Origin of solar systems
30 June - 2 July 2009
Klaus Jockers (jockers@mps.mpg.de)
Max-Planck-Institut of Solar System Science
Katlenburg-Lindau

Each day there are three time slots:
1. 10:00 – 11:00
2. 11:15 – 12:15
3. 14:00 – 15:00

Included in this time: 15 minutes for discussion
If there is more than one speaker for one lesson, speaking time and discussion time must be shared.
Overview

Tue. June 30:
10:00  1. Physics and chemistry of molecular regions in the interstellar medium (dark clouds and dissociation regions): Supriya Deshpande, Antoine Genetelli and Judith de Patoul
11:15  2. Cloud collapse (Jeans mass, free fall time, opacity limit) + numerical calculations of cloud collapse (movies from internet), initial mass function: Tilaye Tadesse, Klaus Jockers
14:00  3. Early phases of protostars – observations: microwave, infrared, visual, X-rays: Maria Dasi, Yeon Joo Lee, and Julia Thalmann

Wed. July 01:
10:00  4. Equilibria of protostars and gas planets: Armando Gonzalez and Peter Kollmann, talk to be given by Klaus Jockers
11:15  5. Condensation and growth of solid bodies in proto-planetary disks: Klaus Jockers
14:00  6. Equilibrium condensation of a solar nebula + explanation of the minerals involved: Klaus Jockers

Thu. July 02:
10:00  7. Dynamical and physical properties of extrasolar planets: Ronny Lutz and Anne Angsmann
11:15  8. Dynamics of comets and the Kuiper belt: Jean-Baptiste Vincent and Yacine Saidi
14:00  9. Physical properties of Trans Neptunian Objects and comets: Tadese Ejeta and Manuela Lippi
Physics and chemistry of molecular regions in the interstellar medium

Lecture on the Origins of the solar system
30th of June 2009

Supriya Desphande
Judith de Patoul
Antoine Genetelli
Overview

• The Interstellar Medium
• Molecular Clouds
• Observations
• Interstellar Chemistry
• Photodissociation Regions
• Summary


Lequeux J., "The interstellar medium", Springer 2005 => Overview + Chapter, 9-10
The Interstellar Medium
The Interstellar Medium (ISM)

70% Hydrogen, 28% Helium and 2% of heavier elements (C,N,O,Mg,Si)

≈0.5% of the total mass of the Galaxy

- ‘Empty space’ between the stars
- 99 % gas, 1% dust
Example

- NGC 7331 (Pegasus constellation, 50 Million light years)
- Shorter wavelengths (3.6 to 4.5 μm): older and cooler stars than the sun
- Longer wavelengths (5 to 8 μm): glow from dust of interstellar dust
Another one: M51
Molecular Clouds
Molecular Clouds

- Diffuse Molecular Clouds
  - No star formation (from observation)
- Translucent Clouds
  - No star formation (from observation)
- Dark Molecular Clouds (DMC)
  - Low mass star formation
- Giant Molecular Clouds (GMC)
  - Low and Massive star formation

? The Sun was formed in DMC or GMC ?

=> Both scenarios need to be investigated
Dark Molecular Clouds (DMC)

- Observation:
  - Visible as dark patches on the sky
  - Show a complex morphology
- Places for stars formation with low mass
- Sites of many complex molecules
- EX: Taurus
Giant Molecular Cloud

- Similarity complex morphology
- Similarity density
- More Massive
- More warm
- Place for Low and Massive-star formation
- Eg: Orion Molecular Cloud

<table>
<thead>
<tr>
<th>Giant Molecular Clouds</th>
<th>Density ($\text{cm}^{-3}$)</th>
<th>$T$ (K)</th>
<th>Mass ($M_\odot$)</th>
<th>$A_V$ (mag)</th>
<th>Size (pc)</th>
<th>$\Delta V$ (km s$^{-1}$)</th>
<th>Examples</th>
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<tbody>
<tr>
<td>complex</td>
<td>100 – 300</td>
<td>15 – 20</td>
<td>$10^5$ – $3 \times 10^6$</td>
<td>1 – 2</td>
<td>20 – 80</td>
<td>6 – 15</td>
<td>M 17, Orion</td>
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<td>clouds</td>
<td>$10^2$ – $10^4$</td>
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<td>3 – 12</td>
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<td>warm clumps</td>
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<td>1 – $10^3$</td>
<td>5 – 1000</td>
<td>0.05 – 3</td>
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<td>hot cores</td>
<td>$10^7$ – $10^9$</td>
<td>100 – 200</td>
<td>10 – $10^3$</td>
<td>50 – 1000</td>
<td>0.05 – 1</td>
<td>1 – 10</td>
<td>Orion hot core</td>
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</table>
Molecular Clouds

Table 3.1 Physical Properties of Molecular Clouds

<table>
<thead>
<tr>
<th>Cloud Type</th>
<th>$A_V$ (mag)</th>
<th>$n_{tot}$ (cm$^{-3}$)</th>
<th>$L$ (pc)</th>
<th>$T$ (K)</th>
<th>$M$ (M$_{\odot}$)</th>
<th>Examples</th>
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</thead>
<tbody>
<tr>
<td>Diffuse</td>
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<td>500</td>
<td>3</td>
<td>50</td>
<td>50</td>
<td>ζ Ophiuchi</td>
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<tr>
<td>Giant Molecular Clouds</td>
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<td>100</td>
<td>50</td>
<td>15</td>
<td>10$^4$</td>
<td>Orion</td>
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<tr>
<td>Dark Clouds</td>
<td></td>
<td></td>
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<tr>
<td>Complexes</td>
<td>5</td>
<td>500</td>
<td>10</td>
<td>10</td>
<td>10$^4$</td>
<td>Taurus-Auriga</td>
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<tr>
<td>Individual</td>
<td>10</td>
<td>$10^3$</td>
<td>2</td>
<td>10</td>
<td>30</td>
<td>B1</td>
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<tr>
<td>Dense Cores/Bok Globules</td>
<td>10</td>
<td>$10^4$</td>
<td>0.1</td>
<td>10</td>
<td>10</td>
<td>TMC-1/B335</td>
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</tbody>
</table>

Figure 1.1 A portion of the Northern sky. The Milky Way is depicted as light grey, while the darker patches indicate giant molecular clouds. Also shown, according to their relative brightness, are the more prominent stars, along with principle constellations.
Molecular Clouds

### Table 3.1 Physical Properties of Molecular Clouds

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<th>$n_{\text{tot}}$ (cm$^{-3}$)</th>
<th>$L$ (pc)</th>
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<th>$M$ ($M_\odot$)</th>
<th>Examples</th>
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<td>TMC-1/B335</td>
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</tbody>
</table>

The Formation of Stars, Steven W. Stahler and Francesco Pallis
Copyright © 2004 Wiley-VCH Verlag GmbH & Co. KGaA, Weinheim
ISBN: 3-527-40559-9

![Image of molecular clouds with labels](image_url)
Taurus Molecular cloud seen in:
- visible (Left side)
- the CO line (right side)
Molecular Clouds

- Some other examples of Dark Molecular cloud:
  - Right Picture: DMC in the plane of the Milky-Way
  - Left Picture: DMC in the Eagle nebulae (M16)
# Molecular Clouds

## Table 3.1 Physical Properties of Molecular Clouds

<table>
<thead>
<tr>
<th>Cloud Type</th>
<th>A\textsubscript{V} (mag)</th>
<th>n\textsubscript{tot} (cm\textsuperscript{-3})</th>
<th>L (pc)</th>
<th>T (K)</th>
<th>M (M\textsubscript{\odot})</th>
<th>Examples</th>
</tr>
</thead>
<tbody>
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<td>ζ Ophiuchi</td>
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<td>Giant Molecular Clouds</td>
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<td>50</td>
<td>15</td>
<td>10\textsuperscript{4}</td>
<td>Orion</td>
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<tr>
<td>Dark Clouds</td>
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<tr>
<td>Complexes</td>
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<tr>
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<tr>
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<td>0.1</td>
<td>10</td>
<td>10</td>
<td>TMC-1/B335</td>
</tr>
</tbody>
</table>

Observations
Observations

• Detection of DMC:
  ✷ Infra-Red
  ✷ Microwave
  ✷ millimeter and submillimeter

• Examples:
  ➡ Orion Region
  ➡ DMC B68 in Ophiuchus
Observations

Why DMC are detected in Infrared?

- Presence of dust
  - Absorption more in shorter wavelengths than in Infrared wavelengths
- The dust reemit in the Infrared

Horsehead nebula in Orion: visible (left), infrared (right)
Observations

- Spitzer (2003):
  - IRAC: InfraRed Array Camera
  - MIPS: Multiband Imaging Photometer for Spitzer

- IRAM: 30m radiotelescope

- Odin (2001):
  - OSIRIS: Odin Spectrometer and InfraRed Imaging System

- And others ...
Interstellar Chemistry
# Interstellar Chemistry

What are the processes involved?

