The Interiors of Giant Planets
Models and Outstanding Questions

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Abstract
We know that giant planets played a crucial role in the making of our Solar System. The discovery of giant planets orbiting other stars is a formidable opportunity to learn more about these objects, what is their composition, how various processes influence their structure and evolution, and most importantly how they form. Jupiter, Saturn, Uranus and Neptune can be studied in detail, mostly from close spacecraft flybys. We can infer that they are all enriched in heavy elements compared to the Sun, with the relative global enrichments increasing with distance to the Sun. We can also infer that they possess dense cores of varied masses. The intercomparison of presently characterised extrasolar giant planets show that they are also mainly made of hydrogen and helium, but that they either have significantly different amounts of heavy elements, or have had different orbital evolutions, or both. Hence, many questions remain and are to be answered for significant progresses on the origins of planets.
1 Introduction

Looking at a starry sky, it is quite vertiginous to think that we are at one of these special epochs in history: Just before, we only knew of the planets in our Solar System. Now, more than 150 giant planets are known to orbit solar-like stars. Our giant planets, Jupiter, Saturn, Uranus and Neptune are no longer the only ones that we can characterize. We now know of six extrasolar giant planets transiting in front of their stars, for which we can measure with a fair accuracy their mass and radius. We lie on the verge of a true revolution: With ground-based and future space-based transit search programs, we should soon be able to detect and characterize many tens, probably hundreds of planets orbiting their stars, with the hope of inferring their composition and hence the mechanisms responsible for the formation of planets.

It is a daunting task too, because we should expect that, like for the planets in our Solar System, a rich variety of giant planets is found, with different compositions, different histories and a number of new or unexpected physical mechanisms at work. We will have to classify observations, test theories, and be aware that although simplicity is appealing, it is not always what Nature has in store for us.

This review aims at providing a synthetic approach to the problems posed by “old” and “new” giant planets, in the Solar System and outside. It updates a previous review by Stevenson (1982), and expands on the review by Hubbard et al. (2002) by focusing on the mass-radius relations and compositions of giant planets. In Section 2, we see how to construct interior and evolution models of giant planets. Section 3 is devoted to our giant planets, what we can infer from observations, and the questions that remain. I then turn to the new field of extrasolar giant planets, focusing on the close-in, transiting “Pegasi planets” (also called “hot Jupiters”). The last section is an attempt to summarize some of the known facts concerning giant planets and provide a few expected milestones for future studies.

2 The calculation of interior models

2.1 A simple model

To tell our story, I will use a simple model, based on the following assumptions:

1. Giant planets are made of a fluid envelope and possibly a dense central core of about $\sim 15 M_\oplus$ (Earth masses);

2. The envelope is mostly made of hydrogen and helium and trace species (heavy elements); The core is made of an unknown combination of refract-
tory material (“rocks”) and more volatile species (“ices” including molecular species such as H2O water, CH4 methane and NH3 ammonia in the fluid state);

3. Contrary to solid planets, viscosity is negligible throughout;

4. In most cases, rotation and magnetic fields can be neglected;

5. Giant planets were formed from an extended, high-entropy, high-luminosity state.

These assumptions can only be justified \textit{a posteriori}: they are the result of our knowledge of observed giant planets and of inferences about the mechanisms that led to their formation. We will see how this simple model predicts the global properties of giant planets between 1/20 to 20 M\textsubscript{J} (about 15 and 6000 M\textsubscript{\oplus}), and how they compare with observations.

2.2 Basic equations

As a consequence of our assumptions, the structure and evolution of a giant planet is governed by the following hydrostatic, thermodynamic, mass conservation and energy conservation equations:

\[
\frac{\partial P}{\partial r} = -\rho g \\
\frac{\partial T}{\partial r} = \frac{\partial P}{\partial r} \frac{T}{P} \nabla T.
\]

\[
\frac{\partial m}{\partial r} = 4\pi r^2 \rho.
\]

\[
\frac{\partial L}{\partial r} = 4\pi r^2 \rho \left( \dot{\epsilon} - T \frac{\partial S}{\partial t} \right),
\]

where \(P\) is the pressure, \(\rho\) the density, and \(g = Gm/r^2\) the gravity (\(m\) is the mass, \(r\) the radius and \(G\) the gravitational constant). The temperature gradient \(\nabla T \equiv (d\ln T/d\ln P)\) depends on the process by which the internal heat is transported. \(L\) is the intrinsic luminosity, \(t\) the time, \(S\) the specific entropy (per unit mass), and \(\dot{\epsilon}\) accounts for the sources of energy due e.g. to radioactivity or more importantly nuclear reactions. Generally it is a good approximation to assume \(\dot{\epsilon} \sim 0\) for objects less massive than \(\sim 13\ M_\text{J}\), i.e. too cold to even burn deuterium (but we will see that in certain conditions this term may be useful, even for low mass planets).

The boundary condition at the center is trivial: \(r = 0; (m = 0, L = 0)\). The external boundary condition is more difficult to obtain because it depends on how energy is transported in the atmosphere. One possibility is to use the Eddington approximation, and to write (e.g. Chandrasekhar 1939): \(r = R; (T_0 = T_{\text{eff}}, P_0 = 2/3 g/\kappa)\), where \(T_{\text{eff}}\) is the effective temperature, and \(\kappa\) is the opacity in cm\(^2\)g\(^{-1}\). Note for example that in the case of Jupiter \(T_{\text{eff}} = 124\ K\),
\[ g = 2600 \text{ cm s}^{-2} \text{ and } \kappa \approx 5 \times 10^{-2} (P/1 \text{ bar}) \text{ cm}^2 \text{ g}^{-1}. \] This implies \( P_0 \approx 0.2 \text{ bar}, \) which is actually close to Jupiter’s tropopause, where \( T \approx 110 \text{ K}. \)

More generally, one has to use an atmospheric model relating the temperature and pressure at a given level to the radius \( R \), intrinsic luminosity \( L \) and incoming stellar luminosity \( L_{\star p} \): \( r = R; (T_0 = T_{00}(R, L, L_{\star p}), P_0 = P_{00}(R, L, L_{\star p})). \) \( P_0 \) is chosen to satisfy the condition that the corresponding optical depth at that level should be much larger than unity. If the stellar flux is absorbed mostly in a convective zone, then the problem can be simplified by using \( T_0(R, L, L_{\star p}) \approx T_0(R, L + L_{\star p}, 0) \) (e.g. Hubbard 1977). An example of such a model is described by Saumon et al. (1996) and Hubbard et al. (2002) and is used hereafter to model the planets in the low irradiation limit.

### 2.3 High pressure physics & equations of state

In terms of pressures and temperatures, the interiors of giant planets lie in a region for which accurate equations of state (EOS) are extremely difficult to calculate. This is because both molecules, atoms and ions can coexist, in a fluid that is partially degenerate (free electrons have energies that are determined both by quantum and thermal effects) and partially coupled (coulombian interactions between ions are not dominant but must be taken into account). The presence of many elements and their possible interactions further complicate matters. For lack of space, this section will mostly focus on hydrogen whose EOS has seen the most important developments in recent years. A phase diagram of hydrogen (fig. 1) illustrates some of the important phenomena that occur in giant planets.

The photospheres of giant planets are generally relatively cold (50 to 3000 K) and at low pressure (0.1 to 10 bar, or \( 10^4 \) to \( 10^6 \) Pa), so that hydrogen is in molecular form and the perfect gas conditions apply. As one goes deeper into the interior hydrogen and helium progressively become fluid. (The perfect gas relation tends to underestimate the pressure by 10% or more when the density becomes larger than about \( 0.02 \text{ g cm}^{-3} \) (\( P \sim 1 \text{ kbar in the case of Jupiter}).)

Characteristic interior pressures are considerably larger however: as implied by Eqs. 1 and 3: \( P_e \approx GM^2/R^4 \), of the order of 10-100 Mbar for Jupiter and Saturn. At these pressures and the corresponding densities, the Fermi temperature \( T_F \) is larger than \( 10^5 \text{ K}. \) This implies that electrons are degenerate. Figure 1 shows that inside Jupiter, Saturn, HD209458b, but also for giant planets in general for most of their history, the degeneracy parameter \( \theta = T/T_F \) is between 0.1 and 0.03. Therefore, the energy of electrons in the interior is expected to be only slightly larger than their non-relativistic, fully degenerate limit: \( u_e \geq 3/5 kT_F = 15.6 (\rho/\mu_e)^{2/3} \text{ eV} \), where \( k \) is Boltzmann’s constant, \( \mu_e \) is the number of electrons per nucleon and \( \rho \) is the density in \( \text{g cm}^{-3} \). For pure hydrogen, when the density reaches \( \sim 0.8 \text{ g cm}^{-3} \), the average energy of electrons becomes larger than hydrogen’s ionization potential, even at zero temperature: hydrogen pressure-ionizes and becomes metallic. This molecular to metallic transition occurs near Mbar pressures, but exactly how this happens remains unclear because of the complex interplay of thermal, coulombian, and
Figure 1: Phase diagram for hydrogen with the main phase transitions occurring in the fluid or gas phase. The temperature-pressure profiles for Jupiter, Saturn, Uranus, Neptune, and HD209458 b are shown. The dashed nearly vertical line near 1 Mbar is indicative of the molecular to metallic transition (here it represents the so-called plasma phase transition as calculated by Saumon et al. 2000). The region in which hydrogen is in solid phase (Datchi et al. 2000; Gregoryanz et al. 2003) is represented as a dashed area. The three phases (I,II,III) of solid hydrogen are shown (see Mao & Hemley 1994). Values of the degeneracy parameter $\theta$ are indicated as dotted lines to the upper right corner of the figure.
degeneracy effects (in particular, whether hydrogen metallizes into an atomic state \( \text{H}^+ \) — as suggested in Fig. 1 — or first metallizes in the molecular state \( \text{H}_2 \) remains to be clarified).

Recent laboratory measurements on fluid deuterium have been able to reach pressures above \( \gtrsim 1 \text{ Mbar} \), and provide new data in a region where the EOS remains most uncertain. Gas-guns experiments have been able to measure the reshock temperature (Holmes et al. 1995), near \( T \sim 5000 \text{ K}, P \sim 0.8 \text{ Mbar} \), and a rise in the conductivity of molecular hydrogen up to \( T \sim 3000 \text{ K}, P \sim 1.4 \text{ Mbar} \), a sign that metallicity may have been reached (Weir et al. 1996). The following few years have seen the development of laser-induced shock compression (Da Silva et al. 1997, Collins et al. 1998), pulsed-power shock compression (Kmdsson et al. 2002, 2004), and convergent shock wave experiments (Belov et al. 2002; Boriskov et al. 2003) in a high-pressure \( (P = 0.3 – 4 \text{ Mbar}) \) high-temperature \( (T \sim 6000 – 10^5 \text{ K}) \) regime. Unfortunately, experimental results along the principal Hugoniot of deuterium do not agree in this pressure range. Laser compression data give a maximum compression of \( \sim 6 \) while both the pulsed-power compression experiments and the convergent shock wave experiments find a value of \( \sim 4 \). Models that are partly calibrated with experimental data (Saumon, Chabrier & Van Horn 1995; Ross 1998; Saumon et al. 2000, Ross & Yang 2001) obtain a generally good agreement with the laser-compression data. However, the fact that independent models based on first principles (Militzer & Ceperley 2001; Desjarlais 2003; Bonev et al. 2004) yield low compressions strongly favors this solution.

The question of the existence of a first-order molecular to metallic transition of hydrogen (i.e. both molecular dissociation and ionisation occur simultaneously and discontinuously at the so-called plasma phase transition, or PPT) remains however. The critical line shown in fig. \( \text{I} \) corresponds to calculations by Saumon et al. (2000), but may be caused by artefacts in the free energy calculation. Recent Density Functional Theory (DFT) simulations by Bonev et al. (2004) indicate the possibility of a first order liquid-liquid transition but other path-integral calculations (Militzer & Ceperley 2001) do not. It is crucial to assess the existence of such a PPT because it would affect both convection and chemical composition in the giant planets.

A clear result from fig. \( \text{I} \) at least is that, as first shown by Hubbard (1968), the interiors of the hydrogen-helium giant planets are fluid, whatever their age (an isolated Jupiter should begin partial solidification only after at least \( \sim 10^3 \text{ Ga of evolution} \)). For Uranus and Neptune, the situation is actually more complex because at large pressures they are not expected to contain hydrogen, but numerical simulations show that ices in their interior should be fluid as well (Cavazzoni et al. 1999).

