinement they also require the presence of an (undepleted) nebula and pose a lifetime constraint for the nebula, namely that nebula dispersal can only begin after a cooling time, that is $\sim 10^6\,\mathrm{yr}$ (Bodenheimer 1985). To determine whether GGPPs are convectively stable so that the non-turbulent core growth scenario can be applied, a detailed calculation of their thermal structure during contraction is necessary (DeCampli and Cameron (1979) found largely convective GGPPs).

One of us (GW) checked convective stability of GGPPs by a radiation hydrodynamical calculation. Alexander and Ferguson (1994) opacities and time-dependent MLT-convection were used in the description of energy transfer. The initial condition was a Jeans-critical nebula condensation of $\mathrm{M_J}$ and a temperature of $10\,\mathrm{K}$. Initially the GGPP had similar properties as Boss's (1998) banana-shaped density enhancements (mean density $8 \times 10^{-10}\,\mathrm{g\,cm^{-3}}$, central density $3.3 \times 10^{-8}\,\mathrm{g\,cm^{-3}}$). According to the new calculation the GGPP needs $1.8 \times 10^4\,\mathrm{yr}$ to contract into the tidal radius and is essentially fully convective from $< 100\,\mathrm{yr}$ to $2 \times 10^5\,\mathrm{yr}$, when a radiative zone spreads out from the planet's center.

B. Nucleated Instability

Planetesimals in the solar nebula are small bodies surrounded by gas. A rarefied equilibrium atmosphere forms around such objects. Early work in the nucleated instability hypothesis which assumes that such solid 'cores' trigger giant planet formation was motivated by the idea that at a certain critical core-mass the atmosphere could not be sustained and isothermal, shock-free accretion (Bondi and Hoyle 1944, Bondi 1952) would set in. Determinations of this critical mass were made for increasingly detailed description of the envelopes: adiabatic (Perri and Cameron 1974), isothermal (Sasaki 1989), isothermal-adiabatic (Harris 1978, Mizuno et al. 1978), radiative and convective energy transfer (Mizuno 1980). By then, modeling the formation and evolution of a proto-giant planet had become essentially a miniature stellar structure calculation with energy dissipation of impacting planetesimals replacing the nuclear reactions as the energy source. Present results on the critical mass are reviewed in the next section. Already, Safronov and Ruskol (1982) pointed out that even after the instability at the critical mass *the rate of gas accretion is determined not by the rate of delivery of mass to the planet [as in Bondi-accretion] but by the energy losses from the contracting envelope.* The planet's accretion rate is limited by the delivery of mass only when $M_P \gtrsim 100\,\mathrm{M}_\oplus$. Consequently, the energy budget of the envelope has been modeled more carefully taking into account the heat generated by gravitational contraction (quasi-hydrostatic models by Bodenheimer an Pollack 1986).

Major progress since PPIII has been made by a detailed treatment of planetesimal accretion to calculate the core growth rate and the capture, dissolution and sinking that determines how much and where in the envelope the planetesimal kinetic energy is liberated (Pollack et al. 1996). That made possible the first study of the coupling between gas accretion and solid accretion. Additionally, the description of the mechanics of contraction has been improved by hydrodynamic studies that determine the flow velocity of the gas by solving an equation of motion for the envelope gas in the framework of convective radiation-fluid-dynamics (e.g., Wuchterl 1993, 1995b, 1999). That allows the study of collapse of the envelope, accretion with finite Mach-number and an access to the study of linear adiabatic (Tajima and Nakagawa 1995) and nonlinear, non-adiabatic pulsational stability and pulsations of the envelope. Furthermore, the treatment of convective energy transfer has been improved by calculations using a time dependent mixing length theory of convection (Wuchterl 1995a, 1996, 1997) in hydrodynamics. The first hydrodynamic calculations with rotation in the quasi-spherical approximation have been undertaken by (Götz 1993).

Most aspects of early envelope growth, up to $\sim 10\,\mathrm{M}_\oplus$, can be understood on the basis of a simplified analytical model given by Stevenson (1982a) for a protoplanet with constant opacity κ_0, core-mass accretion-rate $\dot{M}_{\mathrm{core}}$, core-density ρ_{core}, inside the tidal radius R_T. The key properties of Stevenson's model come from the 'radiative zero solution' for spherical protoplanets with static, fully radiative envelopes, i.e., in hydrostatic and thermal equilibrium. We present here the solution relevant to the structure of an envelope in the gravitational potential of a constant mass, for zero external temperature and pressure and using a generalized opacity law of the form $\kappa = \kappa_0 P^a T^b$. The critical mass, defined as the largest mass to which a core can grow while forced to retain a static envelope, is then given by

$$M_{\rm core}^{\rm crit} = \left[\frac{3^3}{4^4} \left(\frac{\mathcal{R}}{\mu} \right)^4 \frac{1}{4\pi G} \frac{4-b}{1+a} \frac{3\kappa_0}{\pi\sigma} \left(\frac{4\pi}{3} \rho_{\rm core} \right)^{\frac{1}{3}} \frac{\dot{M}_{\rm core}}{\ln(R_T/r_{\rm core})} \right]^{\frac{3}{7}} , \tag{2}$$