## TABLE A1

### Classes of Chemical Reactions

<table>
<thead>
<tr>
<th>Type</th>
<th>Process</th>
<th>Rate Coefficient</th>
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</thead>
<tbody>
<tr>
<td><strong>Formation Processes</strong></td>
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<tr>
<td>Radiative association</td>
<td>$X + Y \rightarrow XY + h\nu$</td>
<td>$10^{-16} - 10^{-9}$</td>
</tr>
<tr>
<td>Grain surface formation</td>
<td>$X + Y: g \rightarrow XY + g$</td>
<td>$\sim 10^{-18}$</td>
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<tr>
<td><strong>Destruction Processes</strong></td>
<td></td>
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<tr>
<td>Photodissociation</td>
<td>$XY + h\nu \rightarrow X + Y$</td>
<td>$\sim 10^{-10} - 10^{-8} \text{ s}^{-1}$</td>
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<tr>
<td>Dissociative recombination</td>
<td>$XY^+ + e \rightarrow X + Y$</td>
<td>$\sim 10^{-6}$</td>
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<tr>
<td>Collisional dissociation</td>
<td>$XY + M \rightarrow X + Y + M$</td>
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<tr>
<td><strong>Chemical Processes</strong></td>
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<tr>
<td>Ion-molecule exchange</td>
<td>$X^+ + YZ \rightarrow XY^+ + Z$</td>
<td>$\sim 10^{-9}$</td>
</tr>
<tr>
<td>Charge-transfer</td>
<td>$X^+ + YZ \rightarrow X + YZ^+$</td>
<td>$\sim 10^{-9}$</td>
</tr>
<tr>
<td>Neutral-neutral</td>
<td>$X + YZ \rightarrow XY + Z$</td>
<td>$\sim 10^{-12}$</td>
</tr>
</tbody>
</table>

*Approximate rate coefficients appropriate for cold dark clouds. All rate coefficients are sensitive to temperature. For photodissociation, the rates in $\text{s}^{-1}$ in the unattenuated interstellar radiation field are listed.*
Interstellar Chemistry
Build-up of complex molecules

- Essential facts and assumptions:
  - $H_2$ is the most abundant constituent
  - Presence of sufficient amount of reactive ions (e.g. C, S, Si in diffuse clouds)

- Cosmic Rays are the most important source of ionization in DMC
  - Cosmic Rays penetrate into molecular clouds up to column density $10^{24}$ cm$^{-2}$.

- Cosmic Rays ionize Hydrogen at rates $10^{-17}$ to $10^{-16}$ s$^{-1}$.

- $H_2^+ + H_2 \rightarrow H_3^+ + H$

  $X + H_3^+ \rightarrow XH^+ + H_2$  
  Neutral atom X (X=O, C, S)

- $H_3^+$ is the key to molecule formation in Dark Molecular Clouds
Interstellar Chemistry
But why?

- Gas phase reactions with O:
  \[ H^+ + O \rightarrow O^+ + H \]
  \[ O^+ + H_2 \rightarrow \text{OH}^+ + H \]
- or:
  \[ H_3^+ + O \rightarrow \text{OH}^+ + H_2 \]
  → exothermal
- Similar reactions are possible with C.
- N reactions must proceed in a different way as
  \[ N + H_3^+ \rightarrow \text{NH}^+ + H_2 \] is endothermic.
  \[ N^+ \] must be formed first.

Chemical Network of O
Interstellar Chemistry

Interstellar molecules are involved and formed

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</tbody>
</table>

Table 1. Interstellar molecules (September 1993).
Interstellar Chemistry

$H_2$ not directly observed in DMC!

- $H_2$ is the most abundant molecule in space
- Vibrational transition
  $v = 1 \rightarrow 0$; $S(1)$ at $2.1 \mu m$
- Typical DMC Temperature
  $\approx 10$ K

$\Rightarrow$ No observation of rotational or vibrational transitions in DMC

![Diagram](image)

**Fig. 1.** Schematic of the first levels of the $H_2$ molecule. The even-$J$ levels are the para-hydrogen, while the odd-$J$ levels are from the ortho-hydrogen. The coupled levels obey $\Delta J = \pm 2$, for quadrupolar transitions, and the two species are not radiatively coupled. The four first rotational lines of hydrogen $S(J)$ (where $J$ is the lower state) are represented — wavelengths in $\mu m$, $S(3) = 9$, $S(2) = 12$, $S(1) = 17$ and $S(0) = 28$. 

\[ h\nu = hc/\lambda = kT \]

\[ T = hc/k \lambda \]

\[ = 1.439/(\lambda \text{ in cm}) \]
### Table 5.1 Some Useful Molecules

<table>
<thead>
<tr>
<th>molecule</th>
<th>abundance$^a$</th>
<th>transition</th>
<th>type</th>
<th>λ</th>
<th>$T_0^b$</th>
<th>$A_{ul}$</th>
<th>$n_{crit}^c$</th>
<th>comments</th>
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<tbody>
<tr>
<td>H$_2$</td>
<td>1</td>
<td>1→0 S(1)</td>
<td>vibrational</td>
<td>2.1 μm</td>
<td>6600</td>
<td>8.5×10$^{-7}$</td>
<td>7.8×10$^7$</td>
<td>shock tracer</td>
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<tr>
<td>CO</td>
<td>8×10$^{-5}$</td>
<td>J= 1 → 0</td>
<td>rotational</td>
<td>2.6 mm</td>
<td>5.5</td>
<td>7.5×10$^{-8}$</td>
<td>3.0×10$^3$</td>
<td>low density probe</td>
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<tr>
<td>OH</td>
<td>3×10$^{-7}$</td>
<td>$^2\Pi_{3/2}$;J=3/2</td>
<td>Λ-doubling</td>
<td>18 cm</td>
<td>0.08</td>
<td>7.2×10$^{-11}$</td>
<td>1.4×10$^0$</td>
<td>magnetic field probe</td>
</tr>
<tr>
<td>NH$_3$</td>
<td>2×10$^{-8}$</td>
<td>(J,K)=(1,1)</td>
<td>inversion</td>
<td>1.3 cm</td>
<td>1.1</td>
<td>1.7×10$^{-7}$</td>
<td>1.9×10$^4$</td>
<td>temperature probe</td>
</tr>
<tr>
<td>H$_2$CO</td>
<td>2×10$^{-8}$</td>
<td>$2_{12}$→$1_{11}$</td>
<td>rotational</td>
<td>2.1 mm</td>
<td>6.9</td>
<td>5.3×10$^{-6}$</td>
<td>1.3×10$^6$</td>
<td>high density probe</td>
</tr>
<tr>
<td>CS</td>
<td>1×10$^{-8}$</td>
<td>J= 2 → 1</td>
<td>rotational</td>
<td>3.1 mm</td>
<td>4.6</td>
<td>1.7×10$^{-5}$</td>
<td>4.2×10$^5$</td>
<td>high density probe</td>
</tr>
<tr>
<td>HCO$^+$</td>
<td>8×10$^{-9}$</td>
<td>J= 1 → 0</td>
<td>rotational</td>
<td>3.4 mm</td>
<td>4.3</td>
<td>5.5×10$^{-5}$</td>
<td>1.5×10$^5$</td>
<td>tracer of ionization</td>
</tr>
<tr>
<td>H$_2$O</td>
<td>$&lt;7×10^{-8}$</td>
<td>$6_{16}$→$5_{23}$</td>
<td>rotational</td>
<td>1.3 cm</td>
<td>1.1</td>
<td>1.9×10$^{-9}$</td>
<td>1.4×10$^3$</td>
<td>maser</td>
</tr>
</tbody>
</table>

$^a$ number density of main isotope relative to hydrogen, as measured in the dense core TMC–1

$^b$ equivalent temperature of the transition energy; $T_0 \equiv \Delta E_{ul}/k_B$

$^c$ evaluated at T=10 K, except for H$_2$ (T=2000 K) and H$_2$O at 527 μm (T=20 K)
Polycyclic Aromatic Hydrocarbon (PAH)

- Large robust organic molecules
- Found in interstellar medium, in comets, and in meteorites.
- PAHs with grains play an important role in the heating of interstellar gas
- Candidate molecule to act as a basis for the earliest forms of life.

The PAH world hypothesis (biological hypothesis) PAH was a means for a pre-RNA World basis for the origin of life. As yet it is untested, though in 2007 Cassini spacecraft found the presence of heavy negative ions of tholin in the upper regions of Titan’s atmosphere.
Photodissociation Regions
Photodissociation Regions (PDRs)

- Part of the interstellar medium where the ultraviolet radiation field is strong enough to photodissociate molecules: \( XY + h\nu \rightarrow X + Y \)

- Heating and chemistry are regulated by farultraviolet photons.

- The study of the PDRs is understanding
  - The effects of stellar far-ultraviolet photons on the structure, chemistry and thermal balance
  - The evolution of the neutral interstellar medium.
  - Understanding the process of star formation:
    *Farultraviolet photons illuminate star-forming regions, causing them to glow in infrared emission, and play an important role in regulating the star formation process*
Photodissociation Regions (PDRs)
Orion Bar Region

- PAH feature (blue),
- H2 emission (yellow),
- CO emission (red).
- The PDR is seen edge on (Green).
- Position (0,0) corresponds to the a star.
- The illuminating source, the star and the ionized gas are in upper right
Photodissociation Regions (PDRs)
Schematic of PDRs
Photodissociation Regions (PDRs)

Schematic of PDRs

Ionisation front
H⁺ → H
O⁺ → O
C → C⁺
Photodissociation Regions (PDRs)

Schematic of PDRs

FUV absorbed by
- PAHs
- grains

Absorbed energy is used
- excite the PAHs
- heat the grains

(~90%) Energy is converted to
- PAH infrared features
- Far infrared continuum radiation of the cooling grains.