Models of the interiors of giant planets require thermodynamically consistent EOSs calculated over the entire domain of pressure and temperature spanned by the planets during their evolution. Elements other than hydrogen, most importantly helium, should be consistently included. Such a calculation is a daunting task, and the only recent attempt at such an astrophysical EOS for substellar objects is that by Saumon et al. (1995). Another set of EOSs reproducing either
the high- or low-compression results was calculated by Saumon & Guillot (2004) specifically for the calculation of present-day models of Jupiter and Saturn.

These EOSs have so far included other elements (including helium), only in a very approximative way, i.e. with EOSs for helium and heavy elements that are based on interpolations between somewhat ideal regimes, using an additive volume law, and neglecting the possibility of existence of phase separations (see Hubbard et al. 2002 and Guillot et al. 2004 for further discussions).

2.4 Heat transport

Giant planets possess hot interiors, implying that a relatively large amount of energy has to be transported from the deep regions of the planets to their surface. This can either be done by radiation, conduction, or, if these processes are not sufficient, by convection. Convection is generally ensured by the rapid rise of the opacity with increasing pressure and temperature. At pressures of a bar or more and relatively low temperatures (less than 1000 K), the three dominant sources of opacities are water, methane and collision-induced absorption by hydrogen molecules.

However, in the intermediate temperature range between $\sim 1200$ and $1500$ K, the Rosseland opacity due to the hydrogen and helium absorption behaves differently: the absorption at any given wavelength increases with density, but because the temperature also rises, the photons are emitted at shorter wavelengths, where the monochromatic absorption is smaller. As a consequence, the opacity can decrease. This was shown by Guillot et al. (1994) to potentially lead to the presence of a deep radiative zone in the interiors of Jupiter, Saturn and Uranus.

This problem must however be reanalyzed in the light of recent observations and analyses of brown dwarfs. Their spectra show unexpectedly wide sodium and potassium absorption lines (see Burrows, Marley & Sharp 2000), in spectral regions where hydrogen, helium, water, methane and ammonia are relatively transparent. It thus appears that the added contribution of these elements (if they are indeed present) would wipe out any radiative region at these levels (Guillot et al. 2004).

At temperatures above $1500 \sim 2000$ K two important sources of opacity appear: (i) the rising number of electrons greatly enhances the absorption of $\text{H}_2$ and $\text{H}^-$; (ii) TiO, a very strong absorber at visible wavelengths is freed by the vaporization of CaTiO$_3$. Again, the opacity rises rapidly which ensures a convective transport of the heat. Still deeper, conduction by free electrons becomes more efficient, but the densities are found not to be high enough for this process to be significant, except perhaps near the central core (see Hubbard 1968: Stevenson & Salpeter 1977).

However, because irradiated giant planets do develop a radiative zone, Rosseland opacity tables covering the proper range of temperatures and pressures are needed. A pure hydrogen-helium mixture table has been calculated by Lenzuni et al. (1991). Opacities for solar composition including grains are available from Alexander & Ferguson (1994), but they do not include alkali metals and
up-to-date data on water, methane and TiO absorption. Guillot (1999a) provides a grain-free, alkali-free table which is limited to low-temperature regimes. The calculations hereafter use opacities provided by F. Allard on the basis of calculations for brown dwarfs of solar composition, including grains and alkali metals (Allard et al. 2001).

2.5 The contraction and cooling histories of giant planets

The interiors of giant planets is expected to evolve with time from a high entropy, high \( \theta \) value, hot initial state to a low entropy, low \( \theta \), cold degenerate state. The essential physics behind can be derived from the well-known virial theorem and the energy conservation which link the planet’s internal energy \( E_i \), gravitational energy \( E_g \) and luminosity through:

\[
\xi E_i + E_g = 0, \quad \tag{5}
\]

\[
L = -\frac{\xi - 1}{\xi} \frac{dE_g}{dt}, \quad \tag{6}
\]

where \( \xi = \int_0^M 3(P/\rho)dm/\int_0^M \rho dm \approx <3P/\rho u> \) and \( u \) is the specific internal energy. For a diatomic perfect gas, \( \xi = 3.2 \); for fully-degenerate non-relativistic electrons, \( \xi = 2 \).

Thus, for a giant planet or brown dwarf beginning its life mostly as a perfect \( \text{H}_2 \) gas, two third of the energy gained by contraction is radiated away, one third being used to increase \( E_i \). The internal energy being proportional to the temperature, the effect is to heat up the planet. This represents the slightly counter-intuitive but well known effect that a star or giant planet initially heats up while radiating a significant luminosity.

Let us now move further in the evolution, when the contraction has proceeded to a point where the electrons have become degenerate. For simplicity, I will ignore Coulombian interactions and exchange terms, and assume that the internal energy can be written as \( E_i = E_{el} + E_{ion} \), and that furthermore \( E_{el} \gg E_{ion} \) (\( \theta \) is small). Because \( \xi \approx 2 \), we know that half of the gravitational potential energy is radiated away and half of it goes into internal energy. The problem is to decide how this energy is split into an electronic and an ionic part. The gravitational energy changes with some average value of the interior density as \( E_g \propto 1/R \propto \rho^{1/3} \). The energy of the degenerate electrons is essentially the Fermi energy: \( E_{el} \propto \rho^{2/3} \). Therefore, \( \dot{E}_{el} \approx 2(E_{el}/E_g)\dot{E}_g \). Using the virial theorem, this yields:

\[
\dot{E}_{el} \approx -\dot{E}_g \approx 2L \quad \tag{7}
\]

\[
L \approx -\dot{E}_{ion} \propto -\dot{T}. \quad \tag{8}
\]

The gravitational energy lost is entirely absorbed by the degenerate electrons, and the observed luminosity is due to the thermal cooling of the ions.

Several simplifications limit the applicability of this result (that would be valid in the white dwarf regime). In particular, the coulombian and exchange
terms in the EOS introduce negative contributions that cannot be neglected. However, the approach is useful to grasp how the evolution proceeds: in its very early stages, the planet is very compressible. It follows a standard Kelvin-Helmoltz contraction. When degeneracy sets in, the compressibility becomes much smaller (αT ≈ 0.1, where α is the coefficient of thermal expansion), and the planet gets its luminosity mostly from the thermal cooling of the ions. The luminosity can be written in terms of a modified Kelvin-Helmoltz formula:

$$L \approx \eta \frac{G M^2}{R \tau},$$

(9)

where τ is the age, and η is a factor that hides most of the complex physics. In the approximation that Coulombian and exchange terms can be neglected, η ≈ θ/(θ+1). The poor compressibility of giant planets in their mature evolution stages imply that η ≪ 1: the luminosity is not obtained from the entire gravitational potential, but from the much more limited reservoir constituted by the thermal internal energy. Equation 9 shows that to first order, log L ∝ −log τ: very little time is spent at high luminosity values. In other words, the problem is (in most cases) weakly sensitive to initial conditions.

Figure 2: Luminosity versus mass for giant planets after 4.5 Ga of evolution compared to measured values for our four giant planets (including the significant uncertainty on Uranus’ luminosity). The lines correspond to: H+He: a pure hydrogen-helium composition with a helium mass mixing ratio Y = 0.25; (a): a model with Y = 0.30 and a 15 M\(\oplus\) core; (b): the same model but with Y = 0.36.
Figure 2 shows calculated luminosities in the framework of our simple model. Compared to (9), calculated luminosities are consistent with $\eta \approx 0.01$ to 0.03. The lower luminosities obtained in the presence of a core and of more heavy elements are due to an earlier contraction, and quicker loss of the internal heat. As model (b) would be appropriate to explain Saturn’s radius (see next section), it can be seen that the planet emits more heat than predicted by homogeneous contraction models. The cases of Uranus and Neptune is more complex and cannot be directly compared with the models in fig. 2 which neglect the thermal heat content of the central core.

### 2.6 Mass-radius relation

![Diagram](image)

Figure 3: Radius versus mass for giant planets after 4.5 Ga of evolution compared to measured values for our four giant planets and four known extrasolar planets. As in fig. 2 the lines correspond to: H+He: a pure, $Y = 0.25$, hydrogen-helium composition ($Y=0.25$); (a): a model with $Y = 0.30$ and a 15 $M_{\oplus}$ core; (b): the same model but with $Y = 0.36$. An approximate mass-radius relation for zero-temperature water and olivine planets is shown as dashed and dash-dotted lines, respectively (Courtesy of W.B. Hubbard).

The relation between mass and radius has very fundamental astrophysical applications. Most importantly it allows one to infer the gross composition of an object from a measurement of its mass and radius. This is especially relevant in the context of the discovery of extrasolar planets with both radial velocimetry
and the transit method, as the two techniques yield relatively accurate determination of $M$ and $R$.

Figure 3 shows the mass-radius relation for isolated or nearly-isolated gaseous planets, based on our simple simple model and various assumption on their composition. The curves have a local maximum near $4 \, M_J$: at small masses, the compression is rather small so that the radius increases with mass. At large masses, degeneracy sets in and the radius decreases with mass.

This can be understood on the basis of polytropic models based on the assumption that $P = K \rho^{1+1/n}$, where $K$ and $n$ are constants. Because of degeneracy, a planet of large mass will tend to have $n \to 1.5$, while a planet a smaller mass will be less compressible ($n \to 0$). Indeed, it can be shown that in their inner 70 to 80% in radius isolated planets of 10, 1 and 0.1 $M_J$ have $n = 1.3, 1.0$ and 0.6, respectively. From polytropic equations (e.g. Chandrasekhar 1939):

$$R \propto K^{\frac{1}{n}} M^{\frac{3-n}{2}}.$$  \hspace{1cm} (10)

Assuming that $K$ is independent of mass, one gets $R \propto M^{0.16}, M^0$, and $M^{-0.18}$ for $M = 10, 1$ and 0.1 $M_J$, respectively, in relatively good agreement with fig. 3. (the small discrepancies are due to the fact that the intrinsic luminosity and hence $K$ depend on the mass considered).

Figure 3 shows already that the planets in our Solar System are not made of pure hydrogen and helium: their radii lie below that predicted for $Y = 0.25$ objects. Indeed, Jupiter, Saturn, and the two ice-giants Uranus and Neptune contain a growing proportion of heavy elements. The theoretical curves for olivine and ice planets predict even smaller radii however: even Uranus and Neptune contain 10 to 20% of their mass as hydrogen and helium.

The extrasolar planets detected so far (see table 3 hereafter) all lie above the pure hydrogen-helium curve. This is due to the fact that these planets have their evolutions dominated by the intense stellar irradiation they receive. Thermal effects are no longer negligible: Using the Eddington approximation, assuming $\kappa \propto P$ and a perfect gas relation in the atmosphere, one can show that $K \propto (M/R^2)^{-1/2n}$ and that therefore $R \propto M^{1/2n}$. With $n = 1$, one finds $R \propto M^{-1/2}$. Strongly irradiated hydrogen-helium planets of small masses are hence expected to have the largest radii which qualitatively explain the positions of the extrasolar planets in fig. 3. Note that this estimate implicitly assumes that $n$ is constant throughout the planet. The real situation is more complex because of the growth of a deep radiative region in most irradiated planets, and because of structural changes between the degenerate interior and the perfect gas atmosphere.

### 2.7 Rotation and the figures of planets

The mass and radius of a planet informs us on its global composition. Because planets are also rotating, one is allowed to obtain more information on their deep
interior structure. The hydrostatic equation becomes more complex however:

\[
\frac{\nabla P}{\rho} = \nabla \left( G \iiint \frac{\rho(r')}{|r - r'|} d^3r' \right) - \mathbf{\Omega} \times (\mathbf{\Omega} \times \mathbf{r}),
\]

(11)

where \( \mathbf{\Omega} \) is the rotation vector. The resolution of eq. (11) is a complex problem. It can however be somewhat simplified by assuming that \( |\mathbf{\Omega}| \equiv \omega \) is such that the centrifugal force can be derived from a potential. The hydrostatic equilibrium then writes \( \nabla P = \rho \nabla U \), and the figure of the rotating planet is then defined by the \( U = \text{cte} \) level surface.