and $M_{\rm core}^{\rm crit}/M_{\rm tot}^{\rm crit} = 3/4$; $\mathcal{R}$, G, σ denote the gas constant, the gravitational constant, and the Stefan-Boltzmann constant respectively. The critical mass depends on neither the midplane density $\varrho_{\rm Neb}$, nor on the temperature $T_{\rm Neb}$ of the nebula in which the core is embedded. The outer radius, R_T, enters only logarithmically. The strong dependence of the analytic solution on molecular weight, μ, led Stevenson (1984) to propose 'superganymedean puffballs' with atmospheres assumed to be enriched in heavy elements; such objects would have low critical masses, providing a way to form giant planets rapidly (see also Lissauer et al. 1995). Equation (2) permits a glimpse on the effect of the run of opacity via the power law exponents a and b. Except for the weak dependences discussed above, a proto-giant planet essentially has the same global properties for a given core wherever it is embedded in a nebula. Even the dependence on $\dot{M}_{\rm core}$ is relatively weak: Detailed radiative/convective envelope models show that a variation of a factor of 100 in $\dot{M}_{\rm core}$ leads only to a 2.6 variation in the critical core mass.

This similarity in the static structure of proto-giant envelopes yields similar dynamical behaviors characterized by pulsation-driven mass loss for solar composition nebula opacities (see Section V.B). However, other static solutions are found for protoplanets with *convective* outer envelopes, which occur for somewhat larger midplane densities than in minimum mass nebulae (Wuchterl 1993). These largely convective proto-giant planets have larger envelopes for a given core and a reduced critical core mass. Their properties can be illustrated by a simplified analytical solution for fully convective, adiabatic envelopes with constant first adiabatic exponent, Γ_1:

$$M_{\rm core}^{\rm crit} = \frac{1}{\sqrt{4\pi}} \frac{\sqrt{\Gamma_1 - \frac{4}{3}}}{(\Gamma_1 - 1)^2} \left(\frac{\Gamma_1}{G} \frac{\mathcal{R}}{\mu} \right)^{\frac{3}{2}} T_{\rm Neb}^{\frac{3}{2}} \rho_{\rm Neb}^{-\frac{1}{2}} , \tag{3}$$

and $M_{\rm core}^{\rm crit}/M_{\rm tot}^{\rm crit} = 2/3$. In this case, the critical mass depends on the nebula gas properties and therefore the location in the nebula, but it is independent of the core accretion rate. Of course, both the radiative zero and fully convective solutions are approximate because they only roughly estimate envelope gravity and all detailed calculations show radiative *and* convective regions in proto-giant planets. In Figure 3 the transition from 'radiative' to 'convective' protoplanets is shown by results from detailed static radiative/convective calculations for $\dot{M}_{\rm core} = 10^{-6}\,{\rm M}_\oplus\,{\rm yr}^{-1}$ (Wuchterl 1993). Nebula conditions are varied from low densities, resulting in radiative outer envelopes, to enhanced densities that result in largely convective proto-giant planets. The critical mass can be as low as $1\,M_\oplus$, and subcritical static envelopes can grow to $48\,M_\oplus$. Calculations with updated opacity and improved, mixing length convection (Wuchterl 1999), and the inclusion of rotational effects in the quasi-spherical approximation (Götz 1993) show a reduction of the critical core mass from 13 to $7\,M_\oplus$ for the 'radiative' proto-giant planets at the low nebula densities. The new, lower values are in better agreement with the new interior models.

The early phases of giant planet formation discussed above are dominated by the growth of the core. The envelopes adjust rapidly to the changing size and gravity of the core. As a result, the envelopes of proto-giant planets remain very close to static and in equilibrium below the critical mass (Mizuno 1980, Wuchterl 1993). This must change when the envelopes become more massive and cannot reequilibrate as fast as the cores grow. The nucleated instability was assumed to set in at the critical mass, originally as a hydrodynamic instability analogously to the Jeans instability. With the recognition that energy losses from the proto-giant planet envelopes control the further accretion of gas, it followed that quasi-hydrostatic contraction of the envelopes would play a key role.

V. DETAILED NUCLEATED INSTABILITY MODELS FOR THE GIANT PLANETS IN OUR SOLAR SYSTEM

Major progress has been made since PPIII by calculating the growth of the cores from planetesimal dynamics and the growth of the envelopes using hydrodynamics. We review these results below.

A. Quasi-Hydrostatic Models with Detailed Core-Accretion

Pollack et al. (1996) constructed models in which they simulated the concurrent accretion rates of both the gaseous and solid components of giant planets. Pollack et al. used an evolutionary model having three major components: a calculation of the three-body accretion rate of a single dominant-mass protoplanet surrounded by a large number of planetesimals, a calculation of the interaction of accreted planetesimals with the gaseous envelope of the growing giant protoplanet, and a calculation of the gas accretion rate using a sequence of quasi-hydrostatic models having a core/envelope structure. These three components of the calculation were updated every time step in a self-consistent fashion in which relevant information from one component was used in the other components.