(~1%) Energy is converted to
- energetic photoelectrons that heat the gas ("photoelectric heating").
Photodissociation Regions (PDRs)

Schematic of PDRs

- Dissociated front: $\text{H} \rightarrow \text{H}_2$

- FUV flux

- Hot Stars

- Molecular Cloud

- $T_{\text{gas}} = 10^4K \quad 10^3K \quad 300K \quad 10^2K \quad 20K$

- AV (in log scale)

- PAHs + Grains

- $\text{H}^+ \quad \text{H}^+/\text{H}$
Photodissociation Regions (PDRs)

Schematic of PDRs

Transition: $\text{C}^+ \rightarrow \text{C} \rightarrow \text{CO}$
Photodissociation Regions (PDRs)

Schematic of PDRs

Transition: O $\rightarrow$ O$_2$
Photodissociation Regions (PDRs)
Structure of the PDR in Orion (1/2)
Photodissociation Regions (PDRs)
Structure of the PDR in Orion (2/2)
Photodissociation Regions (PDRs)

Penetration of Far-Ultraviolet radiation
Photodissociation Regions (PDRs)

Chemistry in PDRs (1/2)

Can we observe H2 in the PDR? YES

- Pumping,
- Dissociation,
- Heating.

Photodissociation:

\[ XY + h\nu \rightarrow X + Y \]

(Destruction process)
Photodissociation Regions (PDRs)

Chemistry in PDRs (2/2)

Chemical network: Example of the Oxygen

In DMC

In PDRs
Summary

• Interstellar Medium
  Hydrogen, Helium and traces of heavier elements (C,N,O,Mg,Si)
  Highly Inhomogeneous

• Molecular Clouds
  Dark molecular clouds
  Giant molecular clouds
  Observation

• Interstellar Chemistry
  Different reactions
  Molecular network
  Set of complex molecules + PAHs

• Photodissociation Regions (PDRs)
  Schema and Structure of the PDRs
  Penetration of Far-Ultraviolet radiation in the PDRs
  Chemistry, H₂ and chemical network
Cloud collapse

T. Tadesse

30 June - 2 July 2009, Course on Origin of solar system

Max-Planck Institute for Solar System Research, Katlenburg-Lindau
Outline

- Introduction
- Protostar Formation
- Jeans criterion
  - Jeans mass & radius
- HR-diagram & Hayashi limit
**Cloud collapse**

**Introduction**

- *Stars are formed out of gas and dust which is giant molecular clouds.*
- *Stars form as the result of the gravitational instability and collapse of condensations or dense regions in the interstellar medium*
- *A protostar is a star in the very earliest stage of development, when interstellar gas is still undergoing gravitational collapse. (Pre-nuclear burning object).*
Protostar Formation

- Stars form when over dense regions inside molecular clouds collapse.

How does density get high enough to cause collapse?
Can happen a variety of ways

- collisions between clouds
- interactions with nearby galaxies
- pressure from stellar winds, supernovae (shock wave from supernova or colliding gas clouds) causes a local region of the ISM to compress.

Lots of ongoing research!

- How does each mechanism operate?
- Which mechanisms are most important?
Protostar Formation

If molecular clouds are the sites of star formation what conditions must exist for collapse to occur?

- James Jeans investigated this with simplified assumptions, considering the effects of small deviations from hydrostatic equilibrium with spherical symmetric collapse and neglecting the effects due to rotation and galactic magnetic fields.

- The Jeans instability causes the collapse of interstellar gas clouds and subsequent star formation. It occurs when the internal gas pressure is not strong enough to prevent gravitational collapse of a region filled with matter.

- For stability, the cloud must be in hydrostatic equilibrium

\[
\frac{\partial P}{\partial r} = - \frac{G \rho M_{\text{enc}}}{r^2}
\]
Hydrostatic Equilibrium

- The equilibrium is stable if small perturbations are damped and unstable if they are amplified. The cloud is unstable if it is either very massive at a given temperature or very cool at a given mass for gravity to overcome the gas pressure.

- This is known as the law of *hydrostatic equilibrium*.

- When stars are *not* in hydrostatic equilibrium, they will either expand or contract.

- Let us estimate the conditions for cloud collapse by Virial theorem considering isolated cloud with no external pressure.
**Estimate of the conditions for cloud collapse**

**The Virial Theorem**

- Applies to any system of particles with pair interactions for which the distribution of particles does not vary with time.
- Theorem states that total energy of system $E$ is related to gravitational potential energy $U$ by:
  \[ E = \frac{1}{2} U \]
- But we know that total energy is sum of the kinetic and potential energy:
  \[ K + U = \frac{1}{2} U \]
  or \[ 2K + U = 0 \]
- Can be applied to a system of gravitationally interacting bodies such as stars forming a cloud, clusters of stars, clusters of galaxies, etc.
- As $K = \frac{3}{2} N k T$ and
  \[ U = -\frac{3GM_c^2}{5R_c} = 2K + U = 3NkT - \frac{3GM_c^2}{5R_c} = 0 \]
- Where $U$ is gravitational potential energy of a spherical cloud of constant density.
Jeans Mass- Applying Virial Theorem

• VT can be used to estimate conditions for cloud collapse:

1. If $2K > U \Rightarrow$ gas pressure (energy) will exceed gravitational potential energy and expand.
2. If $2K < U \Rightarrow$ gravitational energy will exceed gas pressure and collapse.

➢ The boundary between these two cases describes the critical condition for stability.

➢ We know that $\rho = \frac{M_c}{V_c} = \frac{M_c}{(4/3\pi R_c^3)} = 3M_c/(4\pi R_c^3)$

$\Rightarrow R_c = (3M_c/4\pi \rho)^{1/3}$

and $N = M_c/\mu m_H$ where $\mu$ is the mean molecular weight and $m_H$ is the mass of a proton. Assuming constant density.
Jeans Mass

Can we derive an expression for the critical mass?

• Hence we can get expression for Jeans mass as:

$$\frac{3M_c kT}{\mu m_H} = \frac{3GM_c^2}{5}\left(\frac{3M_c}{4\pi \rho}\right)^{-1/3}$$

$$2K = -U$$

$$= > M_J = \left(\frac{5kT}{G\mu m_H}\right)^{3/2}\left(\frac{3}{4\pi \rho}\right)^{1/2}$$

• The Jeans Criterion is: \( M_c > M_J \)

• Jeans mass:- Minimum or critical mass that is necessary to initiate spontaneous collapse of the cloud.

• If \( M_c > M_J \) => cloud will collapse.

• \( M_J \) can be written as function of temperature and density as:

$$M_J = \left(\frac{375k^3}{4\pi \mu^3 m_H^3 G^3}\right)^{1/2}\left(\frac{T^3}{\rho}\right)^{1/2}$$

$$= > M_J = \frac{T^3}{\rho}^{1/2}$$
Cloud collapse

Jeans Radius

• *Is there a critical radius that corresponds to the critical mass?*

• *We know* $M_c = \frac{4}{3} \pi R_c^3 \rho$. *Equate this to the Jeans Mass gives:*

$$\frac{4}{3} \pi R^3 \rho = \left( \frac{375k^3}{4\pi \mu^3 m_H^3 G^3} \right)^{1/2} \left( \frac{T^3}{\rho} \right)^{1/2}$$

$$\Rightarrow R^3 = \left( \frac{375}{4\pi} \right)^{1/2} \left( \frac{k}{\mu m_H G} \right)^{3/2} \left( \frac{T}{\rho} \right)^{3/2} \left( \frac{3}{4\pi} \right)$$

$$\therefore R_J = \left( \frac{15k}{4\pi \mu m_H G} \right)^{1/2} \left( \frac{T}{\rho} \right)^{1/2}$$

*The Jeans Radius*

• *If the cloud is compressed to a radius smaller than this critical radius, stable hydrostatic equilibrium is no longer possible and the cloud must collapse.*

• *If* $R_c > R_J$ *=> stable. If* $R_c < R_J$ *=> unstable and collapse.*
**Gravitational Collapse in ISM**

- **Properties of the ISM:**
  
<table>
<thead>
<tr>
<th></th>
<th>Diffuse HI Cloud</th>
<th>H$_2$ Cloud Core</th>
</tr>
</thead>
<tbody>
<tr>
<td>$T$</td>
<td>50 $K$</td>
<td>10 $K$</td>
</tr>
<tr>
<td>$\rho$</td>
<td>500 cm$^{-3}$</td>
<td>$10^8$ cm$^{-3}$</td>
</tr>
<tr>
<td>$M_c$</td>
<td>1-100 $M_{\text{Sun}}$</td>
<td>10-1000 $M_{\text{Sun}}$</td>
</tr>
</tbody>
</table>

- **We know from the Jeans Criterion that if $M_c>M_J$ collapse occurs.**

- **Substituting the values from the table into**
  
  \[ M_J = \left( \frac{5kT}{G\mu m_H} \right)^{3/2} \left( \frac{3}{4\pi\rho} \right)^{1/2} \]

  - **Diffuse HI cloud:** $M_J \sim 1500$ $M_{\text{Sun}}$ $\Rightarrow$ stable as $M_c<M_J$.  
  - **Molecular cloud core:** $M_J \sim 8$ $M_{\text{sun}}$ $\Rightarrow$ unstable as $M_c>M_J$.