One can show (e.g. Zharkov & Trubitsyn 1978) that the hydrostatic equation of a fluid planet can then be written in terms of the mean radius \( \bar{r} \) (the radius of a sphere containing the same volume as that enclosed by the considered equipotential surface):

\[
\frac{1}{\rho} \frac{\partial P}{\partial \bar{r}} = -\frac{GM}{\bar{r}^2} + \frac{2}{3} \omega^2 \bar{r} + \frac{GM}{\bar{R}^3} \bar{R} \phi_{\omega},
\]

(12)

where \( M \) and \( \bar{R} \) are the total mass and mean radius of the planet, and \( \phi_{\omega} \) is a slowly varying function of \( \bar{r} \). (In the case of Jupiter, \( \phi_{\omega} \) varies from about \( 2 \times 10^{-3} \) at the center to \( 4 \times 10^{-3} \) at the surface.) Equations (2-4) remain the same with the hypothesis that the level surfaces for the pressure, temperature, and luminosity are equipotentials. The significance of rotation is measured by the ratio of the centrifugal acceleration to the gravity:

\[
q = \frac{\omega^2 \bar{R}^3}{GM},
\]

(13)

The external gravitational potential of the planet is (assuming hydrostatic equilibrium):

\[
V_{\text{ext}}(r, \cos \theta) = \frac{GM}{r} \left[ 1 - \sum_{n=1}^{\infty} \left( \frac{a}{r} \right)^{2n} J_{2n}(P_{2n}(\cos \theta)) \right],
\]

(14)

where the coefficients \( J_{2n} \) are the planet’s gravitational moments, and the \( P_{2n} \) are Legendre polynomials. The \( J \)’s can be measured by a spacecraft coming close to the planet, preferably on a polar orbit. Together with the mass, this provides a constraint on the interior density profile (see Zharkov & Trubitsyn 1974):

\[
M = \iiint \rho(r, \theta) d^3\tau,
\]

\[
J_{2i} = -\frac{1}{MR_{\text{eq}}^{2i}} \iiint \rho(r, \theta) r^{2i} P_{2i}(\cos \theta) d^3\tau,
\]

where \( d\tau \) is a volume element and the integrals are performed over the entire volume of the planet.
Figure 4 shows how the different layers inside a planet contribute to the mass and the gravitational moments. The figure applies to Jupiter, but would remain very similar for other planets. Measured gravitational moments thus provide information on the external levels of a planet. It is only indirectly, through the constraints on the outer envelope that the presence of a central core can be inferred. As a consequence, it is impossible to determine this core’s state (liquid or solid), structure (differentiated, partially mixed with the envelope) and composition (rock, ice, helium...).

![Figure 4: Contribution of the level radii to the gravitational moments of Jupiter.](image)

For planets outside the solar system, although measuring their gravitational potential is utopic, their oblateness may be reachable with future space transit observations (Seager & Hui 2002). Since the oblateness $e$ is, to first order, proportional to $q$:

$$ e = \frac{R_{eq}}{R_{eq} - R_{pol}} \approx \left( \frac{3}{2} \Lambda_2 + \frac{1}{2} \right) q $$

(15)

(\text{where } \Lambda_2 = J_2/q \approx 0.1 \text{ to } 0.2), it may be possible to obtain their rotation rate, or with a rotation measured from another method, a first constraint on their interior structure.
3 Jupiter, Saturn, Uranus and Neptune

3.1 Main observational data

The mass of the giant planets can be obtained with great accuracy from the observation of the motions of their natural satellites: 317.834, 95.161, 14.538 and 17.148 times the mass of the Earth ($1\,M_{\oplus} = 5.97369 \times 10^{27}\,g$) for Jupiter, Saturn, Uranus and Neptune, respectively. The more precise determination of their gravity fields listed in table 1 have been obtained by the Pioneer and Voyager space missions.

|                | Jupiter          | Saturn          | Uranus          | Neptune         |
|----------------|------------------|-----------------|-----------------|-----------------|
| $M \times 10^{-29}\,[g]$ | 18.986112(15)$^a$ | 5.684640(30)$^b$ | 0.8683205(34)$^c$ | 1.0243542(31)$^d$ |
| $R_{eq} \times 10^{-9}\,[cm]$ | 7.1492(4)$^e$ | 6.0268(4)$^f$ | 2.559(4)$^g$ | 2.4766(15)$^h$ |
| $R_{pol} \times 10^{-9}\,[cm]$ | 6.6854(10)$^a$ | 5.4364(10)$^f$ | 2.4973(20)$^g$ | 2.4342(30)$^h$ |
| $R \times 10^{-9}\,[cm]$ | 6.9894(6)$^h$ | 5.8210(6)$^h$ | 2.5364(10)$^f$ | 2.4625(20)$^f$ |
| $\bar{\rho} \,[g\,cm^{-3}]$ | 1.3275(4) | 0.6880(2) | 1.2704(15) | 1.6377(40) |
| $J_2 \times 10^2$ | 1.4697(1)$^a$ | 1.6332(10)$^b$ | 0.35160(32)$^c$ | 0.3539(10)$^d$ |
| $J_4 \times 10^4$ | $-5.84(5)^a$ | $-9.19(40)^b$ | $-0.354(41)^c$ | $-0.28(22)^d$ |
| $J_6 \times 10^6$ | 0.31(20)$^a$ | 1.04(50)$^b$ | $\ldots$ | $\ldots$ |
| $P_2 \times 10^{-4}\,[s]$ | 3.57297(41)$^f$ | 3.83577(47)$^f$ | 6.206(4)$^k$ | 5.800(20)$^j$ |
| $q$ | 0.08923(5) | 0.15491(10) | 0.02951(5) | 0.02609(23) |
| $C/\bar{M}R_{eq}^2$ | 0.258 | 0.220 | 0.230 | 0.241 |

The numbers in parentheses are the uncertainty in the last digits of the given value. The value of the gravitational constant used to calculate the masses of Jupiter and Saturn is $G = 6.67259 \times 10^{-8}\,\text{dyn}\cdot\text{cm}^2\cdot\text{g}^{-1}$ (Cohen & Taylor, 1987).

$^a$ Campbell & Synott (1985)

$^b$ Campbell & Anderson (1989)

$^c$ Anderson et al. (1987)

$^d$ Tyler et al. (1989)

$^e$ Lindal et al. (1981)

$^f$ Lindal et al. (1985)

$^g$ Lindal (1992)

$^h$ From 4th order figure theory

$^i$ $(2R_{eq} + R_{pol})/3$ (Clairaut’s approximation)

$^j$ Davies et al. (1986)

$^k$ Warwick et al. (1986)

$^l$ Warwick et al. (1989)

Table 1 also indicates the radii obtained with the greatest accuracy by radio-occultation experiments. By convention, these radii and gravitational moments correspond to the 1 bar pressure level. The rotation periods are measured from the variations of the planets’ magnetic fields (system III) and are believed to be tied to the interior rotation. The giant planets are relatively fast rotators,
with periods of about 10 hours for Jupiter and Saturn, and about 17 hours for Uranus and Neptune. The fact that this fast rotation visibly affects the figure (shape) of these planets is seen by the significant difference between the polar and equatorial radii.

A first result obtained from the masses and radii (using the planets’ mean radii, as defined in section 2.7) indicated in Table 1 is the fact that these planets have low densities. These densities are similar, but considering that compression strongly increases with mass, one is led to a sub-classification between the hydrogen-helium giant planets Jupiter and Saturn, and the “ice giants” Uranus and Neptune.

The values of the axial moment of inertia $C$ have been calculated using the Radau-Darwin approximation (Zharkov & Trubitsyn 1978). Our four giant planets all have an axial moment of inertia substantially lower than the value for a sphere of uniform density, i.e. $2/5 M R^2$, indicating that they have dense central regions. This does not necessarily mean that they possess a core, but simply that the density profile departs significantly from a uniform value.

### Table 2: Energy balance as determined from Voyager IRIS data$^a$.

|                  | Jupiter  | Saturn  | Uranus  | Neptune |
|------------------|----------|---------|---------|----------|
| Absorbed power [$10^{23}$ erg.s$^{-1}$] | 50.14(248) | 11.14(50) | 0.526(37) | 0.204(19) |
| Emitted power [$10^{23}$ erg.s$^{-1}$] | 83.65(84) | 19.77(32) | 0.560(11) | 0.534(29) |
| Intrinsic power [$10^{23}$ erg.s$^{-1}$] | 33.5(26) | 8.63(60) | 0.034(38) | 0.330(35) |
| Intrinsic flux [erg.s$^{-1}$.cm$^{-2}$] | 5440.(430) | 2010.(140) | 42.(47) | 433.(46) |
| Bond albedo [] | 0.343(32) | 0.342(30) | 0.300(49) | 0.290(67) |
| Effective temperature [K] | 124.4(3) | 95.0(4) | 59.1(3) | 59.3(8) |
| 1-bar temperature$^b$ [K] | 165.(5) | 135.(5) | 76.(2) | 72.(2) |

$^a$ After Pearl & Conrath (1991)
$^b$ Lindal (1992)

Jupiter, Saturn and Neptune are observed to emit significantly more energy than they receive from the Sun (see Table 2). The case of Uranus is less clear. Its intrinsic heat flux $F_{\text{int}}$ is significantly smaller than that of the other giant planets. Detailed modeling of its atmosphere however indicate that $F_{\text{int}} \gtrsim 60$ erg cm$^{-2}$ s$^{-1}$ (Marley & McKay 1999). With this caveat, all four giant planets can be said to emit more energy than they receive from the Sun. Hubbard (1968) showed in the case of Jupiter that this can be explained simply by the progressive contraction and cooling of the planets.

It should be noted that the 1 bar temperatures listed in Table 2 are retrieved from radio-occultation measurements using a helium to hydrogen ratio which, at least in the case of Jupiter and Saturn, was shown to be incorrect. The new values of $Y$ are found to lead to increased temperatures by $\sim 5$ K in Jupiter and $\sim 10$ K in Saturn (see Guillot 1999a). However, the Galileo probe found a 1 bar temperature of 166 K (Seiff et al. 1998), and generally a good agreement with the Voyager radio-occultation profile with the wrong He/H$_2$ value.
3.2 Atmospheric composition

The most important components of the atmospheres of our giant planets are also among the most difficult to detect: $\text{H}_2$ and He have a zero dipolar moment. Also their rotational lines are either weak or broad. On the other hand, lines due to electronic transitions correspond to very high altitudes in the atmosphere, and bear little information on the structure of the deeper levels. The only robust result concerning the abundance of helium in a giant planet is by \textit{in situ} measurement by the Galileo probe in the atmosphere of Jupiter (von Zahn \textit{et al.} 1998). The helium mole fraction (\textit{i.e.} number of helium atoms over the total number of species in a given volume) is $q_{\text{He}} = 0.1359 \pm 0.0027$. The helium mass mixing ratio $Y$ (\textit{i.e.} mass of helium atoms over total mass) is constrained by its ratio over hydrogen, $X$: $Y/(X + Y) = 0.238 \pm 0.05$. This ratio is by coincidence that found in the Sun’s atmosphere, but because of helium sedimentation in the Sun’s radiative zone, it was larger in the protosolar nebula: $Y_{\text{proto}} = 0.275 \pm 0.01$ and $(X + Y)_{\text{proto}} \approx 0.98$. Less helium is therefore found in the atmosphere of Jupiter than inferred to be present when the planet formed.

Helium is also found to be depleted compared to the protosolar value in Saturn’s atmosphere. However, in this case the analysis is complicated by the fact that Voyager radio occultations apparently led to a wrong value. The current adopted value is now $Y = 0.18 - 0.25$ (Conrath & Gautier 2000), in agreement with values predicted by interior and evolution models (Guillot 1999a; Hubbard \textit{et al.} 1999). Finally, Uranus and Neptune are found to have near-protosolar helium mixing ratios, but with considerable uncertainty (Gautier & Owen 1989).

The abundance of “heavy elements”, \textit{i.e.} elements other than hydrogen and helium, bears crucial information for the understanding of the processes that led to the formation of these planets. Again, the most precise measurements are for Jupiter, thanks to the Galileo probe. As shown by fig. 5, most of the heavy elements are enriched by a factor 2 to 4 compared to the solar abundance (Niemann \textit{et al.} 1998; Owen \textit{et al.} 1999). One exception is neon, but an explanation is its capture by the falling helium droplets (Roustlon & Stevenson 1995). Another exception is water, but this molecule is affected by meteorological processes, and the probe was shown to have fallen into a dry region of Jupiter’s atmosphere. There are strong indications that its abundance is at least solar. Possible very high interior abundances ($\sim 10$ times the solar value) have also been suggested as a scenario to explain the delivery of heavy elements to the planet (Gautier \textit{et al.} 2001, Hersant \textit{et al.} 2004).