The model of Pollack et al. (1996) is very detailed in many respects (core accretion rate, planetesimal dissolution in the envelope, treatment of energy loss via radiation and convection, equation of state), but it includes the following simplifying assumptions:

1. The planet is assumed to be spherically symmetric.

2. Hydrodynamic effects are not considered in the evolution of the envelope.

3. The opacity in the outer envelope is determined by a solar mixture of small grains in most of the simulations. Solar abundances are also used to calculate the opacity in deeper regions of the envelope where molecular opacities dominate.

4. The equation of state for the envelope is that for a solar mixture of elements.

5. During the entire period of growth of a giant planet, it is assumed to be the sole dominant mass in the region of its feeding zone, i.e., there are no competing embryos, and planetesimal sizes and random velocities remain small. A corollary of this assumption is that accretion can be described as a quasi-continuous process, as opposed to a discontinuous one involving the occasional accretion of a massive planetesimal.

6. Planetesimals are assumed to be well-mixed within the planet's feeding zone, which grows as the planet's mass increases, but planetesimals are not allowed to migrate into or out of the planet's feeding zone as a consequence of their own motion. Tidal interaction between the protoplanet and the disk, or migration of the protoplanet (Chapters by Artymowicz, Ward and Hahn and by Lin et al.), are not considered. It is not at all obvious that these various assumptions are valid, but no well-defined quantitatively justifiable alternative assumptions are available.

The parameters in the calculations of Pollack et al. (1996) were adjusted to fit the properties of giant planets in the Solar System and observations of disks around young stars. They judged the applicability of a given simulation to planets in our Solar System using two basic criteria. One criterion is provided by the time required to reach the runaway gas accretion phase. This time interval should be less than the lifetime of the *gas* component of the solar nebula, t_{sn}, for successful models of Jupiter and Saturn and greater than t_{sn} for successful models of Uranus and Neptune. Limited observations of accretion disks around young stars suggest that $t_{sn} \lesssim 10^7$ years, based on observations of the dust component (Chapters by Calvet et al., Natta et al., Lagrange et al.). The lifetime of the gas component is less well constrained observationally (Strom et al. 1993). See the chapter by Wadhwa and Russell and Podosek and Cassen (1994) for a review of nebula-lifetime estimates.

A second criterion is provided by the amount of high-Z mass accreted, M_Z. In the case of Jupiter and Saturn, M_Z at the end of a successful simulation should be comparable to, but somewhat smaller than, the current high-Z masses of these planets, since additional accretion of planetesimals occurred between the time they started runaway gas accretion and the time they contracted to their current dimensions and were able to gravitationally scatter planetesimals out of the Solar System. Updated values of the constraints on high-Z material in the jovian planets are discussed in Section II of this chapter.

In Pollack et al.'s (1996) models, there are three main phases to the accretion of Jupiter and Saturn. Phase 1 is characterized by rapidly varying rates of planetesimal and gas accretion. Throughout phase 1, dM_Z/dt exceeds the rate of gas accumulation, dM_{XY}/dt. Initially, there is a very large difference (many orders of magnitude) between these two rates. However, they become progressively more comparable as time advances. Over much of phase 1, dM_Z/dt increases steeply. After a maximum at 5×10^5 years, it declines sharply. Meanwhile, dM_{XY}/dt grows steadily from its extremely low initial value. The second phase of accretion is characterized by relatively time-invariant values of dM_Z/dt and dM_{XY}/dt, with $dM_{XY}/dt > dM_Z/dt$. Finally, phase 3 is defined by rapidly increasing rates of gas and planetesimal accretion, with dM_{XY}/dt exceeding dM_Z/dt by steadily increasing amounts. The accretion of Uranus and Neptune was terminated during phase 2, presumably as a result of the dissipation/dispersal of the gas in the protoplanetary disk.

The models of Pollack et al. (1996) imply that the crossover mass, at which the solids and gas components of the planet are equal in mass, depends almost exclusively on the surface mass density of solids and the distance from the Sun. The crossover time is a rapidly decreasing function of the initial surface mass density of solids. Surface mass density of $\approx 10\,\mathrm{g\,cm^{-2}}$ at Jupiter yields both a small enough condensables mass and rapid enough gas accretion to be consistent with observations for nominal values of other parameters. Good fits for Saturn and Uranus are obtained if surface density of solids drops off with distance from the Sun as r^{-2}. Constraints on the surface density are quite restrictive in the 'baseline' case, but a lower value is allowed if either the opacity of the outer envelope is low because grains sink or if planetesimal heating is reduced because accreted planetesimals dissolve well above the core and their residue does not sink to the core or planetesimal accretion stops during phase 2, e.g., as a result accretion by neighboring embryos (Pollack et al. 1996). Increasing the mean molecular weight of the envelope also increases the gas accretion rate (cf. eq. 4), but this parameter variation was not modeled by Pollack et al. (1996). The model results are relatively insensitive to moderately large changes in the gas density and temperature. Planetesimal size (which affects the velocity dispersion) is important in determining the duration of phase 1; for nominal parameters this has a small effect on the overall growth time for Jupiter, but the accretion time of Uranus is more profoundly affected by changes in planetesimal size.