- **So deep inside molecular clouds the cores are collapsing to form stars.**
Time-Scale for Collapse

• If Jeans criteria for gravitational collapse has been satisfied, the collapsing molecular cloud is essentially in free-fall during the first part of its evolution; any existing pressure gradients are too small to influence the motion.

• Throughout the free-fall phase, the temperature of the gas remain nearly constant (i.e., the collapse is said to be isothermal).

    ---- Because as long as the cloud remains optically thin and the gravitational potential energy released during the collapse can be efficiently radiated away.

• As long as the original density of the spherical cloud was uniform, all parts of the cloud will take the same amount of time to collapse and density will increase at the same rate everywhere (Homologous collapse).
Time-Scale for Collapse

• The collapse time-scale $t_{ff}$ when $M_r > M_J$ is given by the time a mass element at the cloud surface needs to reach the centre.

• In free-fall, an mass element is subject to acceleration $g = \frac{GM_r}{R^2}$

• The time to cover a distance $R$ can therefore be estimated from:

$$R = \frac{1}{2} gt_{ff}^2 = \frac{1}{2} \frac{GM_r}{R^2} t_{ff}^2$$

• By approximating $R$ using $R^3 \sim M/\rho \Rightarrow t_{ff} \approx (G \rho)^{-1/2}$

• Higher density at cloud center $\Rightarrow$ faster collapse.

• For typical molecular cloud, $t_{ff} \sim 10^3$ years (i.e. very short).

$t_{ff} \sim 27$ minutes for our Sun
Fragmentations

- *If cloud cools enough, Jeans instability allows gravity to overtake thermal energy. The densest parts of the cloud become gravitationally unstable and fragmentation occurs.*
• Since the Jeans mass is a minimum mass, the unstable collapsing cloud could begin with a very large mass, may be even thousands of solar masses. Why, then, don’t we get such massive stars? There are a number of reasons, but one is that as the density increases hugely.

\[
M_J = \left( \frac{375k^3}{4\pi\mu^3 m_H^3 G^3} \right)^{1/2} \left( \frac{T^3}{\rho} \right)^{1/2}
\]

• Thus the Jeans mass becomes smaller, and so large clouds will fragment. Final fragments will tend to be on the order of a solar mass. Thus, an initial massive collapsing cloud will fragment into many small collapsing protostars ⇒ star cluster.

• Why does the temperature remain approximately constant? With such a low initial density(≈ 3x10^{-24}g/cm³), the opacity is low, and the cloud can easily radiate away its thermal in the form of IR photons.
Why doesn’t the fragmentation continue to very small masses? That is, why do stars form at all?

• The isothermal collapse cannot continue forever. As the fragment becomes more dense, the opacity rises, and the collapse goes from being isothermal to adiabatic. Beginning with the adiabatic relationship:

\[ P = K \rho^\gamma \]

• Substituting into the ideal gas law, we can show that for an adiabatic collapse,

\[ T \propto \rho^{\gamma-1} \]

i.e., the temperature rises in an adiabatic collapse. This means that

\[ M_J \propto \rho^{(3\gamma-4)/2} \propto \rho^{1/2} \]

for atomic hydrogen, \( \gamma = 5/3 \)

and this means that the Jeans mass begins to increase, stopping the fragmentation. As \( M_J > M_C \) locally.
Cloud collapse

Protostar contraction

• The increased pressure that occurs when the collapse becomes more adiabatic slows the rate of collapse near the core. At this point the central region is nearly in hydrostatic equilibrium.

• Eventually, the density in the core becomes so great significantly opacity develops. This slows the collapse, and the core begins to heat. When the temperature reaches about 2000K, molecular dissociation occurs, and the molecular H dissociates into atomic hydrogen.
Cloud collapse

Hertzsprung-Russell diagram

This shows the progression from a cold nebula (maybe $T = 30$ K) with a total mass of about 1 solar mass into a star like our Sun.

<table>
<thead>
<tr>
<th>Class</th>
<th>Temperature (°K)</th>
<th>Star Color</th>
</tr>
</thead>
<tbody>
<tr>
<td>O</td>
<td>30,000 - 60,000</td>
<td>Blue</td>
</tr>
<tr>
<td>B</td>
<td>10,000 - 30,000</td>
<td>Blue</td>
</tr>
<tr>
<td>A</td>
<td>7,500 - 10,000</td>
<td>White</td>
</tr>
<tr>
<td>F</td>
<td>6,000 - 7,500</td>
<td>White (yellowish)</td>
</tr>
<tr>
<td>G</td>
<td>5,000 - 6,000</td>
<td>Yellow (like the Sun)</td>
</tr>
<tr>
<td>K</td>
<td>3,500 - 5,000</td>
<td>Orange</td>
</tr>
<tr>
<td>M</td>
<td>2,000 - 3,500</td>
<td>Red</td>
</tr>
</tbody>
</table>

Spectral classification includes 7 main types: **O, B, A, F, G, K, M**.
How long does it take a protostar to "reach" the main-sequence, i.e. start nuclear fusion? It depends on its mass:
The Hayashi Limit:

- Hayashi discovered that for a star of a given mass and chemical composition, at each luminosity there exists a maximum radius (and a minimum temperature) at which a star becomes fully convective. This defines, for a given star, a nearly vertical line in the H-R diagram called Hayashi limit.

- To the right of the Hayashi limit, stars are dynamically unstable (not in hydrostatic equilibrium) and evolve on dynamical timescale. To the left of the Hayashi limit, stars can achieve hydrostatic equilibrium, and thus evolve more slowly. On the Hayashi limit, stars are completely convective in their interiors.
**Cloud collapse**

**Protostar contraction**

- The star would slowly contract, radiating away all of its gravitational potential energy. If it does so at its present luminosity, the time it takes to radiate away its energy is given by the total energy divided by the rate the energy is lost (which is the luminosity). This time is Kelvin – Helmholtz (K-H) timescale.

- The region to the right (cool side) of the Hayashi line is the so-called forbidden zone, where there are no equilibrium solutions to the equations of stellar structure. The few objects which are observed to occur in the forbidden zone are assumed to be either: collapsing protostars or other objects not in hydrostatic equilibrium.
Cloud collapse

1. Cloud exceeding $M_J \Rightarrow$ collapse
2. first: Optically thin $\Rightarrow$ isothermal collapse
3. Core density reaches $10^{-13}\text{cm}^{-3} \Rightarrow$ optically thick $\Rightarrow$ Adiabatic collapse, pressure slows down the collapse at the core.
4. Above core: material in supersonic free-fall $\Rightarrow$ shock front $\Rightarrow E_{\text{kin}} \Rightarrow E_{\text{ther}}$
5. When $T > 1000\text{K}$: dust evaporates $\Rightarrow$ opacity drops $\Rightarrow L$ drops
6. When $T > 2000\text{K}$: $\text{H}_2$ dissociates (absorbs energy) $\Rightarrow$ second collapse

A theoretical evolutionary track of the gravitational collapse of a $1\,M_\odot$ cloud.
- With the high H- opacity near the surface, the star is completely convective during 1 million of years of collapse (about point 2).
- During this early period of collapse, the first stage of nuclear burning occurs in the star’s center (about point 1). The reaction involved is deuterium burning.
- The radiative core allows energy to escape into the convective envelope more readily, causing the luminosity of the star to increase again (about 3).
- Rate of nuclear E becomes great at the core (about point 5)

Pre-main sequence evolutionary tracks of stars with various masses.
Discussions

- A cold, dense molecular cloud begins to collapse because gravity overcomes thermal pressure.
- Jeans mass is the minimum or critical mass that is necessary to initiate the spontaneous collapse of the cloud.
- During the initial parts of its contraction, a cloud won't heat up -- radiation escapes and carries away the energy. The Jeans mass decreases because T is staying constant but the density is increasing. Eventually the cloud breaks into fragments which in turn collapse to become stars.
  - Hayashi line: In the Hertzsprung-Russell diagram the (nearly vertical) transition line along which the star establishes hydrostatic equilibrium.
- Protostars in the process of contracting towards the ZAMS, prior to the development of a radiative core and the initiation of hydrogen fusion, are fully convective. The evolutionary track of the contracting star follows the Hayashi line in the beginning, until it develops a growing core in radiative equilibrium when it moves to the left (higher $T_e$) and finally reaches the ZAMS with the initiation of hydrogen burning.
References:-

• An introduction to modern astrophysics, by Bradley W. Carroll & Dale A. Ostlie, 1996

• Stellar structure and Evolution, by R. Kippenhahn & A. Weigert, 1990


• The formation of stars, Stahler and Palla,, Wiley-VCH, 2004
Origin of solar systems
30 June - 2 July 2009
by Klaus Jockers (jockers@mps.mpg.de)
Max-Planck-Institut of Solar System Science
Katlenburg-Lindau

Part 2b
Cloud collapse (numerical calculations), initial mass function, summary
Numerical calculations of cloud collapse

Matthew R. Bate, Ian A. Bonnell and Volker Bromm, 2003
The formation of a star cluster: predicting the properties of stars and brown dwarfs
MNRAS 339, 577-599.