In the case of Saturn, both carbon in the form of methane and nitrogen as ammonia appear to be significantly enriched, but with large error bars (Atreya \textit{et al.} 2003). In Uranus and Neptune, methane is probably between 30 and 60 times the solar value (Gautier & Owen 1989; Hersant \textit{et al.} 2004).
Figure 5: Elemental abundances measured in the tropospheres of Jupiter (circles) and Saturn (squares) in units of their abundances in the protosolar nebula. The elemental abundances for Jupiter are derived from the in situ measurements of the Galileo probe (e.g. Mahaffy et al. 2000; Atreya et al. 2003). Note that the oxygen abundance is considered to be a minimum value due to meteorological effects (Roos-Serote et al. 2004). The abundances for Saturn are spectroscopic determination (Atreya et al. 2003 and references therein). The solar or protosolar abundances used as a reference are from Lodders (2003). The arrows show how abundances are affected by changing the reference protosolar abundances from those of Anders & Grevesse (1989) to those of Lodders (2003). The horizontal dotted lines indicate the locus of a uniform 2- and 4-times solar enrichment in all elements except helium and neon, respectively.
Figure 6: Schematic representation of the interiors of Jupiter and Saturn. The range of temperatures is estimated using homogeneous models and including a possible radiative zone indicated by the hashed regions. Helium mass mixing ratios $Y$ are indicated. The size of the central rock and ice cores of Jupiter and Saturn is very uncertain (see text). In the case of Saturn, the inhomogeneous region may extend down all the way to the core which would imply the formation of a helium core. [Adapted from Guillot 1999b].
3.3 Interior models: Jupiter and Saturn

As illustrated by fig. 6, the simplest interior models of Jupiter and Saturn matching all observational constraints assume the presence of three main layers: (i) an outer hydrogen-helium envelope, whose global composition is that of the deep atmosphere; (ii) an inner hydrogen-helium envelope, enriched in helium because the whole planet has to fit the H/He protosolar value; (iii) a central dense core. Because the planets are believed to be mostly convective, these regions are expected to be globally homogeneous. (Many interesting thermo-chemical transformations take place in the deep atmosphere, but they are of little concern to us).

A large part of the uncertainty in the models lies in the existence and location of an inhomogeneous region in which helium separates from hydrogen to form helium-rich droplets that fall deeper into the planet due to their larger density. Models have generally assumed this region to be relatively narrow, because helium was thought to be most insoluble in low-pressure metallic hydrogen (e.g. Stevenson 1982). However, DFT calculations have indicated that the critical temperature for helium demixing may rise with pressure (Pfaffenzeller et al. 1995), presumably in the regime where hydrogen is only partially ionized and bound states remain. This opens up the possibility that the inhomogeneous regions may be more extended. In particular, in the case of Saturn, Fortney & Hubbard (2003) have shown that explaining Saturn’s age may require that helium fall all the way to the core, thereby yielding the formation of a helium core (or of a helium shell around a rock or ice core).

With these caveats, the three-layer models can be used as a useful guidance to a necessarily hypothetical ensemble of allowed structures and compositions of Jupiter and Saturn. Figure 7 shows such an ensemble for Jupiter, based on calculations by Saumon & Guillot (2004). The calculations assume that only helium is inhomogeneous in the envelope (the abundance of heavy elements is supposed to be uniform across the molecular/metallic hydrogen transition). Many sources of uncertainties are included however; among them, the most significant are on the equations of state of hydrogen and helium, the uncertain values of $J_{4}$ and $J_{6}$, the presence of differential rotation deep inside the planet, the location of the helium-poor to helium-rich region, and the uncertain helium to hydrogen protosolar ratio.

These results show that Jupiter’s core is smaller than $\sim 10 M_{\oplus}$, and that its global composition is pretty much unknown (between 10 to $42 M_{\oplus}$ of heavy elements in total). The models indicate that Jupiter is enriched compared to the solar value, particularly with the new, low value of $Z_{\odot}$ (Lodders 2003) used in fig. 6. This enrichment could be compatible with a global uniform enrichment of all species near the atmospheric Galileo values. Alternatively, species like oxygen (as mostly water) may be significantly enriched.

Most of the constraints are derived from the values of the radius (or equivalently mass) and of $J_{2}$. The measurement of $J_{4}$ allows to further narrow the ensemble of possible models, and in some cases, to rule out EOS solutions (in particular those indicating relatively large core masses, between 10 and $20 M_{\oplus}$).
Figure 7: Constraints on Jupiter’s interior structure based on Saumon & Guillot (2004). The value of the core mass ($M_{\text{core}}$) is shown in function of the mass of heavy elements in the envelope ($M_Z$) for models matching all available observational constraints. The dashed region corresponds to models matching the laser compression experiments. The plain box corresponds to models matching the pulsed power and convergent shock compression experiments (see text). Grey lines indicate the values of $M_Z$ that imply uniform enrichments of the envelope in heavy elements by factors 2 to 8 times the solar value ($Z_\odot = 0.0149$), respectively.
As discussed in Guillot (1999a) and Saumon & Guillot (2004), most of the uncertainty in the solution arises because very different hydrogen EOSs are possible. The fact that more laboratory and numerical experiments seem to indicate relatively low-compressions for hydrogen at Mbar pressures points towards smaller core masses and a larger amount of heavy elements in the planet (plain box in fig. 7). However, this relies on uncertain temperature gradients, because the EOSs are based on laboratory data obtained at temperatures higher than those relevant to the planetary interiors.

Results slightly outside the boxes of fig. 7 are possible in the presence of a discontinuity of the abundance of heavy elements in the interior. Thus, Guillot (1999a) found slightly larger core masses (up to 12 M\(_\oplus\)) in the case of the Saumon-Chabrier EOS with a first order plasma-phase transition.

Figure 8: Same as fig. 7 in the case of Saturn. Note that smaller core masses could result either from allowing a variation of the abundance of heavy near the molecular/metallic transition (Guillot 1999a), or from the presence of a helium shell around the core (Fortney & Hubbard 2003).

In the case of Saturn (fig. 8), the solutions depend less on the hydrogen EOS because the Mbar pressure region is comparatively smaller. The total amount of heavy elements present in the planet can therefore be estimated with a better accuracy than for Jupiter. However, because Saturn’s metallic region is deeper into the planet, it mimics the effect that a central core would have on \( J_2 \). If we allow for variations in the abundance of heavy elements together with the helium discontinuity, then the core mass can become much smaller, and even solutions with no core can be found (Guillot 1999a). These solutions depend on the
hypothetic phase separation of an abundant species (e.g. water), and generally cause an energy problem because of the release of considerable gravitational energy. However, another possibility is through the formation of an almost pure helium shell around the central core, which could lower the core masses by up to $7\,M_{\oplus}$ (Fortney & Hubbard 2003; Hubbard, personal communication).

3.4 Interior models: Uranus and Neptune

![Figure 9: Schematic representation of the interiors of Uranus and Neptune. (Adapted from Guillot 1999b).]

Although the two planets are relatively similar, fig. 8 already shows that Neptune’s larger mean density compared to Uranus has to be due to a slightly different composition: either more heavy elements compared to hydrogen and helium, or a larger rock/ice ratio. The gravitational moments impose that the density profiles lie close to that of “ices” (a mixture initially composed of $H_2O$, $CH_4$ and $NH_3$, but which rapidly becomes an ionic fluid of uncertain chemical composition in the planetary interior), except in the outermost layers, which have a density closer to that of hydrogen and helium (Marley et al. 1995; Podolak et al. 2000). As illustrated in fig. 8, three-layer models of Uranus and Neptune consisting of a central “rocks” core (magnesium-silicate and iron material), an ice layer and a hydrogen-helium gas envelope have been calculated (Podolak et al. 1991; Hubbard et al. 1995).

The fact that models of Uranus assuming homogeneity of each layer and adiabatic temperature profiles fail in reproducing its gravitational moments seem to imply that substantial parts of the planetary interior are not homogeneously mixed (Podolak et al. 1995). This could explain the fact that Uranus’ 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. The existence of these non-
homogeneous, partially mixed regions are 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 $\sim 2.5$. 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% rocks, 60−70% ices, and 5−15% hydrogen and
helium (Podolak et al. 1991, 1995; Hubbard et al. 1995). Assuming both ices
and rocks are present in the envelope, an upper limit to the amount of hydrogen
and helium present is $\sim 4.2 \, M_\oplus$ for Uranus and $\sim 3.2 \, M_\oplus$ for Neptune (Podolak
et al. 2000). A lower limit of $\sim 0.5 \, M_\oplus$ for both planets can be inferred by
assuming that hydrogen and helium are only present in the outer envelope at
$P \lesssim 100$ kbar.

3.5 Are the interiors adiabatic?

As discussed, the near-adiabaticity of the interiors of the giant planets is a conse-
quence of the rapid rise of opacities with increasing pressure and temperatures.
Several exceptions are possible:
(i) In the “meteorological layer”, the temperature gradient could become
either subadiabatic (because of latent heat release and moist convection) or su-
peradiabatic (because of molecular weight gradients created by condensation
and precipitation). Locally, a depletion of an efficient radiative absorber (e.g.
water or methane) could imply that convection is suppressed, either because of
a lowered radiative gradient, or because sunlight can then be deposited to this
level. In Uranus and Neptune, a superadiabatic region at $P \sim 1 − 2$ bar is cor-
related with methane condensation (Lindal 1992, Guillot 1995). In Jupiter, the
Galileo probe measured a nearly-adiabatic profile, with a slight static stability
($N < 0.2 \, K \, km^{-1}$) down to 20 bars (Magalhães et al. 2002).
(ii) At the Plasma Phase Transition between molecular and metallic hy-
drogen, if it exists, with an entropy jump that could be of order $1 \, k_B$/baryon
(Stevenson & Salpeter 1977; Saumon et al. 1995).
(iii) In the hydrogen-helium phase separation region, where a slow droplet
formation may inhibit convection and yield a significant superadiabacity (Steven-
son & Salpeter 1977).
(iv) Near the core/envelope interface (whether it is abrupt or not) where an
inhibiting molecular weight gradient occurs and, in the case of Jupiter, conduction
might play a role.
(v) Throughout the planets, even though mixing-length arguments predict
that the superadiabacity is extremely small ($\sim 10^{-6}$ or less), rotation and mag-
netic fields may increase it, although probably by modest amounts (Stevenson
1982; See also discussion in Guillot et al. 2004).
3.6 What are the ages of our giant planets?

If we understand something of the formation of our Solar System and of other stars, our giant planets should have formed 4.55 Ga ago (e.g. Bodenheimer & Lin 2002). The model ages show significant deviations from that value, however.

In the case of Jupiter, the present radius and luminosity are obtained after 3.5 to 5.5 Ga of evolution, but most realistic EOSs predict ages above 4.5 Ga (Saumon & Guillot 2004). Several processes, among which core erosion, could lead to a reduction of that value (Guillot et al. 2004). For Saturn, homogeneous evolution models predict ages of order 2 Ga (Stevenson 1982; Saumon et al. 1992; Guillot et al. 1995). In both planets, the presence of a phase separation of helium is likely and would tend to lengthen the cooling.

The case of Uranus and Neptune is less clear-cut because of the uncertainties both on the properties of their atmospheres (in particular their evolution with time), and on the global specific heat of material inside. It appears however that both planets have luminosities that are too small. This could be due to a cold start (relatively low initial temperatures), a rapid loss of the internal heat, or a strong molecular weight gradient that prevents interior regions from cooling (Podolak et al. 1991; Hubbard et al. 1995).

3.7 Do some elements separate from hydrogen at high pressures?

Where?

Helium is strongly suspected of separating from hydrogen in Jupiter and Saturn because its lower than protosolar abundance in the atmosphere, and in Saturn because without this additional energy source, the planet would evolve to its present state in ~2 Ga. However, it has not been shown so far that a hydrogen helium mixture at Mbar pressures has a critical demixing temperature that is above that required in Jupiter and Saturn.

Helium demixing should occur in the metallic hydrogen region, but it is not clear that the critical temperature should decrease with pressure as for fully ionized plasmas (Stevenson 1982), or increase with pressure (Pfaffenzeller et al. 1995). The first scenario would imply the existence of a small inhomogeneous region near the molecular/metallic transition as illustrated in fig. 6. The second one would yield a more extended inhomogeneous region.

Evolution models including the two phase diagrams by Fortney & Hubbard (2003) show that in order to reconcile Saturn’s age with that of the Solar System and the atmospheric helium abundance derived by Conrath & Gautier (2000), sufficient energy $\Delta M_{\text{He}} g H$ is required. This implies maximizing $H$, the distance of sedimentation of helium droplets, and hence favors the Pfaffenzeller-type phase diagram and the formation of a helium core.

The question of a phase separation of other elements is still open. It is generally regarded as unlikely at least in Jupiter and Saturn because of their small abundances relative to hydrogen and the fact that the critical demixing temperature depends exponentially on that abundance.
3.8 How do the planetary interiors rotate?