B. Hydrodynamic Accretion Beyond the Critical Mass

Static and quasi-hydrostatic models discussed so far rely on the assumption that gas accretion from the nebula onto the core is very subsonic, and the inertia of the gas and dynamical effects such as dissipation of kinetic energy do not play a role. To check whether hydrostatic equilibrium is achieved and whether it holds, especially beyond the critical mass, hydrodynamical investigations are necessary. Two types of hydrodynamical investigations have been undertaken since PPIII: (1) linear adiabatic dynamical stability analysis of envelopes evolving quasi-hydrostatically (Tajima and Nakagawa 1997) and (2) nonlinear, convective radiation hydrodynamical calculations of core-envelope proto-giant planets (Wuchterl 1993, 1995b, 1996, 1997, 1999) that follow the evolution of a proto-giant planet without *a priori* assuming hydrostatic equilibrium and which *determine* whether envelopes are hydrostatic, pulsate or collapse, and at which rates mass flows onto the planet. Wuchterl's models solve the flow-equations for the envelope gas essentially only assuming that spherical symmetry holds. They determine the net gain and loss of mass from the equations of motion for the gas in spherical symmetry, while quasi-hydrostatic calculations add mass according to some prescription and then calculate the structure for the updated mass, yielding a new equilibrium. While the other assumptions made in the hydrodynamic calculations agree with those listed in the previous section for the quasi-hydrostatic models, there is a second important difference: the core accretion-rate is, for simplicity, assumed to be either constant or calculated according to the particle-in-a-box approximation (see e.g., Lissauer 1993).

The first hydrodynamical calculation of the nucleated instability (Wuchterl 1989, Wuchterl 1991a,b) started at the static critical mass and brought a surprise: Instead of collapsing, the proto-giant planet envelope begin to pulsate after a very short contraction phase (see Wuchterl 1990 for a simple discussion of the driving κ-mechanism). The pulsations of the inner protoplanetary envelope expanded the outer envelope and the outward traveling waves caused by the pulsations resulted in a mass loss from the envelope into the nebula. The process can be described as a pulsation-driven wind. After a large fraction of the envelope mass has been pushed back into the nebula, the dynamical activity fades and a new quasi-equilibrium state is found that resembles Uranus and Neptune in core and envelope mass (Wuchterl 1991a,b). The mass loss process occurs in a very similar way for nebula conditions at Jupiter to Neptune positions and for core mass accretion rates from 10^{-7} to $10^{-5}\,\mathrm{M}_{\oplus}\,\mathrm{yr}^{-1}$. Starting the hydrodynamics at low core mass rather than at the critical mass does not change the eventual mass loss (Wuchterl 1995b).

Pulsations and mass loss do not occur when 'no dust' zero metallicity opacities are used; the lack of dust makes conditions most favorable for energy loss from the envelope and therefore for accretion. It is interesting to note that even for zero metallicity opacities, the static critical core mass is between 1.5 and $3\,\mathrm{M}_\oplus$ for $\dot{M}_{\rm core} = 10^{-8}$ to $10^{-6}\,\mathrm{M}_\oplus\,\mathrm{yr}^{-1}$, respectively. Envelope accretion becomes independent of the core accretion at about $15\,\mathrm{M}_\oplus$, the quasi-hydrostatic assumptions holds until inflow velocities reach a Mach-number of 0.01 at about $50\,\mathrm{M}_\oplus$. At a total mass of about $100\,\mathrm{M}_\oplus$ the nebula gas influx approaches the Bondi accretion rate and at $300\,\mathrm{M}_\oplus$ the envelope collapses overall (cf. Wuchterl 1995b). This result shows that there must be an opacity-dependent transition from pulsation-driven winds to efficient gas accretion at the critical mass.

The main question concerning the hydrodynamics was then to ask for conditions that allow gas accretion, i.e., damp envelope pulsations for 'realistic' solar composition opacities that include dust. Wuchterl (1993) derived conditions for the breakdown of the radiative zero solution by determining nebulae conditions that would make the outer envelope of a 'radiative' critical mass proto-giant planet convectively unstable. The resulting criterion gives a minimum nebula density that is necessary for a convective outer envelope. Protoplanets that grow under nebula conditions above that density have larger envelopes for a given core and a reduced critical mass as described in Section IV.B. Since convection is of great importance in damping stellar pulsations of RR-Lyrae and δ-Cepheid stars at the cool, so-called 'red end' of the stellar instability strip, a similar behaviour may be expected in proto-giant planet envelopes. Wuchterl (1995b) calculated the growth of giant planets from low core masses hydrodynamically for a set of nebula conditions reaching from below the critical density to somewhat above. As the density was increased, the envelopes became increasingly more convective at the critical mass, but still showed the mass loss. At a nebula density of $10^{-9}\,\mathrm{g\,cm^{-3}}$, i.e., increased by a factor 6.7 relative to Mizuno's (1980) minimum reconstituted mass nebula value, the dynamical behaviour was different: the pulsations were damped and rapid accretion of gas set in and proceeded to $300\,\mathrm{M}_\oplus$. Apparently the spreading of convection in the outer envelope had damped the pulsations, thereby inhibiting the onset of a wind and leading to accretion. The critical core masses required for the formation of this class of proto-giant planets are significantly smaller than for the Uranus/Neptune-type (see Wuchterl 1993, 1995b).