Matthew R. Bate, 2009
Stellar, brown dwarf and multiple star properties from hydrodynamical simulations of
star cluster formation
MNRAS 392, 590-616.

If mean atomic mass = 2.46 amu = 4.084 $10^{-24}$g, then the number density of $10^4$ cm$^{-3}$
(clouds, clumps) corresponds to a mass density of $4.084 \times 10^{-20}$g cm$^{-3}$

Aim of the calculations: Not only to visualize the condensation process but also to
determine and to understand the initial mass function, i. e. the mass distribution of
the forming protostars, how many of them are multiple systems, ect.
The opacity limit for fragmentation:

As long as the gravitational energy gained by contraction can be radiated away, the polytropic index \( \gamma = \frac{d \log[p]}{d \log[\rho]} \approx 1 \). This allows the possibility of fragmentation because the Jeans mass decreases with increasing density if \( \gamma < 4/3 \).

\[
T \propto n^{\gamma - 1} \\
M_J \propto T^{3/2} \cdot n^{-1/2} \propto n^{3/2}(\gamma - 1)^{-1/2}
\]

for \( \gamma = 1 \), \( M_J \propto n^{-1/2} \)

If the gravitational energy gained by contraction exceeds the rate that can be radiated away, the gas heats up with \( \gamma > 4/3 \), the Jeans mass increases and the unstable clump quickly becomes stable.

\[
\text{for } \gamma = \frac{4}{3} \quad M_J \propto n^0
\]

For an initial temperature \( T = 10 \) K the critical density \( \rho_{\text{crit}} \approx 10^{-13} \) g cm\(^{-3} \).

Minimum “stellar” mass \( \approx 10 \) \( M_J \) and the minimum separation between stars = 10AU (Size of pressure supported fragment).
Computational method:

3d Smoothed Particle Hydrodynamics (SPH), originally developed by Benz. Parallelized using OpenMP.

Equation of state:

To model the opacity limit for fragmentation, discussed in Section 2, without performing radiative transfer, we use a barotropic equation of state for the thermal pressure of the gas \( p = K \rho^n \), where \( K \) is a measure of the entropy of the gas. The value of the effective polytropic exponent \( \eta \), varies with density as

\[
\eta = \begin{cases} 
1, & \rho \leq 10^{-13} \text{ g cm}^{-3}, \\
7/5, & \rho > 10^{-13} \text{ g cm}^{-3}.
\end{cases}
\]

We take the mean molecular weight of the gas to be \( \mu = 2.46 \). The value of \( K \) is defined such that when the gas is isothermal \( K = c_s^2 \), with the sound speed \( c_s = 1.84 \times 10^4 \text{ cm s}^{-1} \) at 10 K, and the pressure is continuous when the value of \( \eta \) changes.
Figure 1. Comparison of our barotropic equation of state (dotted line) with the temperature–density relation during the spherically-symmetric collapse of a molecular cloud core as calculated with frequency-dependent radiative transfer (solid line Masunaga & Inutsuka 2000). The curves differ for densities less than $10^{-14} \text{ g cm}^{-3}$ simply because Masunaga & Inutsuka chose parameters such that their initial core had a temperature of 5 K rather than our assumption of 10 K. However, in the non-isothermal regime, from $10^{-13}$ to $10^{-8}$, our parametrization matches the radiative transfer result to an accuracy of better than 20 per cent. The second collapse (discussed in Section 2) occurs from densities of $\approx 5 \times 10^{-8}$ to $\approx 3 \times 10^{-3}$ and is not modelled.
“Sink” particles

Sink particles must be introduced into the numerical code to provide a lower limit of the scale length.

If $\rho > 1000 \rho_{\text{crit}}$, a sink particle is inserted. It replaces the SPH particles contained within $r_{\text{acc}} = 5\text{AU}$ by a point mass with the same mass and momentum.

Sink particles interact with the gas only via gravity and accretion.

All stars and brown dwarfs start as sink particles.

Gravitation between sink particles is Newtonian but softened if the particles approach each other by less than 4 AU. Maximum acceleration occurs when the distance = 1AU (minimum separation of components of double stars), but part of the calculation was redone without this softening.

Sink particles merge when they pass within 0.02 AU from each other. (23 mergers within the whole run).
Multiple stellar systems

Multiple stellar systems are determined after the run by constructing a structure tree.

Some of the binaries turn out to be very wide (several 1000 AU). They consist of ejected objects that happen to have nearly the same velocity.
Initial conditions:

A 500 M☉ molecular cloud with radius 0.404 pc = 83300 AU.
At a temperature of 10 K the mean thermal Jeans mass is 1 M☉.

+ **supersonic turbulent velocity field:**

Initially the kinetic energy of the turbulence equals the magnitude of the gravitational potential energy of the cloud, i.e. the cloud has enough turbulent energy to support itself against gravity.

The initial rms Mach number of the turbulence = 13.7.

At 10 K the sound speed is 184 m s⁻¹, i.e. the mean turbulent speed = 2.52 km s⁻¹. This unrealistically high value is necessitated by the large size of the cloud (see next projection).

Supersonic velocity field is an essential ingredient in the stability of a molecular cloud.
Resolution:

Minimum Jeans mass must be resolved. At $\rho_{\text{crit}} = 10^{-13}$ g cm$^{-3}$ it is 0.0011 $M_\odot$.

This requires $3.5 \times 10^7$ smoothed model particles.

Total computing time $10^5$ CPU hours ($\sim$4000 days) on a 1.65GHz IBM p570 computer node.

$>459$ stars and $<795$ brown dwarfs formed, total mass 191 $M_\odot$.
I.e. 38% of the cloud was transformed into stars.

The movie, produced by M. Bate and coworkers, Exeter, UK, can be found at http://www.astro.ex.ac.uk/people/mbate/Research/Cluster/cluster3d.html
Figure 3. Histograms giving the IMF of the 1254 stars and brown dwarfs that had been produced by the end of the main calculation. The single-hashed region gives all objects, while the double-hashed region gives those objects that have stopped accreting. Parametrizations of the observed IMF by Salpeter (1955), Kroupa (2001) and Chabrier (2003) are given by the magenta line, red broken power law and black curve, respectively. The numerical IMF broadly follows the form of the observed IMF, with a Salpeter-like slope above $\sim 0.5 \, M_\odot$ and a turnover at low masses. However, it clearly overproduces brown dwarfs by a factor of $\approx 4$. 


Summary

A molecular cloud becomes unstable to collapse simply because in a homogeneous gas cloud with constant density and pressure, gravitational energy rises faster than volume, while thermal energy is proportional to volume, i.e. if one increases the size of a homogeneous cloud a point of collapse will be reached.

Instability increases with increasing mass of the cloud and with decreasing temperature.

An important issue are the “quasi-random” velocities in a molecular cloud. Numerical models of cloud collapse assume a random velocity field of large enough velocities to stabilize the cloud initially. As the temperature is very low these velocities are supersonic (larger than the thermal velocity). If the cloud increases in size, these velocities must increase to unrealistic levels (because the gravitational energy in the cloud increases too rapidly).

Numerical models allow to calculate the initial mass function in a collapsing cloud, but there are theoretical and observational limits to an accurate determination of this initial mass function.
EARLY PHASES OF PROTOSTARS:
Star formation and Protoplanetary Disks

How do we observe them?
Formation of stellar systems: 1. Introduction

- Observation of stars at different stages of their life has enabled us to retrace the processes of their formation and of their different stages in their evolution.

- By studying them we can constrain theories of stellar and planetary formation and can develop timescales for the evolution of planetary development.

- And finally compare our own solar system to others.
• **Nebular hypothesis:** Most widely accepted model explaining the formation and evolution of the solar system. This hypothesis of planetary formation is also thought to occur throughout the Universe.

“(...) the solar system condensed from a large rotating nebula.”

*Immanuel Kant, 1755*[^1]

“(...) the planets have formed from gas rings ejected from the equator of the collapsing Sun.”

*Laplace, 1796*[^1]

• It is now believed that our sun, like other stars, formed from dense interstellar clouds, composed mainly of molecular hydrogen (H₂).

Dark clouds in the star forming region IC 2944 from HST's WFPC2.
Formation of stellar systems: Molecular clouds

<table>
<thead>
<tr>
<th>Dense molecular clouds</th>
</tr>
</thead>
</table>

<table>
<thead>
<tr>
<th>Temperature</th>
<th>Density</th>
<th>state of hydrogen</th>
<th>Size</th>
<th>Mass</th>
</tr>
</thead>
<tbody>
<tr>
<td>10 -20 K</td>
<td>&gt; 300 H cm$^{-3}$</td>
<td>molecular (H$_2$)</td>
<td>10-50 pc</td>
<td>~ 100 000 M$_{\odot}$</td>
</tr>
</tbody>
</table>

They are the densest and coldest forms of interstellar medium and form 50% of the overall mass of the interstellar medium.