Interior rotation is important because it affects the gravitational moments and their interpretation in terms of density profiles (Zharkov & Trubitsyn 1978). It is presently not known whether the observed atmospheric zonal flow patterns are tied to the planetary interiors or whether they are surface phenomena, with the interior rotating close to a solid body with the rate given by the magnetic field. Interior rotation affects more significantly gravitational moments of higher order. Using extrema set by solid rotation and by a model in which the zonal wind pattern is projected into a cylindrical rotation (Hubbard 1982), one can show that interior rotation introduces an uncertainty equivalent to the present error bar for $J_4$, of the order of the spread in interior models for $J_6$, and that becomes dominant for $J_8$ and above. Measurements of high order gravitational moments $J_8 - J_{14}$ should tell whether atmospheric zonal flow penetrate into the deep interior or whether the deep rotation is mainly solid (Hubbard 1999).

3.9 What can we tell of the giant planets’ cores? Are they primordial?

Confronted to diagrams such as figs. 6 and 9, there is the tendency to think that the giant planets cores as well defined, separate entities. It is not necessarily the case: first, as shown by fig. 4, solutions with a well-defined central core are equivalent to solutions with cores that have been diluted into the central half of the planet. Second, convection does not necessarily guarantee the presence of globally homogeneous regions, and can efficiently oppose the settling of species, as observed in thermohaline convection. Finally, the history of core formation, and in particular the epoch at which planetesimals were accreted and their sizes matter (e.g. Stevenson 1985).

Once formed, the cores of the giant planets are difficult to erode, as this demands both that heavy elements are (at least partially) soluble in the hydrogen helium envelope, and that enough energy is present to overcome the molecular weight barrier that is created (Stevenson 1982). However, in the case of Jupiter at least, the second condition may not be that difficult to obtain, as only 10% of the energy in the first convective cell (in the sense of the mixing length approach) needs to be used to dredge up about 20 $M_\oplus$ of core material (Guillot et al. 2004). Evaluating whether the first condition is satisfied would require knowing the core’s composition and its state, but one can nevertheless note that the initially high central temperatures ($\sim 30,000$ K) favor solubility. Such an efficient erosion would not occur in Saturn (and much less so in Uranus and Neptune) because of its smaller total mass.

3.10 Do we understand the planets’ global compositions?

This may be the hardest question because it requires tying all the different aspects of planet formation to the observations of the atmospheres of the giant planets and the constraints on their interior structures. So far, most of the focus
has been on explaining the presence of a central core of $\sim 10 \, M_\oplus$ in Jupiter, Saturn, Uranus and Neptune. The new interior data suggest that Jupiter’s core is probably smaller, and that Saturn’s may be larger. More importantly, the envelopes of all planets appear to be enriched in heavy elements, and this has to be explained as well.

The possibility that Jupiter could have been formed by a direct gravitational instability (e.g. Boss 2000) may be appealing in view of its small inferred core. However, the enrichment of its envelope in heavy elements is difficult to explain within that scenario, given the low accretion rate of a fully-formed Jupiter (Guillot & Gladman 2000).

The leading scenario therefore remains the standard “core accretion” scenario (Pollack et al. 1996), with the addition that Jupiter, Saturn, Uranus and Neptune were closer together (5-20 AU) just after their formation (Levison & Morbidelli 2003). While this scenario require core masses $\gtrsim 10 \, M_\oplus$, the possibility of an erosion of Jupiter’s core is appealing because it would both explain the difference in size with Saturn and an enrichment of its envelope. Although more limited, a small $\sim 2 \, M_\oplus$ erosion of Saturn’s core could provide part of the enrichment of the envelope (Guillot et al. 2004).

The fact that Jupiter’s atmosphere is also enriched in noble gases, in particular Ar which condenses at very low temperatures ($\sim 30 \, K$) is still a puzzle. Presently invoked explanations include a clathration of noble gases in ices (Gautier et al. 2001; Hersant et al. 2004), and the delivery of planetesimals formed at very low temperatures (Owen et al. 1999).

Finally, the large enrichments in C and possibly N of the atmospheres of Uranus and Neptune probably indicate that a significant mass of planetesimals ($\gtrsim 0.1 \, M_\oplus$) impacted the planets after they had captured most of their present hydrogen-helium envelopes. Along with the other problems related to this section, this requires quantitative work.

4 Extrasolar planets

4.1 Observables

More than 145 extrasolar planets have been discovered to date (see J. Schneider’s Extrasolar Planets Encyclopedia on [http://www.obspm.fr/planets]), but only those for which a determination of both the planetary mass and radius are useful for the purposes of this review. This can only be done for planets which transit in front of their star, which, by probabilistic arguments limits us to planets that orbit close to their star. I will therefore only be concerned with “Pegasi planets”, giant planets similar to 51 Peg b and HD209458b (both in the constellation Pegasus), with semi-major axes smaller than 0.1 AU.

Six transiting Pegasi planets have been discovered so far. Their main characteristics are listed in Table 3. The first one, HD209458b (Charbonneau et al. 2000; Henry et al. 2000), has been shown to possess sodium in its atmosphere (Charbonneau et al. 2002) and to have an extended, evaporating atmosphere
Table 3: Systems with transiting Pegasi planets discovered so far

| System      | Age [Ga] | [Fe/H] | a [AU]  | $T_{\text{eq}}$ [K] | $M_p/M_J$ | $R_p/10^{10}$ cm |
|-------------|----------|--------|---------|----------------------|-----------|-------------------|
| HD209458$^a$ | 4 – 7    | 0.00(2) | 0.0462(20) | 1460(120) | 0.69(2)  | 1.02(9)        |
| OGLE-56$^b$  | 2 – 4    | 0.0(3)  | 0.0225(4)  | 1990(140) | 1.45(23) | 0.88(11)       |
| OGLE-113$^c$ | ?        | 0.14(14) | 0.0228(6)  | 1330(80)  | 0.765(25) | 0.77(±2)       |
| OGLE-132$^d$ | ?        | 0.04(18) | 0.0307(5)  | 2110(150) | 1.19(13) | 0.81(6)        |
| OGLE-111$^e$ | ?        | 0.12(28) | 0.0470(10) | 1040(160) | 0.53(11) | 0.71(±9)       |
| TrES-1$^f$   | ?        | 0.00(4)  | 0.0393(11) | 1180(140) | 0.75(7)  | 0.77(4)        |

$^a$ Equilibrium temperature calculated on the basis of a zero planetary albedo

$^b$ Torres et al. (2004), Sasselov (2003), Konacki et al. (2003)

$^c$ Bouchy et al. (2004), Konacki et al. (2004)

$^d$ Moutou et al. (2004)

$^e$ Pont et al. (2004)

$^f$ Laughlin et al. (2004), Sozzetti et al. (2004), Alonso et al. (2004)

(Vidal-Madjar et al. 2003, 2004). Four others have been discovered by the photometric OGLE survey and subsequent radial velocity measurements (Konacki et al. 2003, 2004; Bouchy et al. 2004, Pont et al. 2004). One is a result of the TrES network survey (Alonso et al. 2004). Present photometric surveys have a strong detection bias towards very short periods. Associated to a probability of transiting that is inversely proportional to the orbital distance, this shows that Table 3 represents only a tiny fraction of planets which may have a low probability of existence.

A crucial parameter for the evolution models is the equilibrium temperature $T_{\text{eq}} = T_\star \sqrt{R_\star/2a}$ (assuming a zero albedo, i.e. that all incoming stellar light is absorbed by the planetary atmosphere). With values of $T_{\text{eq}}$ between $\sim 1000$ and $2000$ K, the present sample of transiting planets is already quite rich.

### 4.2 Observed vs. calculated radii of “Pegasi planets”

Contrary to the giant planets in our Solar System, Pegasi planets are subject to an irradiation from their central star that is so intense that the absorbed stellar energy flux is about $\sim 10^4$ times larger than their intrinsic flux (estimated from $T_{\text{eq}}$, or calculated directly). The atmosphere is thus prevented from cooling, with the consequence that a radiative zone develops and governs the cooling and contraction of the interior (Guillot et al. 1996). Typically, for a planet like HD209458b, this radiative zone extends to kbar levels, $T \sim 4000$ K, and is located in the outer 5% in radius (0.3% in mass) (Guillot & Showman 2002).

Problems in the modeling of the evolution of Pegasi planets arise mostly because of the uncertain outer boundary condition. The intense stellar flux implies that the atmospheric temperature profile is extremely dependant upon the opacity sources considered. Depending on the chosen composition, the opacity data used, the assumed presence of clouds, the geometry considered, resulting
temperatures in the deep atmosphere can differ by up to \( \sim 600 \text{ K} \) (Seager & Sasselov 1998, 2000; Goukenleuque et al. 2000; Barman et al. 2001; Sudarsky et al. 2003; Iro et al. 2004). Because of this problem, and in the framework of our simple model, the following discussion will be based on an outer boundary condition at 1 bar and a fixed temperature \( T_1 = 1500 \) or 2000 K\(^1\).

Another related problem is the presence of the radiative zone. Again, the composition is unknown and the opacity data are uncertain in this relatively high temperature \((T \sim 1500 - 3000 \text{ K})\) and high pressure (up to \( \sim 1 \text{ kbar} \)) regime. Results from our models are based on opacities from Allard et al. (2001). Other calculations using e.g. the widely used Alexander & Ferguson (1994) opacities do yield only a slightly faster cooling even though the Rosseland opacities are lower by a factor \( \sim 3 \) in this regime.

The resulting mass-radius relations are shown in fig. 10 for \( T_1 = 1500 \) and 2000 K, and compared to the observations for the planets listed in Table 3. For each case, an upper limit on the radius is obtained from a pure hydrogen-helium composition with \( Y = 0.25 \). An ad hoc lower limit comes from a model with a 15 M\(_\oplus\) central core, and a \( Y = 0.30 \) envelope. In both case, the opacity table is unchanged.

Figure 10 shows that within uncertainties, the measurements for 4 planets out of 6 can be explained in the framework of our simple model. However, two cases stand out: OGLE-TR-132b appears too small for its age implying that it may contain significant amounts of heavy elements in a core or in its deep interior. The case of HD209458b is more problematic: the constraints on its age, mass, an deep atmospheric temperature that should be \( \sim 1500 - 2000 \text{ K} \) yield radii that are about 10 to 20\% smaller than measured (Bodenheimer et al. 2001, 2003; Guillot & Showman 2002; Baraffe et al. 2003). The fact that the measured radius corresponds to a low-pressure (\( \sim \text{nbar} \)) level while the calculated radius corresponds to a level near 1 bar is not negligible (Burrows et al. 2004) but too small to account for the difference. This is problematic because while it is easy to invoke the presence of a massive core to explain the small size of a planet, a large size such as that of HD209458b may require an additional energy source.

Bodenheimer et al. (2001) proposed that this large radius may be due to a small forced eccentricity \((e \sim 0.03)\) of HD209458b, and subsequent tidal dissipation in the planet interior. In this case, \( \dot{e} > 0 \) in the energy conservation equation (1). Because of the relatively limited amount of energy available in the (non-circular) orbit and the presumably rapid dissipation (due to a tidal \( Q \) that is presumably similar to that of Jupiter, i.e. \( Q \sim 10^3 - 10^6 \)), this requires the presence of an unseen eccentric companion. The search for this companion and a possible non-zero eccentricity of HD209458b is ongoing (Bodenheimer et al. 2003).

A natural possibility may be the stellar flux itself, since transporting to deep

\(^1\)Technically, in order to obtain high entropy initial conditions I use \( T_1 \approx T_{eq}(1+L/L_{eq})^{1/4} \), but the precise form does not matter as long as \( L \ll L_{eq} \), or equivalently \(-T_1 dS_1/dt \ll -T_{int} dS_{int}/dt \) where \( S_{int} \) is the characteristic interior entropy. The 1.equality between \( T_1 \) and \( T_{eq} \) is only a very rough estimate guided by present works on atmospheric models of heavily irradiated planets.
Figure 10: Mass-radius relation of strongly irradiated planets with ages of 1 Ga (upper panels) and 5 Ga (lower panels), and 1-bar temperatures equal to 2000 (left panels) and 1500 K (right panels), respectively. The hashed areas have upper and lower envelopes defined by \((Y = 0.25, M_{\text{core}} = 0)\) and \((Y = 0.30, M_{\text{core}} = 15 M_\oplus)\), respectively. Dotted symbols with error bars indicate known objects, plotted as a function of their estimated 1-bar temperatures and ages. Planets whose age is uncertain appear in both upper and lower panels. Results for non-irradiated planets (dotted lines) are shown for an easier comparison.
levels (~ 100 bars or more) only a small fraction of order 0.1% to 1% of the incoming flux would yield a radius that is in agreement with the observations. On this basis, Showman & Guillot (2002) proposed that kinetic energy generated in the atmosphere due to the strong asymmetry in stellar insolation may be transported to deep levels and dissipated there, possibly due to a small asynchronous rotation and its dissipation by stellar tides. Another possibility evoked by the authors that km/s atmospheric winds may maintain the atmosphere into a shear-unstable, quasi-adiabatic state, which would force temperatures in excess of 3000 K at levels between 10 and a few tens of bars.