Improved convective energy transfer and opacities: Most giant planet formation studies use zero entropy gradient convection, i.e., set the temperature gradient to the adiabatic value in convectively unstable layers of the envelope. That is done for simplicity but can be inaccurate, especially when the evolution is rapid and hydrodynamical waves are present (see Wuchterl 1991b). It was, therefore, important to develop a time-dependent theory of convection that can be solved together with the equations of radiation hydrodynamics. Such a time-dependent convection model (Kuhfuß 1987) has been reformulated for self-adaptive grid radiation hydrodynamics (Wuchterl 1995a) and applied to giant planet formation (Götz 1993, Wuchterl 1996, 1997). In a reformulation by Wuchterl and Feuchtinger (1998), it closely approximates standard mixing length theory in a static local limit and accurately describes the solar convection zone and RR-Lyrae lightcurves. In addition, updated molecular opacities (Alexander and Ferguson 1994) are used in a compilation by Götz (1993) to improve the accuracy of radiative transfer in the proto-giant planet envelopes. The effect of these improvements in energy transfer is that the core mass needed to initiate gas accretion to a few hundred Earth masses at various orbital radii is reduced to 8.30, 9.48 and 9.56 $M_\oplus$ at 0.052, 5.2 and 17.2 AU, respectively (see Fig. 4), even in a minimum mass nebula.

VI. FORMATION OF EXTRASOLAR PLANETS

More than a dozen planets have thus far been discovered to orbit main sequence stars other than the Sun; all of these objects are more massive than is Saturn, and most are more massive than Jupiter (Mayor and Queloz 1995; Chapter by Marcy et al. and references therein). The extrasolar planets currently known all orbit nearer to their stars than Jupiter does to the Sun (this is primarily an observational selection effect - high precision radial velocity surveys have not been in operation long enough to have observed a full orbit of more distant planets). Some of these planets orbit on highly eccentric paths, suggesting that after they formed they were subjected to close encounters with other giant planets (Weidenschilling and Marzari 1996, Lin and Ida 1997, Levison et al. 1998) or, in the case of the companion to 16 Cyg B, secular perturbations from the star 16 Cyg A (Holman et al. 1997). Some of the extrasolar planets are separated from their stars by less than one percent of the Jupiter-Sun distance. Guillot et al. (1996) showed that giant planets are stable over the main-sequence lifetime of a 1 $M_\odot$ star even if they are as close as 0.05 AU. Models involving migration caused by disk-planet interactions are favored by many researchers for the formation of these objects (e.g., Lin et al. 1996; Trilling et al. 1998; see also chapters by Ward and Hahn and by Lin et al.). However, simulations also show that it may be possible to form giant planets very close to stars, and we review these models in this section.

A. Hydrostatic Models for In-situ Formation

Bodenheimer et al. (1999) have modeled the formation and evolution of the planets recently discovered in orbit about the stars 51 Pegasi, ρ Corona Borealis and 47 Ursa Majoris, assuming that these planets formed in or near their current orbits. They used updated versions of the quasi-hydrostatic codes developed by Bodenheimer and Pollack (1986) and Pollack et al. (1996). The isolated protoplanet/no migration model of Pollack et al. requires high surface mass density of solids for giant planets to form close to stars within the observed lifetimes of protoplanetary disks. The primary cause of this restriction is that the larger Kepler shear near the star decreases the solid core's isolation mass unless the amount of solids is large; the increase in temperature closer to the star has only a very small effect (Mizuno 1980, Bodenheimer and Pollack 1986) and the higher density of gas acts in the opposite sense (Wuchterl 1996). The planet orbiting 2.1 AU from 47 UMa can form in $\sim 2\,\mathrm{Myr}$ for a surface density of condensed material $\sigma = 90\,\mathrm{g\,cm^{-2}}$, but requires $\sim 18\,\mathrm{Myr}$ for $\sigma = 50\,\mathrm{g\,cm^{-2}}$ (Bodenheimer et al. 1999). A value of $\sigma = 90\,\mathrm{g\,cm^{-2}}$ at 2.1 AU is well above that used by Pollack et al., but still well below that required for local axisymmetric gravitational instabilities (Toomre 1964) assuming a solar composition mix.