**Where to look if we don't see through them in the visible part of the spectrum?**

Stellar forming region DR21 in the constellation of Cygnus. Image taken from NASA's Spitzer Space Telescope.

- The **infrared** image reveals the presence of dust.
- Dust grains absorb UV light from the surrounding stars. It heats up and re-emits in the infrared.
How do we know that young stars are related to the molecular clouds?

1) We see massive, luminous stars which cannot be older than about $10^6$ yr.

2) In association with these we see often peculiar stars with emission lines (TTauri stars), thus assumed to be also young.

3) TT stars have lower luminosities and are redder than the massive O, B stars but are still considerably more luminous than MS stars of the same color.

4) Lower luminosities $\rightarrow$ lower masses than O, B stars $\rightarrow$ contraction times longer

5) SO, if TT stars were formed at the same time as their more massive associates ($\sim 10^6$ yr) they have not had enough time to contract to the MS $\rightarrow$ still contracting! ($t_{\text{cl}} \propto M_c^{-0.5}$)

6) Both O, B and TT stars appear in association with large dust complexes.
**Cloud collapse:** Observations show that new stars can be formed in an environment of dense interstellar clouds. Under certain circumstances these clouds can become gravitationally unstable to contraction.

In equilibrium:

\[-E_g = 2E_k\]

\[3 \frac{GM^2}{5R} = \frac{3}{2} kT \frac{M}{m}\]

During collapse:

\[-\frac{1}{2}E_g > E_k\]

\[M > M_J = 9 \cdot 10^4 M_\odot \left(\frac{T^3(K)}{n(m^{-3})}\right)^{\frac{1}{2}}\]

For a MC with:

\[T=10K \quad n=100 \text{ cm}^{-3}\]

\[M_J \sim 10^2 M_\odot\]

To form a star it is not strictly necessary to have such massive clouds. There are **inhomogeneities** that will cause the cloud to **fragment** leading to the formation of more than one star.
Formation of stellar systems: Cloud collapse (stages)

1) Isothermal phase

2) Adiabatic phase (Class 0)

3) Protostar (Class I)

4) “Classic” T Tauri phase (Class II)

5) “Weak” T Tauri phase (Class III)

Ph. André et. al. 2002, EAS publication series, vol. 2
Formation of stellar systems: Cloud collapse (Phase 1)

Isothermal phase

What is going on?

- $T \sim 10K$
- $\rho$ is tenuous enough for gravitational energy to dissipate through the radiation coming from the thermal excitation of the atoms.
- As a consequence of this, the temperature remains low and it keeps contracting.

Where can we observe it?

Object observable through its infrared thermal emission.

far IR – sub mm radiation
Formation of stellar systems: Cloud collapse (Phase 2)

Adiabatic phase

What is going on?

- $\rho$ increases $\Rightarrow$ $k$ increases
- $k$ reaches the point where the energy released by the contraction cannot escape by radiation: opaque
- As a consequence of this, the temperature rises until contraction stops because of pressure built up:
  Hydrostatic equilibrium

Where can we observe it?

Only the cloud is detectable from radiation from the dust as a black body that peaks in the far infrared $\Rightarrow$ SED's

 PROTOSTAR!! (class 0)
Formation of stellar systems: SEDs

★ What can we say is going on in a given physical system from its spectral energy distribution (SED)?

Central star surrounded by a disk of gas and dust

For every distance from the star the dust will be at a different and unique temperature: each radius emits at a characteristic peak wavelength.

The infrared excess seen is due to the gas and dust in the disk and envelope.

Modeling the SED allows to compare observations with what might be going on.
Formation of stellar systems: SEDs (Multiwavelength Observations)

- **Near-infrared bump**: the inner rim, the infrared dust features from the warm surface layer, and the underlying continuum from the deeper (cooler) disk regions.

- **Near- and mid-infrared**: from small radii.

- **Far-infrared**: from the outer disk regions.

(sub-)mm. emission mostly comes from the mid-plane of the outer disk. This flux probes also the disk mass.

Differences in disk geometry are mainly reflected in the energetic domain.

(Dullemond et al. 2007)

At short λ's SED is in the "Wien domain":

\[ B_\lambda(T) \approx \frac{2hc^2}{\lambda^5} e^{-\frac{hc}{\lambda kT}} \]

At long λ's the SED is in the "Rayleigh-Jeans domain":

\[ B_\nu(T) \approx \frac{2k\nu^2}{c^2} T \]
Formation of stellar systems: Disk structure

**Disk structure** *(Dullemond et al. 2007)*

1) Optically THIN disk:
- Silicate emission from small (<6μm), warm (150-450K) grains.

2) Optically THICK disk:
- IR excess from outer most grains.

**Gas accretion onto the stellar surface:**
- Hα emission lines
- Widths > 200km/s
- Assymetries: e.g. blueshifted absorption

**Inner gaseous disk (mostly H₂)**
- Typically at r<0.05 AU

**Layer of cold, larger (<1mm) dust grains**
- Submm continuum emission

**Dust sublimation due to high temperatures (T>1500K)**

**Central protostar**

**Inner dust rim**

**Flared disk**

**Silicate emission at 9.7μm**
*(IRS specrum: Sicilia et al. 2007)*

**UV excess**
- Hα emission
- 11 - 2146 K6

**IR excess**
Formation of stellar systems: SEDs (modeling)

- **Temperature distribution of the disk**

  \[ T(R) = T_s \left( \frac{R}{R_*} \right)^{-p} \]

  R: distance from star  
  R*: Radius of star  
  T_s: Normalization factor  
  p: power \((0 < p \leq 2)\)

  **Material, size and shape** of the dust particles affects the temperature distribution and thus the shape of the SED.

- **Characteristics of a SED**

  Slope (value of \(p\)):
  - flatter slope --> small \(p\) --> small grains  
  - steep slope --> big \(p\) --> big grains

  Shape:
  - \(\lambda_{\text{max}}\) --> dust dominated (longer \(\lambda\)'s)  
  - star dominated (shorter \(\lambda\)'s)  
  - dips --> gaps in the disk --> planets?

  \((p=0.75)\)
Formation of stellar systems: SEDs (Classes)

3 main classes of young stellar objects (YSO) can be distinguished based on the slope of their SEDs (André et. al. 1994):

\[ \alpha_{IR} = \frac{d\log(\lambda F_{\lambda})}{d\log(\lambda)} \quad \text{for} \quad \lambda = [2.2, 10-25] \mu m \]

\[ \alpha_{IR} > 0 \quad \text{(Class I)} \]
\[ -1.5 < \alpha_{IR} < 0 \quad \text{(Class II)} \]
\[ \alpha_{IR} < -1.5 \quad \text{(Class III)} \]

related to the amount of dust in the disk + envelope

NGC 6240 (www.ipac.caltech.edu)

HD 23514 in the Pleiades star cluster (http://www.gemini.edu)
Formation of stellar systems: 4. Class 0 vs. Class I sources

**CLASS 0**
- Associated with formed (hydrostatic) YSOs
- Visible at $\lambda \geq 25\mu$m
- Indirect evidence for a central YSO:
  - cm radio continuum emission
  - Presence of a collimated CO outflow
- Presence of a dust envelope in sub-mm.
- $L_{\text{submm}} / L_{\text{bol}} \gg 0.5\% \rightarrow M_{\text{env}} \gg M_*$ (André 2003)
- $\dot{M}_{\text{jet}} \sim 10^{-6} M_\odot /\text{yr}$ (Bontemps et al. 1996)
- $\dot{M}_{\text{acc}} \sim 10^{-5} M_\odot /\text{yr}$

![HH 212 outflows in SiO(2-1) (3.4mm) (Codella et al. 2007)](image1)

**CLASS I**
- Associated with late active mass accretion phase
- Rising SEDs at $\lambda \geq 2\mu$m
- Less collimated CO outflows
- $L_{\text{submm}} / L_{\text{bol}} \sim 0.5\% \rightarrow M_{\text{env}} \leq M_*$
- $\dot{M}_{\text{jet}} \sim 2 \times 10^{-8} M_\odot /\text{yr}$ (Bontemps et al. 1996)
- $\dot{M}_{\text{acc}} \sim 2 \times 10^{-7} M_\odot /\text{yr}$

![TMC-1 at 1.6 μm (i=40° – 70°) (Terebey et al. 1993)](image2)
Formation of stellar systems: 4. Class 0 (observations)

OBSERVED CLASS 0 SOURCES:

IRAS 04191
(Belloche et al. 2001)

VLA 1623
(André et al. 1993)

- Dots are the observational data.
- Solid curve is a fit of a 20K blackbody.
- Dashed curve is a “best-fit” resulting from a modeling with a radiative transfer code.
Formation of stellar systems: 4. Class I (observations)

**OBSERVED CLASS I SOURCES:**

- Dots are observational data.
- Dashed line is a greybody model.
- Other lines are examples of circumstellar envelope fits.