It is puzzling that all other recently announced transiting planets do not require an additional energy source to explain their size: this is seen in fig. 10, which shows that all planets except HD209458b are consistent with the evolutionary tracks.

Is there a consistent scenario explaining all the observations? One possibility is that, as proposed by Bodenheimer et al. (2001), HD209458b indeed has an eccentric companion. A second possibility is that their orbital histories have been very different. Finally, the planets may well have different compositions.

4.3 How do tides and orbital evolution affect the contraction and cooling of Pegasi planets?

The small orbital eccentricities of Pegasi planets compared to more distant extrasolar planets tells us that tides raised by the star on the planet have probably played an important role in circularizing their orbits, with a timescale estimated at ~ 1 Ga for a planet at 0.05 AU (Rasio et al. 1996; Marcy et al. 1997). Synchronisation is expected to occur in only Ma timescales (Guillot et al. 1996), maybe much less (Lubow et al. 1997). The tides raised by the planet on the star also tend to spin up the star which leads to a decay of the planetary orbit. It is interesting to note that, with periods of only ~ 1 day the three OGLE planets lie close to the orbital stability threshold (Rasio et al. 1996), or would be predicted to fall into the star in Ga timescales or less (Witte & Savonije 2002; Pätzold & Rauer 2002).

The energies available from circularisation and synchronisation can be usefully compared to the gravitational energy of the planet (e.g. Bodenheimer et al. 2001; Showman & Guillot 2002):

\[
E_{\text{circ}} = \frac{e^2 G M_a M}{a} = 3.6 \times 10^{42} \left( \frac{e}{0.1} \right)^2 \left( \frac{M_a}{M_{\odot}} \right) \left( \frac{M}{M_J} \right) \left( \frac{a}{10R_{\odot}} \right) \text{erg},
\]

\[
E_{\text{sync}} = \frac{1}{2} k^2 M R^2 \Delta \omega^2 = 2.4 \times 10^{41} \left( \frac{k^2}{0.25} \right) \left( \frac{M}{M_J} \right) \left( \frac{R}{10^{10} \text{cm}} \right)^2 \left( \frac{\Delta \omega}{10^{-4} \text{s}^{-1}} \right)^2 \text{erg},
\]

\[
E_{\text{grav}} = \frac{\delta GM^2}{R} = 2.4 \times 10^{42} \left( \frac{\delta}{0.1} \right) \left( \frac{M}{M_J} \right)^2 \left( \frac{R}{10^{10} \text{cm}} \right) \text{erg},
\]

where \( e \) is the initial eccentricity, \( a \) the planet’s orbital distance, \( M \) is its mass, \( R \) is its radius, \( k \) is the dimensionless radius of gyration, \( \Delta \omega \) is the change in
the planet’s spin before and after synchronisation, and $\delta$ is approximatively the change in the planet’s radius (neglecting any structural changes in the calculation of $E_{\text{sync}}$ and $E_{\text{grav}}$). $E_{\text{grav}}$ is the gravitational energy lost by the planet when its radius decreases by a factor $\sim \delta$, or alternatively the minimum energy required to expand its radius by the same factor.

The fact that the three energy sources are comparable imply that very early in the evolution, circularisation and synchronisation may have played a role, perhaps inducing mass loss (Gu et al. 2004). Once a planet has contracted to a degenerate, low $\theta$ state, the gravitational energy becomes large, and circularisation and synchronisation only have a limited role to play. However, two reservoirs can be invoked: the orbital energy of a massive eccentric planet that would force a non-zero eccentricity of the inner one (Bodenheimer et al. 2001) and the absorbed stellar luminosity in its ability to create kinetic energy in the atmosphere (Showman & Guillot 2002).

A major uncertainty related to these processes and how they affect the planetary structure is to know how and where energy is dissipated. Lubow et al. (1997) proposed that a resonant tidal torque is exerted at the outer boundary of the inner convection zone, and that dissipation occurs through the damping of gravity waves propagating in the outer stable radiative region. Contrary to Jupiter, this may be an efficient process because Pegasi planets have a radiative region that extends to great depths. Another possibility is through the excitation of inertial waves in the convective region, a process that would occur also in our giant planets (Ogilvie & Lin 2004). The location of the dissipation is not clear, however. If it occurs in the atmosphere, the effect of tides on the evolution will be limited, whereas they will have a maximum impact if they occur deep into the radiative zone (Guillot & Showman 2002).

If dissipation cannot reach into the deep interior, the planets will not inflate significantly when they migrate to their present location. This would imply that HD209458b must have migrated from several AU to its present location in less than $\sim 10$ Ma (Burrows et al. 2000b). In this framework, one could invoke a late migration of the OGLE planets (in particular OGLE-TR-132b) to explain their relatively small radius compared to HD209458b.

### 4.4 How does the composition affect the structure and evolution?

It is generally believed that giant planets of the mass of Jupiter should have near solar composition and relatively small core masses. However, it may not be the case: first, Jupiter is in fact relatively significantly enriched in heavy elements. Second, while Jupiter is very efficient at ejecting planetesimals from the Solar System, Pegasi planets are unable to do so because the local orbital speed $(GM_*/a)^{1/2} \sim 150 \text{ km s}^{-1}$ is much larger than the planet’s escape velocity $(2GM/R)^{1/2} \sim 50 \text{ km s}^{-1}$ (Guillot & Gladman 2000). Furthermore, most planetesimals on low $e$ orbits close to the planet would end up impacting the planet, not the star (A. Morbidelli, pers. communication 2004). For this reason, models of in situ formation of Pegasi planets generally yield large core masses $\sim 40 \text{ M}_\oplus$ (Bodenheimer et al. 2000). Pegasi planets should therefore be expected to have
very different compositions and core masses, depending on the properties of the disk of planetesimals at their formation, the presence of other planets, and their orbital evolution.

Figure 11: Evolution of giant planets in terms of radius vs. time, for different irradiation levels, and 2 assumed compositions: solar, and 6 times solar. (This calculation ignores second order effects as modifications of the adiabatic temperature gradient and non-linear effects in the opacity calculation, and more importantly modifications of atmospheric properties.).

The presence of a core has a relatively straightforward impact on the evolution of giants planets. As shown in fig. 10, it leads to a much faster contraction and a smaller radius at any given age. An enrichment of the envelope both increases the mean molecular weight and the opacities, with two opposite effects in terms of the planet’s contraction and cooling. Figure 11 shows that for large irradiations (extended radiative zones), the second effect wins and leads to a (limited) increase of the planetary radius. However, planets with a larger mean molecular weight eventually become smaller.

The difference in inferred radii between HD209458b and other transiting planets could hence indicate that stellar tides play a role in slowing or even stalling the contraction of all planets, but that because of different histories,
some planets have a large core mass but HD209458b has not. In that framework, OGLE-TR-132b would probably need a core of \( \sim 20 \, M_\oplus \) or more (or the same amount of heavy elements in its deep interior) to explain its small radius. The large \([\text{Fe}/\text{H}]\) value measured for its parent star (Table \ref{table:stellar_properties}) is an indication that the planet may indeed have grown a large core.

### 4.5 What is the role of the atmosphere for the evolution?

I have purposely used a very simple atmospheric model by setting \( T_1 \propto T_{\text{eq}} = \text{cte} \). Of course, this hides many important complications like opacities, chemistry, gravity dependency, presence of clouds, atmospheric dynamics, dependence on the incoming stellar flux...etc. These complications partially explain differences between several authors (Seager & Sasselov 1998, 2000; Gouken-leuque et al. 2000; Barman et al. 2001; Sudarsky et al. 2003; Iro et al. 2004). These works yield characteristic temperatures at the base of the atmosphere (i.e. where most of the incoming flux has been absorbed) that range from \( \sim 1700 \) to \( \sim 2300 \) K.

However, the largest differences arise from simple geometrical reasons: Because these calculations are one-dimensional, some authors choose to model the atmosphere at the substellar point, some average the received stellar flux over the day-hemisphere (1/2 less flux), and others average it over the entire planet (1/4 less flux). This points to real problems: how does the planet react to this extremely inhomogeneous stellar irradiation, and how do possible inhomogeneities in the atmosphere affect the planetary evolution?

Without atmospheric dynamics, a synchronous Pegasi planet at \( \sim 0.05 \, \text{AU} \) of a G-type star would see its substellar point heated to \( \sim 2500 \, \text{K} \) or more, and its night hemisphere and poles have temperatures \( \sim 100 \, \text{K} \), a clearly unstable situation. Assuming synchronisation of the convective interior and a radiative atmosphere obeying the Richardson shear-instability criterion, Showman & Guillot (2002) showed that the atmosphere of Pegasi planets are likely to develop \( \text{km/s} \) winds, but that spatial photospheric temperature variations of \( \sim 500 \, \text{K} \) are likely. Dynamical models using shallow-water equations by Cho et al. (2003) also yield latitudinal temperature variations, but predict a surprising time-dependent behavior, with a night-side that sometimes becomes hotter than the day side. A time-dependant approach of radiative transfer, in which the atmosphere is allowed to react to a varying irradiation, shows that a \( \text{km/s} \) rotation indeed yields a \( \sim 500 \, \text{K} \) effective temperature variation. It also shows that the conditions required for the shallow-water treatment (a relatively long radiative timescale) are probably not met in Pegasi planets (Iro et al. 2004).

As shown by Guillot & Showman (2002), to first order (i.e. neglecting possible non-linear behavior due to e.g. opacity temperature dependences and/or cloud formation), the cooling with an inhomogeneous boundary condition is faster than if the same amount of heat has been homogeneous distributed. This is because heat tends to escape more rapidly in regions of low atmospheric temperatures. But since the radiative timescale below optical depth unity is
Therefore, there is presently no reason to use for evolution models an atmospheric boundary condition other than that obtained assuming a stellar flux averaged over the entire planet. Of course, more work is to be done, as opacity variations, the presence of clouds either on the day or night side (depending on the kind of circulation), non-equilibrium chemistry, and possible shear instabilities and gravity waves damping can all play an important role.

4.6 Stability and evaporation?

Because Pegasi planets are so close to their star, the question of their survival has been among the first following the discovery of 51 Peg B. Guillot et al. (1996) and Lin et al. (1996) independently concluded to a relatively fast contraction of the planet and to its survival based on non-thermal evaporation rates extrapolated from Jupiter. These evaporation rates $\sim 10^{-16} \text{M}_\oplus \text{a}^{-1}$ turn out to be extremely close to those inferred from observations of HD209458b showing the escape of HI (Vidal-Madjar et al. 2003), OI and CII (Vidal-Madjar et al. 2004). However, the atmospheric escape problem is more complex than initially envisioned, with XUV heating, conduction and gravity waves playing important roles (Lammer et al. 2003; Lecavelier des Etangs et al. 2004).

Generally, a critical question is that of the stability of planets at close orbital distances in their young ages (Baraffe et al. 2004; Gu et al. 2004). Figure 11 shows that the cooling timescale is initially relatively long in the case of intense irradiation (see also fig. 2 of Guillot et al. 1996) and might lead to a significant mass loss in case of a rapid inward migration because of Roche lobe overflow (part of the planetary envelope becomes unbound because of the star’s gravitational potential) (Trilling et al. 1998, 2002). Baraffe et al. (2004) find that another route may be the strong exospheric evaporation. Below a critical mass, the planet would inflate before it can become degenerate enough. However, either the presence of a core and the consequent rapid contraction (see fig. 11), or an internal cooling associated to the decompression upon mass loss may protect the planets from an exponential evaporation.

5 Conclusion & prospects

We are just beginning to discover the diversity of giant planets. Already, a variety of problems that are particular to one planet or a small ensemble of planets have arisen. Given the limited ensemble of objects that we are given to study and the rapid evolution of the subject, any attempt to find general rules it fraught with risk. Some salient conclusions should however resist the trial of time:

- The giant planets of our Solar System all contain a minimum of $10 \text{M}_\oplus$ of heavy elements, and even $\sim 20 \text{M}_\oplus$ for Saturn and probably Jupiter. In Jupiter, most of the heavy elements are mixed in the hydrogen-helium
envelope. On the contrary, Saturn, Uranus and Neptune appear to be significantly differentiated.

- The envelopes of Jupiter, Saturn, Uranus and Neptune are enriched in heavy elements compared to a solar composition, implying that heavy elements were delivered either after the formation (requiring large masses in planetesimals because of the low accretion probabilities) or when the planets, and in particular Jupiter, were not fully formed. In that case, an upward mixing (erosion) of these elements with the envelope is required. A third possibility is that these elements were captured in an enriched nebula.