The surface mass density of solids required to form giant planets at 0.23 AU (ρ CrB) and 0.05 AU (51 Peg) is prohibitively large unless orbital decay of planetesimals is incorporated into the models. On the other hand, Ruzmaikina (1998) has given a model to provide the required amounts of both gas and solids. As no well-constrained method to quantify core growth close to stars is available, Bodenheimer et al. (1999) made the *ad hoc* assumption of a constant rate of solid body accretion for these inner planets. Model results for 51 Peg indicate that if the growth rate of the core is $1 \times 10^{-5}\ \mathrm{M_\oplus\ yr^{-1}}$, then the planet takes $\sim 4 \times 10^6$ years to form and has a final high-Z mass of $\sim 40\ \mathrm{M_\oplus}$. Using the same definition for the planetary radius and the same planetesimal accretion rate as used by Bodenheimer et al. (1999), Wuchterl obtained a critical core-mass of about $25\,\mathrm{M_\oplus}$. The two groups are currently attempting to resolve this discrepancy, an effort which will include calculations with identical opacities.

B. Hydrodynamical Models of Giant Planet Formation Near Stars

A major result of the hydrodynamical studies is that proto-giant planets may pulsate and develop pulsation-driven mass loss. Only if the pulsations are damped can gas-accretion produce Jupiter-mass envelopes. Since all extrasolar planets discovered so far have minimum masses $\gtrsim 0.5\,\mathrm{M_J}$, they probably require efficient gas accretion and should satisfy the convective outer envelope criterion (Wuchterl 1993). A glance at Wuchterl's (1993) Fig. 2 shows that proto-giant planets located somewhat inside of Mercury's orbit in the Hayashi et al. (1985) minimum mass nebula fulfill this condition. Convective radiation hydrodynamical calculations of core-envelope growth at 0.05 AU, for particle-in-a-box core mass accretion at nebula temperatures of 1250 and 600 K, show gas accretion beyond $300\,M_\oplus$ at core masses of $13.5\,M_\oplus$ and $7.5\,M_\oplus$, respectively (Wuchterl 1996, 1997).

It is interesting to apply the arguments based on the convection-controlled bifurcation in hydrodynamic accretional behavior to an ensemble of preplanetary nebula models, to simulate a variety of initial conditions for planet formation that might have been present around other stars. Wuchterl (1993) has shown that almost all nebula conditions, from a literature collection of nebula models, result in radiative outer envelopes at the critical mass. Nonlinear radiation hydrodynamical calculations with zero-entropy gradient convection show that Uranus/Neptune type giant planets are produced under such circumstances (Sect. V.B). Jupiter-mass planets should then be the exception. The first calculations with time-dependent mixing length convection, discussed in Section V.B, show gas accretion to beyond a Jupiter mass for a much wider range of nebula conditions. Apparently the improved description of convection (and the updated opacities) have shifted the instability-strip for pulsations and mass loss at the critical mass. Further calculations and a reanalysis of the conditions for efficient gas accretion for mixing length convection have to be undertaken before an updated expectation concerning the mass-distribution of extrasolar planets can be given. An important requirement for that is an extensive theoretical and observational study of plausible preplanetary nebulae.

VII. CONCLUSIONS

Jupiter and Saturn are composed primarily of hydrogen and helium, yet the heavy elements that they contain may hold the key to the problem of their formation. The density profiles of these planets derived from interior models, as well as the composition of their atmospheres, clearly indicate a significantly larger fraction of heavy elements than was present in the protosolar gas. Were the heavy elements the first to accrete, or did the enrichment occur at later stages? Depending on the scenario, Jupiter and Saturn might have received very different amounts of planetesimals, thereby providing a way to differentiate a very rapid formation (such as in the nebula instability mechanism) from a slower one (such as in the nucleated instability).

Interior and evolution models for Jupiter and Saturn tend to favor core masses that lie within the range of acceptable critical core masses predicted within the nucleated instability hypothesis. The models based on this hypothesis also explain why Uranus and Neptune are mostly core: either because (i) gas accretion is limited to $\sim 1\,M_\oplus$ by a hydrodynamic instability that operates under certain nebula conditions, low gas density being the dominant factor (Wuchterl 1993, 1995b); (ii) their cores grew more slowly than those of Jupiter and Saturn because orbital time scales are longer farther from the Sun, and thus they did not achieve sufficient mass to accrete large quantities of gas before the solar nebula gas was dispersed (Pollack et al. 1996); or (iii) the gas in the Uranus/Neptune region of the nebula was dispersed rapidly via photoevaporation, whereas gas remained in the Jupiter/Saturn region for a much longer period of time (Shu et al. 1993).

The nucleated instability hypothesis thus provides a viable model for the formation of the giant planets observed in our Solar System and beyond. Presently known extrasolar planets may have accreted *in situ* if their preplanetary nebula provided sufficient amounts of gas and solids. Alternatively, according to studies of disk-induced migration (Chapters by Ward and Hahn and by Lin et al.) and gravitational encounters with other planets, they could have formed elsewhere and moved into the present positions. In that case the orbits of most if not all planets known to be bound to main sequence stars *other than the Sun* suffered substantial orbital evolution.