**CIRCUMSTELLAR ENVELOPE MODELS FOR EL 29**

<table>
<thead>
<tr>
<th>Model</th>
<th>(p)</th>
<th>(R_{\alpha}) (AU)</th>
<th>(T_{\text{in}}) (K)</th>
<th>(\langle T_{\text{dust}}\rangle) (K)</th>
<th>(M_{\text{env}}) ((M_\odot))</th>
<th>(A_V)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>0.0</td>
<td>0.1</td>
<td>1600</td>
<td>35.1</td>
<td>0.10</td>
<td>~14</td>
</tr>
<tr>
<td>2</td>
<td>0.5</td>
<td>1</td>
<td>630</td>
<td>35.9</td>
<td>0.09</td>
<td>~20</td>
</tr>
<tr>
<td>3</td>
<td>1.0</td>
<td>300</td>
<td>60</td>
<td>34.2</td>
<td>0.12</td>
<td>~21</td>
</tr>
<tr>
<td>4</td>
<td>1.5</td>
<td>300</td>
<td>60</td>
<td>35.6</td>
<td>0.09</td>
<td>~22</td>
</tr>
<tr>
<td>5</td>
<td>2.0</td>
<td>300</td>
<td>55</td>
<td>34.3</td>
<td>0.10</td>
<td>~31</td>
</tr>
<tr>
<td>6*</td>
<td>Graybody</td>
<td>...</td>
<td>...</td>
<td>40.0</td>
<td>0.08</td>
<td>~13</td>
</tr>
</tbody>
</table>

Taurus star forming region (Padgett et al. 1998)

IRAS 04302+2247 F160W

HST/NICMOS at 1.6μm broadband photometry.
Formation of stellar systems: 4. Accretion and outflows

**ACCRETION**

- Inward motions can be traced by optically thick molecular lines which exhibit asymmetric double-peaked profiles with a stronger blue peak:
  - CS (2-1) transition (~ 3mm) 
    (Belloche et al. 2002)
  - CO (2-0) transition (~2.3 μm) 
    (Antoniucci et al. 2007)

- From thermal fluxes at submm and far-IR wavelengths:
  \[ M_{\text{env}} \]

- It's also possible to retrieve the mass accretion rates through the \( L_{\text{acc}} \) (related to the HI Brγ flux (Antoniucci et al. 2007) ) using the formula:
  \[ \dot{M}_{\text{acc}} \sim \frac{L_{\text{acc}} R_\star}{GM_\star} \]

**OUTFLOWS**

Study of 36 near-IR (class I) and 9 far-IR/submm (class 0) protostars using maps of CO(2-1) line. 
(Bontemps et al. 1996)

From mm transitions of CO(2-1) (1.2mm): 
(Belloche et al. 2002)

\[ \rightarrow \text{outflow activity from jets} \]
\[ \rightarrow \text{difference in classes (evolution)} \]
One expects (independently of details of protostellar theory) YSO to become warmer and to be surrounded by progressively smaller amounts of circumstellar material as they evolve.

- Evolutionary indicators:
  - Peak of the SED $\rightarrow \lambda_{\text{max}}$
  - Decrease of circumstellar mass $\rightarrow M_{\text{env}}$
  - Accretion mass rates $\rightarrow \dot{M}_{\text{acc}}$
  - Jets, outflow mass rates $\rightarrow \dot{M}_{\text{jet}}$
Spectral energy distribution (SED) or Radiative flux as a function of wavelength ($F_\lambda$)

**Blackbody**: Theoretical object that is a perfect emitter and absorber of radiation. Its energy spectrum depends only on the temperature of the body.

$$B_\lambda (T) = \frac{2 \hbar c^2}{\lambda^5} \frac{1}{e^{\frac{\hbar c}{\lambda kT}} - 1}$$

**Wien's law**: The temperature of a blackbody is inversely proportional to the wavelength peak of emission.

$$\lambda_m = 2.89 \times 10^{-3} / T (k)$$

No object is a perfect blackbody, but it is still possible to approximate most objects as blackbodies. Stars act almost exactly like a blackbody.
Formation of stellar systems: SEDs (Multiwavelength Observations)

Optically thick disks:

As soon as the grains absorb the radiation from the central star and emit their own, it immediately is absorbed again by the surrounding grains throughout the entire disk. Only the radiation emitted by the outermost grains in the disk can be observed so the material of the grains and the density of the disk do not play a large role in the SED. The size of the grains is the dominant factor.

Dust grains that are larger than the wavelength of incoming radiation absorb that radiation very efficiently, i.e., grains that are larger than the peak wavelength they emit, emit that radiation effectively, they tend to cool off closer to the star.

Grains smaller than the wavelength of incoming radiation do not absorb that radiation very well, i.e., do not efficiently emit radiation at wavelengths larger than their diameter, they will remain hotter at larger radii.

Optically thin disks:

The energy emitted from the grains within the disk is not immediately absorbed again by the surrounding grains. Energy emission from grains throughout the disk, not just the ones present at the surface is observed so the density of the disk and the material of the disk play a larger role.

The longer a wavelength, the harder it is for grains to emit that wavelength of radiation (does not take effect until considering wavelengths longer than 100 μm).
Formation of stellar systems: Evolution

**Pre-Stellar Dense Core**
- $T_{\text{bol}} \sim 10-20$ K, $M = 0$
- $\sim 10^6$ yr

**Class I**
- $T_{\text{bol}} < 70$ K, $M_* \ll M_{\text{env}}$
- $< 30,000$ yr

**Evolved Accreting Protostar**
- $T_{\text{bol}} \sim 70-650$ K, $M_* > M_{\text{env}}$
- $\sim 200,000$ yr

**Class II**
- $T_{\text{bol}} \sim 650-2880$ K, $M_{\text{Disk}} \sim 0.01 M_\odot$
- $\sim 10^8$ yr

**Class III**
- $T_{\text{bol}} > 2880$ K, $M_{\text{Disk}} < M_{\text{Jupiter}}$
- $\sim 10^9$ yr

**Debris + Planets?**

**Time**
The double peaked profiles show two outflow patterns. From the fit of synthetic spectra one can retrieve the velocities.

Spectra along the axis perpendicular to the outflow in the optically thick CS(1-2) line.
Optical interferometers are ideally suited to directly probe the innermost regions of the circumstellar environment around young stars.

Observational phase-space (spectral domain and angular resolution) for optical interferometers, and for complementary techniques (shaded polygons). Also outlined over the most relevant phase-space regions (rectangular boxes) are the main physical phenomena associated with young stellar objects. (Millan-Gabet et al. 2007)
Gas in the central part of the nebula undergoes fast compression and forms a hot hydrostatic (not contracting) core containing a small fraction of the mass of the original nebula. The gas of a circumstellar disk accretes onto the core. The core gradually grows in mass until it becomes a young hot protostar. At this stage, the protostar and its disk are obscured by the infalling envelope and are not directly observable. Such objects are observed as very bright condensations, which emit mainly millimeter-wave and submillimeter-wave radiation - Class 0 protostars.

**PROTOSTELLAR ENVELOPES:** Class 0 source Barnard 1c in Perseus

- early protostellar collapse stage
- SEDs that resemble blackbodies with $T \leq 30$ K
- majority of the source mass resides in the infalling envelope
- exhibit powerful, bipolar molecular flows
The extinction map of Perseus shows a chain of dark clouds. All of the known dark clouds and star-forming regions are seen with highest extinction.

The morphology of the integrated CO intensity in Perseus is similar to that of the extinction, but reveal complex substructure within the clumps seen in extinction.

$^{13}$CO emission in Perseus, overlaid with the positions of the dense cores detected in submm continuum emission (red circles). Symbol size is proportional to the mass of the core. (Ridge et al. 2006)
PROTOTOSTELLAR ENVELOPES: Class 0 source Barnard 1c (B1-c) in Perseus

Barnard 1c is a potential star-forming in the main molecular core of Barnard 1 in the Perseus cloud complex.

Three-color image of the Perseus region with 3.6 μm (blue), 4.5 μm (green), and 8.0 μm (red). (Jørgensen et al. 2006)

The SED resembles a blackbody with $T \leq 30$ K. At this stage, the protostar and its disk are obscured by the infalling envelope.

SED of the B1-c class 0 object. (original figure from Jørgensen et al. 2006)
The most strongly red-shifted (gray contours) and blue-shifted (black contours) emission lies close to the protostar.

The material at the leading edge of the blue-shifted lobe appears red-shifted relative to the velocity of the source.

**CO 1–0 emission over a 4.5 μm image (greyscale).** (Jørgensen et al. 2006) The cross marks the position of the 3 mm continuum peak. (Matthews et al. 2008)

Schematic of the morphology of B1-c: The centrally heated cavity and outflow cavity are indicated by absence of N₂H⁺. (Matthews et al. 2008)
A central cavity can be identified in N$_2$H$^+$ emission. Its anticorrelation with C$^{18}$O emission suggests that heating in the center has released CO from grain mantles, in turn destroying N$_2$H$^+$.

The declining emissivity of grains and the lower fluxes of dust emission at longer wavelengths makes 850 $\mu$m measurements superior to 1.3 mm observations.

Continuum emission at 3 mm (gray contours), 1 mm (white contours) and 850 $\mu$m (grayscale). (Matthews et al. 2008)
As the envelope's material continues to infall onto the disk, it eventually becomes thin and transparent and the YSO becomes observable (far-IR, visible).

The protostar begins to fuse deuterium and then ordinary hydrogen.

The external appearance of the YSO at this stage corresponds to the spectral class I protostars (YTTS). The forming star has already accreted much of its mass (80-90% of the system).