- The demixing of helium in metallic hydrogen has probably begun in Jupiter, and has been present in Saturn for $2 - 3\, \text{Ga}$.

- Like Jupiter and Saturn, the Pegasi planets discovered so far are mostly made of hydrogen and helium, but their precise composition depends on how tidal effects lead to the dissipation of heat in their interior.

Improvements on our knowledge of the giant planets requires a variety of efforts. Fortunately, nearly all of these are addressed at least partially by adequate projects in the next few years. The efforts that are necessary thus include (but are not limited to):

- Obtain a better EOS of hydrogen, in particular near the molecular/metallic transition. This will be addressed by the construction of powerful lasers such as the NIF in the US and the MégaJoule laser in France, and by innovative experiments such as shocks on pre-compressed samples. One of the challenges is not only obtaining higher pressures, but mostly lower temperatures than currently possible with single shocks. The parallel improvement of computing facilities should allow more extended numerical experiments.

- Calculate hydrogen-helium and hydrogen-water phase diagrams. (Other phase diagrams are desirable too, but of lesser immediate importance). This should be possible with new numerical experiments.

- Have a better yardstick to measure solar and protosolar compositions. This may be addressed by the analysis of the Genesis mission samples, or may require another future mission.

- Improve the values of $J_4$ and $J_6$ for Saturn. This will be done as part of the Cassini-Huygens mission. This should lead to better constraints, and possibly a determination of whether the interior of Saturn rotates as a solid body.

- Detect new transiting extrasolar planets, and hopefully some that are further from their star. The space missions COROT (2006) and Kepler (2007) should provide the detection and characterization of many tens, possibly hundreds of giant planets.
• Improve the measurement of Jupiter’s gravity field, and determine the abundance of water in the deep atmosphere. This would be possible either from an orbiter, or even with a single fly-by (Bolton et al. 2003).

Clearly, there is a lot of work on the road, but the prospects for a much improved knowledge of giant planets and their formation are bright.

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References

Alexander DR & Ferguson JW. 1994. Low-temperature Rosseland opacities. \textit{ApJ} 437:879-91
Allard F, Hauschildt PH, Alexander DR, Tamanai A, & Schweitzer A. 2001. The Limiting Effects of Dust in Brown Dwarf Model Atmospheres. \textit{ApJ} 556:357-72
Alonso R, Brown TM, Torres G, Latham DW, Sozzetti A, et al.. 2004. TrES-1: The Transiting Planet of a Bright K0 V Star. \textit{ApJ} 613:L153
Anders E, Grevesse N. 1989. Abundances of the elements - Meteoritic and solar. \textit{Geochem. Cosmo. Acta} 53:197-214
Anderson JD, Campbell JK, Jacobson RA, Sweetnam DN, Taylor AH. 1987. Radio science with Voyager 2 at Uranus - Results on masses and densities of the planet and five principal satellites. \textit{J. Geophys. Res.} 92:14877-83
Atreya SK, Mahaffy PR, Niemann HB, Wong MH, Owen TC. 2003. Composition and origin of the atmosphere of Jupiter-an update, and implications for the extrasolar giant planets. \textit{Plan. Space Sci.} 51:105-12
Baraffe I, Chabrier G, Barman TS, Allard F, Hauschildt P. 2003. Evolutionary models for cool brown dwarfs and extrasolar giant planets. The case of HD 209458. \textit{A\&A} 402:701-12
Baraffe I, Selsis F, Chabrier G, Barman TS, Allard F et al.. 2004. The effect of evaporation on the evolution of close-in giant planets. \textit{A\&A} 419:L13-16
Barman T, Hauschildt PH, Allard F. 2001. Irradiated Planets. \textit{ApJ} 556:885-95
Belov SI, Boriskov GV, Bykov AI, Il’kaev IL, Luk’yanov NB et al. 2002. Shock Compression of Solid Deuterium. \textit{Soviet Phys. - JETP Lett.} 76:433-5
Bodenheimer P, Hubickyj O & Lissauer JJ. 2000. Models of the in Situ Formation of Detected Extrasolar Giant Planets. \textit{Icarus} 143:2-14
Bodenheimer P, Lin DNC, Mardling R. 2001. On the Tidal Inflation of Short-Period Extrasolar Planets. \textit{ApJ} 548:466-72
Bodenheimer P, Lin DNC. 2002. Implications of Extrasolar Planets for Understanding Planet Formation. *Ann. Rev. Earth Plan. Sci.* 30:113-48

Bodenheimer P, Laughlin G, Lin DNC. 2003. On the Radii of Extrasolar Giant Planets. *ApJ* 592:555-63

Bolton SJ, Allison M, Anderson J, Atreya S, Bagenal F, et al. 2003. A Polar Orbiter to Probe Jupiter’s Deep Atmosphere, Interior Structure and Polar Magnetosphere. *DPS Meeting* #35:#41.08

Bonev SA, Militzer B, Galli G. 2004. Ab initio simulations of dense liquid deuterium: Comparison with gas-gun shock-wave experiments. *Phys. Rev. B* 69:014101

Boss AP. 2000. Possible Rapid Gas Giant Planet Formation in the Solar Nebula and Other Protoplanetary Disks. *ApJ* 536:L101-4

Boriskov GV, Bykov AI, Il'kaev IL, Selemir VD, Sinakov GV et al. 2003. Shock-Wave Compression of Solid Deuterium at a Pressure of 120 GPa. *Dokl. Phys.* 48:553-555

Bouchy F, Pont F, Santos NC, Melo C, Mayor M et al. 2004. Two new “very hot Jupiters” among the OGLE transiting candidates. *A&A* 421:L13-16

Brown TM, Charbonneau D, Gilliland RL, Noyes RW, Burrows A. 2001. Hubble Space Telescope Time-Series Photometry of the Transiting Planet of HD 209458. *ApJ* 552:699-709

Burrows A, Marley MS & Sharp, CM. 2000a. The Near-Infrared and Optical Spectra of Methane Dwarfs and Brown Dwarfs. *ApJ* 531:438-446

Burrows A, Guillot T, Hubbard WB, Marley MS, Saumon D, et al. 2000b. On the Radii of Close-in Giant Planets. *ApJL* 534:L97-100

Burrows A, Hubeny I, Hubbard WB, Sudarsky D, Fortney JJ. 2004. Theoretical Radii of Transiting Giant Planets: The Case of OGLE-TR-56b. *ApJ* 610:L53-6

Campbell JK & Synnott SP. 1985. Gravity field of the Jovian system from Pioneer and Voyager tracking data. *Astron. J.* 90:364-72

Campbell JK & Anderson JD. 1989. Gravity field of the Saturnian system from Pioneer and Voyager tracking data. *Astron. J.* 97:1485-95

Cavazzoni C, Chiarotti GL, Scandolo S, Tosatti E, Bernasconi M, & Parrinello M. 1999. Superionic and Metallic States of Water and Ammonia at Giant Planet Conditions. *Science* 283:44-46

Chandrasekhar S. 1939. “Stellar Structure and Evolution”, The University of Chicago press, Chicago

Charbonneau D, Brown TM, Latham DW, Mayor, M. 2000. Detection of planetary transits across a Sun-like star. *ApJ* 529:L45-48

Charbonneau D, Brown TM, Noyes RW, Gilliland RL. 2002. Detection of an Extrasolar Planet Atmosphere. *ApJ* 568:377-84

Cho JYK, Menou K, Hansen BMS, Seager S. 2003. The Changing Face of the Extrasolar Giant Planet HD 209458b. *ApJ* 587:L117-20

Cody AM, Sasselov DD. 2002. HD 209458: Physical Parameters of the Parent Star and the Transiting Planet. *ApJ* 569:451-58

Cohen ER, Taylor BN. 1986. The 1986 adjustment of the fundamental physical constants.. *Rev. Mod. Phys.* 59:1121
Collins GW, Da Silva LB, Celliers P, et al. 1998. Measurements of the equation of state of deuterium at the fluid insulator-metal transition. *Science* 281:1178-81

Conrath BJ & Gautier D. 2000. Saturn Helium Abundance: A Reanalysis of Voyager Measurements. *Icarus* 144:124-34

da Silva LB, Celliers P, Collins GW, Budil KS, Holmes NC et al. 1997. Absolute Equation of State Measurements on Shocked Liquid Deuterium up to 200 GPa (2 Mbar). *Phys. Rev. Lett.* 78:483-6

Datchi F, Loubeyre P, Letoullec R. 2000. Extended and accurate determination of the melting curves of argon, helium, ice (H2O), and hydrogen (H2). *Phys. Rev. B* 61:6535-46

Davies ME, Abalakin VK, Bursa M, Lederle T, Lieske JH et al. 1986. Report of the IAUIAG COSPAR working group on cartographic coordinates and rotational elements of the planets and satellites: 1985. *Celestial Mech.* 39:102-13

Desjarlais MP. 2003. Density-functional calculations of the liquid deuterium Hugoniot, reshock, and reverberation timing. *Phys. Rev. B* 68:064204

Fortney J, Hubbard WB. 2003. Phase separation in giant planets: inhomogeneous evolution of Saturn. *Icarus* 164:228-43

Gautier D, Owen, T. 1989. The composition of outer planet atmospheres. In *Origin and Evolution of Planetary and Satellite Atmospheres*, ed. SK Atreya, JB Pollack, MS Matthews, 487–512. University of Arizona Press: Tucson

Gautier D, Hersant F, Mousis O & Lunine JI. 2001. Enrichments in Volatiles in Jupiter: A New Interpretation of the Galileo Measurements. *ApJL* 550:L227-230

Goukenleuque C, Bézard B, Joguet B, Lellouch E, Freedman R. 2000. A radiative equilibrium model of 51 Peg b. *Icarus* 143:308

Gregoryanz E, Goncharov AF, Matsuishi K, Mao H, Hemley RJ. 2003. Raman Spectroscopy of Hot Dense Hydrogen. *Phys Rev. Lett.* 90:175701-1-4

Gu P-G, Bodenheimer PH, Lin DNC. 2004. The Internal Structural Adjustment Due to Tidal Heating of Short-Period Inflated Giant Planets. *ApJ* 608:1076-94

Guillot T, Gautier D, Chabrier G, Mosser B. 1994. Are the giant planets fully convective?. *Icarus* 112:337-53

Guillot T, Chabrier G, Gautier D & Morel P. 1995. Effect of radiative transport on the evolution of Jupiter and Saturn. *ApJ* 450:463-72

Guillot T. 1995. Condensation of Methane Ammonia and Water in the Inhibition of Convection in Giant Planets. *Science* 269:1697-99

Guillot T, Burrows A, Hubbard WB, Lunine JI & Saumon D. 1996. Giant planets at small orbital distances. *ApJ* 459:L35-38

Guillot T. 1999a. A comparison of the interiors of Jupiter and Saturn. *Plan. Space. Sci.* 47:1183-200

Guillot T. 1999b. Interior of Giant Planets Inside and Outside the Solar System. *Science* 286:72-77
Guillot T Gladman B. 2000. Late Planetesimal Delivery and the Composition of Giant Planets. In Proceedings of the Disks, Planetesimals and Planets Conference, ASP Conference Series, eds F Garzon et al., 219:475-485.