The next few years will be dedicated to the development of a synoptic understanding of giant planet formation processes for a variety of preplanetary nebulae to work out predictive elements of the formation theories. These theories will be confronted by a representative observational census of giant planets orbiting neighboring stars.

Acknowledgments

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FIGURE CAPTIONS

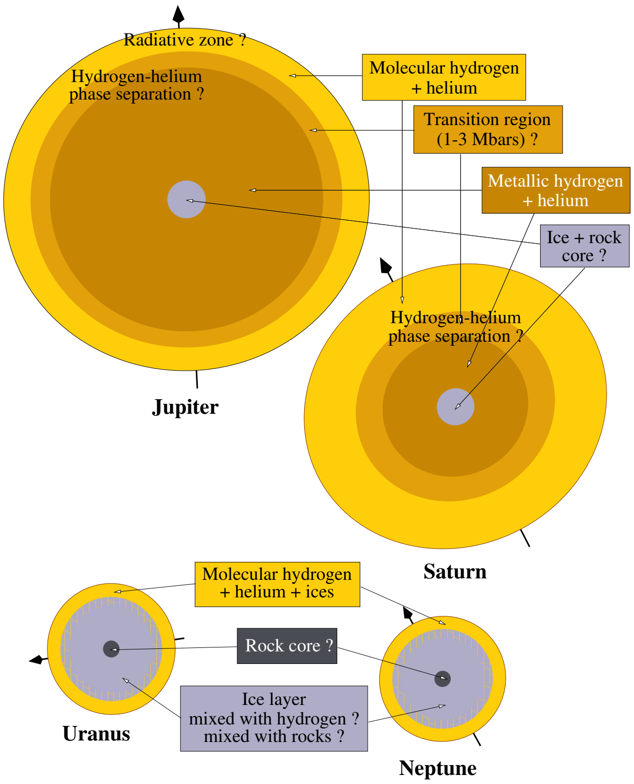

Figure 1. The interiors of Jupiter, Saturn, Uranus and Neptune, according to the conventional wisdom. The sizes and oblatenesses of the planets are represented to scale. Inside Jupiter and Saturn, hydrogen, which is in molecular form (H_2) at low pressures, is thought to become metallic in the 1 to 3 Mbar region. This transition could be abrupt or gradual. The equation of state is very uncertain for a substantial portion of the interiors of both planets. Uranus and Neptune, contain, in a relative sense, more heavy elements. There are indications that their interiors are partially mixed (see text).

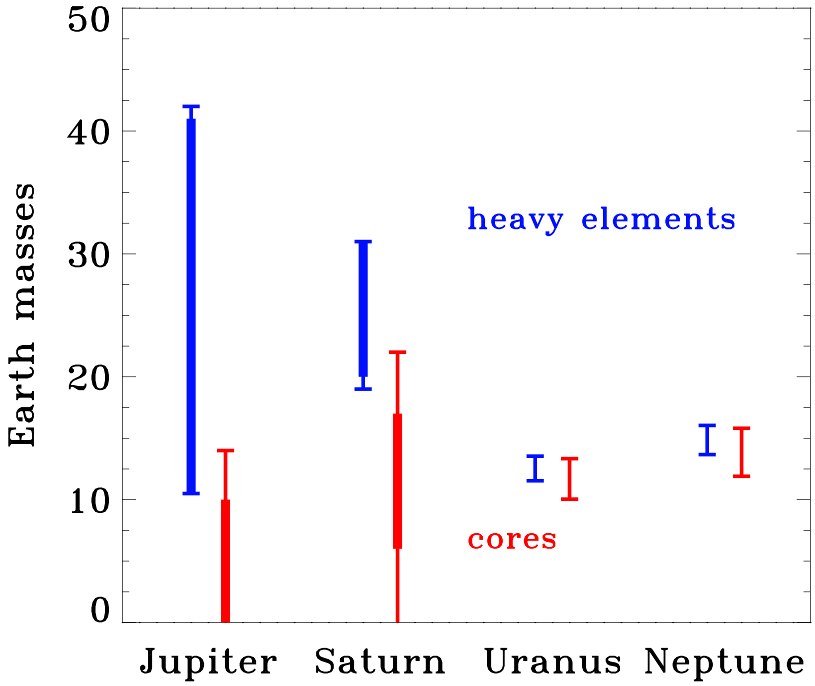

Figure 2. Limits on the abundances of heavy elements in the four jovian planets in our Solar System. For each planet, the point on the left represents the total amount of high-Z material, whereas the (lower) point on the right shows the amount of heavy elements segregated into the planet's core. For Jupiter and Saturn, the thick lines represent solutions with additional constraints obtained from evolution models. Note the high level of uncertainty, especially regarding the core masses of Jupiter and Saturn. Models of Jupiter with small cores (i.e., less than 2 $M_\oplus$) require significant enrichments in heavy elements (i.e., more than 20 $M_\oplus$).