Guillot T & Showman A. 2002. Evolution of “51 Pegasus b-like” planets. A&A 385:156-65

Guillot T, Stevenson DJ, Hubbard WB, Saumon D. ed. F Bagenal, W McKin-non, T Dowling, in press. In The interior of Jupiter, Jupiter: The Planet, Satellites, and Magnetosphere

Henry GW, Marcy GW, Butler RP, Vogt, SS. 2000. A transiting “51 Peg-like” planet. ApJ 529:L41-44

Hersant F, Gautier D, Lunine JI. 2004. Enrichment in volatiles in the giant planets of the Solar System. Plan. Space Sci. 52:623-41

Holmes NC, Ross M, Nellis WJ. 1995. Temperature measurements and dissociation of shock-compressed liquid deuterium and hydrogen. Phys. Rev. B 52:15835-45

Hubbard WB. 1968. Thermal structure of Jupiter. ApJ 152:745-54

Hubbard WB. 1977. The Jovian surface condition and cooling rate. Icarus 30:305-10

Hubbard WB. 1982. Effects of differential rotation on the gravitational figures of Jupiter and Saturn. Icarus 52:509-15

Hubbard WB, Pearl JC, Podolak M, Stevenson DJ. 1995. The Interior of Neptune. In Neptune and Triton, ed. DP Cruikshank. 109–138. Univ. of Arizona Press: Tucson

Hubbard WB, Guillot T, Marley MS, Burrows A, Lunine JI & Saumon DS. 1999. Comparative evolution of Jupiter and Saturn. Plan. Space. Sci. 47:1175-82

Hubbard WB. 1999. Gravitational signature of Jupiter’s deep zonal flows. Icarus 137:196-99

Hubbard WB, Burrows A, Lunine JI. 2002. Theory of Giant Planets. Ann. Rev. Astron. Astrophys. 40:103-36

Iro N, Bézard B, Guillot T. 2004. A Time-dependent radiative model of HD209458b. submitted to Icarus

Konacki M, Torres G, Jha S, Sasselov DD. 2003. An extrasolar planet that transits the disk of its parent star. Nature 421:507-9

Konacki M, Torres G, Sasselov DD, Pietrzynski G, Udalski A et al.. 2004. The Transiting Extrasolar Giant Planet around the Star OGLE-TR-113. ApJ 609:L37-40

Knudson MD, Hanson DL, Bailey JE, Hall CA, Asay JR, Deeney C. 2004. Principal Hugoniot, reverberating wave, and mechanical reshock measurements of liquid deuterium to 400 GPa using plate impact techniques. Phys. Rev. B 69:144209

Knudson MD, Hanson DL, Bailey JE, Hall CA, Asay JR, Anderson WW. 2002. Equation of State Measurements in Liquid Deuterium to 70 GPa. Phys. Rev. Lett. 87:225501

Lammer H, Selsis F, Ribas I, Guinan EF, Bauer SJ, Weiss WW. 2003. Atmospheric Loss of Exoplanets Resulting from Stellar X-Ray and Extreme-
Ultraviolet Heating. *ApJ* 598:L121-4
Laughlin G, Wolf A, Vanmunster T, Bodenheimer P, Fischer D et al. 2004. A Comparison of Observationally Determined Radii with Theoretical Radius Predictions for Short-Period Transiting Extrasolar Planets. *ApJ* submitted

Lecavelier des Etangs A, Vidal-Madjar A, McConnell JC, Hébrard G. 2004. Atmospheric escape from hot Jupiters. *A&A* 418:L1-4

Lenzini P, Chernoff DF, Salpeter EE. 1991. Rosseland and Planck mean opacities of a zero-metallicity gas. *ApJS* 76:759-801

Levison HF, Morbidelli A. 2003. The formation of the Kuiper belt by the outward transport of bodies during Neptune’s migration. *Nature* 426:419-21

Lin DNC, Bodenheimer P, & Richardson DC. 1996. Orbital migration of the planetary companion of 51 Pegasi to its present location. *Nature* 380:606-7

Lindal GF, Wood GE, Levy GS, Anderson JD, Sweetnam DN, et al. 1981. The atmosphere of Jupiter - an analysis of the Voyager radio occultation measurements. *J. Geophys. Res.* 86:8721-7

Lindal GF, Sweetnam DN, Eshleman VR. 1985. The atmosphere of Saturn – an analysis of the Voyager radio occultation measurements. *Astron. J.* 90:1136-46

Lindal GF. 1992. The atmosphere of Neptune – an analysis of radio occultation data acquired with Voyager 2. *Astron. J.* 103:967-82

Lodders K. 2003. Solar System Abundances and Condensation Temperatures of the Elements. *ApJ* 591:1220-47

Lubow SH, Tout CA & Livio M. 1997. Resonant Tides in Close Orbiting Planets. *ApJ* 484:866-70

Magalhães JA, Seiff A, Young RE. 2002. The Stratification of Jupiter’s Troposphere at the Galileo Probe Entry Site. *Icarus* 158:410-33

Mahaffy PR, Niemann HB, Alpert A, Atreya SK, Demick J et al. 2000. Noble gas abundance and isotope ratios in the atmosphere of Jupiter from the Galileo Probe Mass Spectrometer. *JGR* 105:15061-72

Mao H, Hemley RJ. 1994. Ultrahigh-pressure transitions in solid hydrogen. *Rev. Mod. Phys.* 66:671-92

Marcy GW, Butler RP, Williams E, Bildsten L, Graham JR, Ghez AM, Jernigan JG. 1997. The Planet around 51 Pegasi. *ApJ* 481:926-35

Marley MS, Gomez P & Podolak P. 1995. Monte Carlo interior models for Uranus and Neptune. *J. Geophys. Res.* 100:23349-54

Marley MS & McKay CP. 1999. Thermal Structure of Uranus’ Atmosphere. *Icarus* 138:268-86

Militzer B & Ceperley DM. 2001. Path Integral Monte Carlo Simulation of the Low-Density Hydrogen Plasma. *Phys. Rev. E* 63:6404

Moutou C, Pont F, Bouchy F & Mayor M. 2004. Accurate radius and mass of the transiting exoplanet OGLE-TR-132b. *A&A* 424:L31

Niemann HB, Atreya SK, Carignan GR, Donahue TM, Haberman JA, et al. 1998. The composition of the jovian atmosphere as determined by the Galileo probe mass spectrometer.. *J. Geophys. Res.* 103:22831-8
Ogilvie GI, Lin DNC. 2004. Tidal Dissipation in Rotating Giant Planets. *ApJ* 610:477-509

Owen T, Mahaffy P, Niemann HB, Atreya S, Donahue T, Bar-Nun A & de Pater I. 1999. A low-temperature origin for the planetesimals that formed Jupiter. *Nature* 402:269-70

Pätzold M, Rauer H. 2002. Where Are the Massive Close-in Extrasolar Planets?. *ApJ* 568:L117-20

Pearl JC & Conrath BJ. 1991. The albedo, effective temperature, and energy balance of Neptune, as determined from Voyager data. *J. Geophys. Res. Suppl.* 96:18921-9

Pfaffenzeller O, Hohl D & Ballone P. 1995. Miscibility of hydrogen and helium under astrophysical conditions. *Phys Rev. Lett.* 74:2599-602

Podolak M, Hubbard WB, Stevenson DJ. 1991. Model of Uranus’ interior and magnetic field. In *Uranus*, ed. JT Bergstralh, ED Miner, MS Matthews. 29-61. Univ. of Arizona Press: Tucson

Podolak M, Weizman A & Marley MS. 1995. Comparative models of Uranus and Neptune. *Plan. Space. Sci.* 43:1517-22

Podolak M, Podolak JI & Marley MS. 2000. Further investigations of random models of Uranus and Neptune. *Plan. Space. Sci.* 48:143-51

Pollack JB, Hubickyj O, Bodenheimer P, Lissauer JJ, Podolak M & Greenzweig Y. 1996. Formation of the Giant Planets by Concurrent Accretion of Solids and Gas. *Icarus* 124:62

Pont F, Bouchy F, Queloz D, Santos NC, Melo C, Mayor M, Udry S. 2004. The “missing link”: A 4-day period transiting exoplanet around OGLE-TR-111. *A&A* 426:L15

Rasio F, Tout CA, Lubow SH, Livio M. 1996. Tidal Decay of Close Planetary Orbits. *ApJ* 470:1187-91

Ross M. 1998. Linear-mixing model for shock-compressed liquid deuterium. *Phys Rev. B.* 58:669

Ross M, Yang LH. 2001. Effect of chainlike structures on shock-compressed liquid deuterium. *Phys Rev. B* 64:134210

Roos-Serote M, Atreya SK, Wong MK, Drossart P. 2004. On the water abundance in the atmosphere of Jupiter. *Plan. Space Sci.* 52:397-414

Roulston, MS, Stevenson, DJ. 1995. Prediction of neon depletion in Jupiter’s atmosphere. *EOS* 76: 343 [abstract]

Sasselov DD. 2003. The new transiting planet OGLE-TR-56b: Orbit and atmosphere. *ApJ* 596:1327-31

Saumon D, Hubbard WB, Chabrier G, Van Horn HM. 1992. The role of the molecular-metallic transition of hydrogen in the evolution of Jupiter, Saturn and brown dwarfs. *ApJ* 391:827-31

Saumon D, Chabrier G & Van Horn HM. 1995. An equation of state for low-mass stars and giant planets. *ApJS* 99:713-41

Saumon D, Hubbard WB, Burrows A, Guillot T, Lunine JI & Chabrier G. 1996. A Theory of Extrasolar Giant Planets. *ApJ* 460:993-1018

Saumon D, Chabrier G, Wagner DJ, & Xie X. 2000. Modeling Pressure-Ionization of Hydrogen in the Context of Astrophysics. *High Pressure Research*
Saumon D, Guillot T. 2004. Shock Compression of Deuterium and the Interiors of Jupiter and Saturn. *ApJ* 609:1170-80

Seager S & Sasselov DD. 1998. Extrasolar giant planets under strong stellar irradiation. *ApJ* 502:L157-60

Seager S & Sasselov DD. 2000. Theoretical Transmission Spectra during Extrasolar Giant Planet Transits. *ApJ* 537:916-21

Seager S & Hui L. 2002. Constraining the Rotation Rate of Transiting Extrasolar Planets by Oblateness Measurements. *ApJ* 574:1004-10

Seiff A, Kirk DB, Knight TCD, Young RE, Mihalov JD et al. 1998. Thermal structure of Jupiter’s atmosphere near the edge of a 5-μm hot spot in the north equatorial belt. *J. Geophys. Res.* 103:22857-90

Showman AP & Guillot T. 2002. Atmospheric circulation and tides of “51 Pegasus b-like” planets. *A&A* 385:166-80

Sozzetti A, Young D, Torres G, Charbonneau D, Latham DW, et al. 2004. High-Resolution Spectroscopy of the Transiting Planet Host Star TrES-1, ApJL, in press (astro-ph/0410483)

Stevenson DJ & Salpeter EE. 1977. The dynamics and helium distribution in hydrogen-helium fluid planets. *ApJ Suppl* 35:239-61

Stevenson DJ. 1982. Interiors of the giant planets. *Ann. Rev. Earth Planet. Sci.* 10:257-95

Stevenson DJ. 1985. Cosmochemistry and structure of the giant planets and their satellites. *Icarus* 62:4-15

Sudarsky D, Burrows A, Hubeny I. 2003. Theoretical Spectra and Atmospheres of Extrasolar Giant Planets. *ApJ* 588:1121-48

Torres G, Konacki M, Sasselov DD, Jha S. 2004. New Data and Improved Parameters for the Extrasolar Transiting Planet OGLE-TR-56b. *ApJ* 609:1071-5

Trilling DE, Benz W, Guillot T, Lunine JI, Hubbard WB, & Burrows A. 1998. Orbital Evolution and Migration of Giant Planets: Modeling Extrasolar Planets. *ApJ* 500:428-39

Trilling DE, Lunine JI, Benz W. 2002. Orbital migration and the frequency of giant planet formation. *A&A* 394:241-51

Tyler GL, Sweetnam DN, Anderson JD, Borutzki SE, Campbell JK, et al. 1989. Voyager radio science observations of Neptune and Triton. *Science* 246:1466-73

Vidal-Madjar A, Désert J-M, Lecavelier des Etangs A, Hébrard G, Ballester GE et al. 2004. Detection of Oxygen and Carbon in the Hydrodynamically Escaping Atmosphere of the Extrasolar Planet HD 209458b. *ApJ* 604:L69-72

Vidal-Madjar A, Lecavelier des Etangs A, Désert J-M, Ballester GE, Ferlet R, et al. 2003. An extended upper atmosphere around the extrasolar planet HD209458b. *Nature* 422:143-6

Warwick JW, Evans DR, Roming JH, Sawyer CB, Desch MD, et al. 1986. Voyager 2 radio observations of Uranus. *Science* 233:102-6
Warwick JW, Evans DR, Peltzer GR, Peltzer RG, Roming JH, et al. 1989. Voyager planetary radio astronomy at Neptune. *Science* 246:1498-501

Weir ST, Mitchell AC, Nellis WJ. 1996. Metallization of Fluid Molecular Hydrogen at 140 GPa (1.4 Mbar). *Phys. Rev. Lett.* 76:1860-3

Witte MG, Savonije GJ. 2002. Orbital evolution by dynamical tides in solar type stars. Application to binary stars and planetary orbits. *A&A* 386:222-36

von Zahn U, Hunten DM, Lehmacher G. 1998. Helium in Jupiter’s atmosphere: results from the Galileo probe helium interferometer experiment. *J. Geophys. Res.* 103:22815-30

Zharkov VN & Trubitsyn VP. 1974. Determination of the equation of state of the molecular envelopes of Jupiter and Saturn from their gravitational moments. *Icarus* 21:152-6

Zharkov VN & Trubitsyn VP. 1978. “Physics of Planetary Internals”, Ed. WB Hubbard, Pachart:Tucson