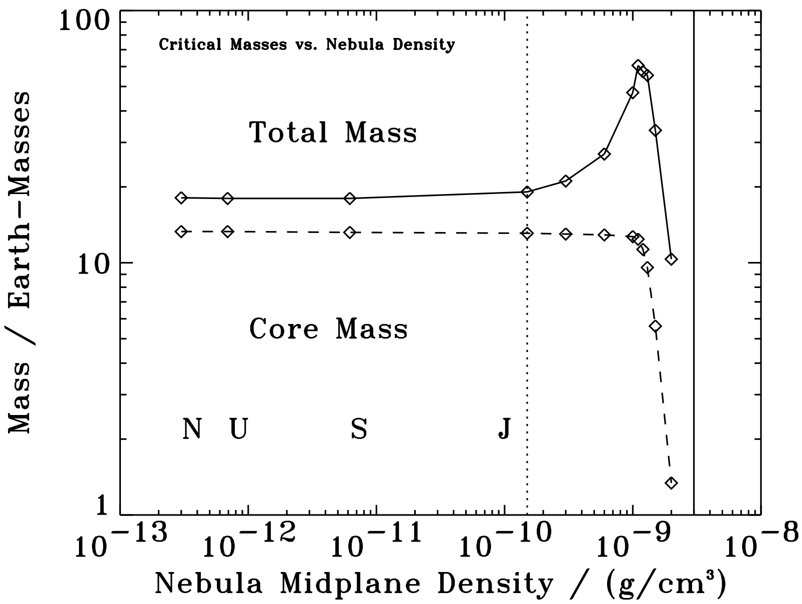

Figure 3. Critical masses of static protoplanets as a function of nebula midplane density. Critical total mass and core mass values are connected by a full and dashed line respectively. Observe the increased envelope masses and decreased core masses for the convective outer envelopes occurring at larger nebula densities. The conditions in the nebula correspond to Mizuno's minimum mass nebula (Mizuno 1980), densities at Neptune, Uranus, Saturn and Jupiter positions are labelled by N,U,S, and J, respectively. They illustrate the constancy of the critical mass in the case of radiative outer envelopes. Densities to the right of the dotted vertical line are arbitrarily enhanced relative to the minimum mass values, so that the outer envelopes become convective (see text). The full vertical line gives an estimate for the critical density of a Jupiter mass nebula fragment at Jupiter's position. The value plotted is the mean density of a condensation that is Jeans critical *and* fits into its Hill-Sphere.

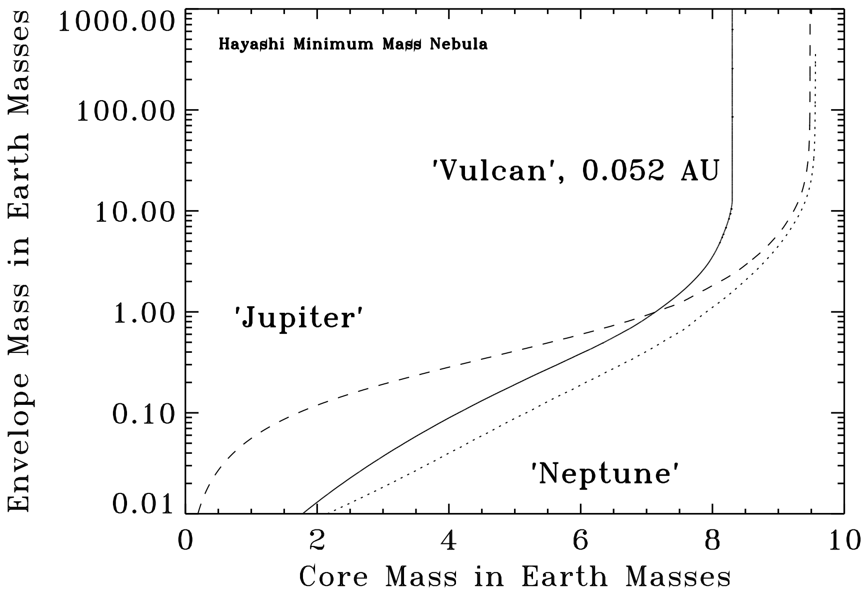

Figure 4. Evolution of proto-giant planets with mixing length convection. Envelope masses as obtained from hydrodynamic accretion calculations are plotted as functions of core mass for locations at 0.05 AU in the Hayashi et al. (1985) nebula (full) and for Mizuno's (1980) 'Jupiter' (dashed) and 'Neptune' (dotted) cases. Nebula temperatures and densities for the three cases are 'Vulcan': 1252 K, 5.3×10^{-6} g/cm^3, 'Jupiter': 97 K, 1.5×10^{-10} g/cm^3, and 'Neptune': 45 K, 3.0×10^{-13} g/cm^3. The core accretion rate is 10^{-6} $M_\oplus$/yr.

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