- later stage of protostellar collapse
- displays very broad SEDs that peak near 100 μm
- envelope masses are similar to the mass of the central pre-main-sequence core
- well-developed accretion disks
- envelopes have bipolar cavities excavated by outflows
The Butterfly star is a class I protostar in the Taurus-Auriga molecular cloud complex whose equatorial plane is inclined edge-on to the LOS (incl. ~90°).

The near-IR appearance is dominated by the totally opaque band extending 900 AU north/south that bisects the scattered light nebulosity.

HST 1.6 μm surface photometry. Contours double in flux with each level. (Padgett et al. 1999)

HST pseudo-true color composite image of 1.1 μm, 1.6 μm, and 2.05 μm observations. (Padgett et al. 1999)

The dark lane coincides with a dense rotating disk of molecular gas and may therefore be a large optically thick circumstellar disk seen precisely edge-on.
Mm-mapping of this source in $^{13}$CO 1-0 indicates that the dark lane coincides with a dense rotating disk of molecular gas.

The dust lane may therefore be a large optically thick circumstellar disk seen precisely edge-on.

Near-IR and mm wavelength images show that the grains in the envelope of this object cannot be distinguished from those of the ISM and that the grains the much denser circumstellar disk have grown via coagulation by up to 2–3 orders of magnitude.

The separated dust grain evolution is in agreement with the theoretical prediction of a sensitive dependence of grain growth on the location in the circumstellar environment of young (proto)stars:

Grain growth is expected to occur on much shorter timescales in the dense region of circumstellar disks than in the thin circumstellar envelope. For the same reason a radial dependence of the dust grain evolution in the disk itself is expected. (Wolf et al. 2003b)
The system is assumed to consist of a circumstellar disk and an infalling envelope.

The circumstellar disk is assumed to be responsible for both the dark lane in the optical/IR wavelength range and the mm structure of the object. An additional envelope is required to explain the extended scattered light structure.

Although an infalling envelope, seen perpendicular to their axis of rotation symmetry, may also create a dark lane that hides the central star at optical and IR wavelengths, the observed sharp transition of the dark lane toward the scattered light regions, points to the presence of a disk. (Wolf et al. 2003a)

HST near-IR images. Assuming a distance of 140 pc, the side length of the images is 900 AU. Contours mark steps of 0.5". (Wolf et al. 2003a)
The result of simulations shows a very good agreement with the observed SED in case ofassuming an envelope+disk system, with the envelope being extended to a maximum distance of 450 AU from the star.

Comparing the dust reemission SED of the disk alone with the reemission of the whole system (disk +envelope), one finds that the SED is dominated by the reemission from the envelope up to a wavelength of about 174 µm. (Wolf et al. 2003)
As the envelope completely disappears, the protostar becomes a classical T Tauri star. The mass of the disk around a classical T Tauri star is about 1–3% of the stellar mass. A pair of bipolar jets is usually present as well. Strong emission lines form as the accreted gas hits the "surface" of the star, which happens around its magnetic poles.

- characterized by the presence of excess infrared emission above that expected for a stellar photosphere
- SED emission peak occurring in the near-IR
- dispersal of the remnant infall envelope by the combined effects of infall and outflow
HV Tau A, HV Tau B, and HV Tau C were revealed to be a triple system with an edge-on disk inclined by 85° around HV Tau C.

They are located in the Taurus molecular cloud (140 pc) and are two of the brightest young stellar object with edge-on disks.

HV Tau C has 10 μm flux excess relative to the estimated spectrum of the central star, while HV Tau A, HV Tau B shows no infrared excess.
HV Tau C exhibits scattered light in an edge-on disk, and its two components are clearly resolved in all the bands.

The width of the dark lane of HV Tau C is estimated by the distance between the peak of the two components of the scattered light from HV Tau C. The width is wider at the shorter wavelengths as expected by the greater extinction at the shorter wavelengths. (Terada et al. 2007)

On each side, bright lobes correspond to photons from the central star scattered back to the observer in the disk’s upper layers and in optically thin bipolar cavities. The central star itself is not seen, being heavily extincted by the disk midplane. (Monin and Bouvier 2000)
The spectra show a deep water ice absorption at 3 μm. In addition, many emission and absorption lines can be seen.

Since HV Tau C is well known to have a strong outflow activity, the most prominent emission lines are due to molecular hydrogen lines excited by the jet.

1.93–4.13 μm spectra of HV Tau C on 18 Jan. 2005. The dotted line indicates the continuum of the spectrum. (Terada et al. 2007)

The scattered light from the central star originates within a radius 50 AU from HV Tau C. Thus, it is likely that the material responsible for the ice absorption is outside the scattered light region. (Stapelfeldt et al. 1998, 2003)

Since the midplane disk is optically thick in the 3 μm band, we can observe only the scattered light through the regions above and below the midplane disk.

Schematic view of geometry for the water ice absorption. Gray elliptical areas: scattered light regions; black elliptical areas: midplane of the flared disk. (Terada et al. 2007)
The disk eventually disappears due to accretion onto central star, planet formation, ejection by jets and photoevaporation by UV-radiation from the central star and nearby stars. As a result the young star becomes a weakly lined T Tauri star, which slowly evolves into an ordinary sun-like star.

- SEDs that resemble a stellar photosphere
- Not known for certain whether Class III sources are more evolved than the Class II sources or whether they have simply lost most of their circumstellar material on a faster timescale
- Stage ends in a zero-age main-sequence star
DoAr 25 is located in the L1688 dark cloud in Ophiuchus.

The map of Ophiuchus reveals a multi-filamentary structure with the very opaque L1688 dark cloud showing highest extinction.

$^{13}$CO emission in Ophiuchus, overlaid with the positions of the dense cores detected in submm continuum emission (red circles). Symbol size is proportional to the mass of the core. (Ridge et al. 2006)
The high spatial resolution 865 μm continuum image indicates a well-resolved source with an inclination of about 62°. (Andrews et al. 2008)

The IR SED shows a relatively small thermal excess above the stellar photosphere out to mid-IR wavelengths. Pronounced dips in the IR SED suggest that large inner holes or gaps have been cleared. (Andrews & Williams 2007)

Despite its bright mm emission, this source exhibits only a comparatively small IR excess, suggesting that the material and structural properties of the inner disk may be in an advanced state of evolution. (Andrews et al. 2008)
Stars form from collapsing clouds of gas and dust and in their earliest infancies are surrounded by complex environments that obscure our view at optical wavelengths.

As evolution proceeds, a stage is revealed with three main components:

- **the young star**
- **a circumstellar disk**
- **an infalling envelope**

Most disks around young low- and intermediate-mass stars fall into one of two categories: *(Watson et al. 2007)*

**YSO disks (characterized by dust growth):**

- optically thick at visible and near-IR wavelengths
- rich in molecular gas
- found around Class I and Class II systems

**Debris disks (characterized by planetesimal destruction):**

- optically thin at optical and near-IR wavelengths
- only trace quantities of gas
- found around Class III and older systems
**PROTOTOSTELLAR ENVELOPES: Summary**

- Star plus nebular disk formation
- Most material falls onto the disk
- An accretion-driven stellar or disk wind begins to clear envelope gas away from the rotational poles
- Object in this early embedded YSO phase are completely invisible at $\lambda < 25 \mu m$

- The later "embedded YSO phase" (YTTS)
- Spectrum rises beyond 2 $\mu m$
- In most embedded objects, most the scattered light from walls of the outflow cavity in the envelope
- The embedded phase is thought to last for a few times $10^5$ yr

- The infalling envelope disperses
- Optically visible "classical" T Tauri star (CTTS) with a circumstellar disk
- Accretion through the disk slows, but continues
- This "Class II" phase lasts from $10^6$-$10^7$ yr
- Spectrum has near-IR to far-IR excess of photospheric values, but flat or falling longward of 2 $\mu m$
- Light reflected from the top and/or bottom surfaces of an optically thick disk will dominate

- Accretion through the disk subsides due to lack of replenishment or disk gaps
- The disk becomes optically thin, and the system evolves into a "weak-lined" T Tauri star (WTTS)
- Spectrum is that of a stellar photosphere with a small amount of infrared excess at mid-IR and far-IR
- No emission lines as strong as those of the CTTS


Grains without ice mantles are more effective at polarizing radiation in the diffuse interstellar medium. (Whittet et al. 2001)

Polarization within the heated cavity of B1-c. $N_2H^+$ (grayscale; white indicates high emission, black indicates diminished $N_2H^+$ emission) on a 850 μm continuum image (gray contours). Length of the vectors indicates polarization percentage. (Matthews et al. 2008)

The uniform value of polarization percentage, and the anti-correlation of the $N_2H^+$ and $^{18}C^1O$ emission in the cavity indicate that all the grains are heated to a temperature sufficient to desorb CO from grain surfaces.

Polarization percentage as a function of measured intensity. Dashed line indicates 1.4% polarization. (Matthews et al. 2008)
Maps of three Stokes parameters ($I$, $Q$, and $U$) are combined to yield the polarization percentage and polarization position angle according to the following relations:

$$p = \frac{\sqrt{Q^2 + U^2}}{I}, \quad \theta = \frac{1}{2} \arctan \left( \frac{U}{Q} \right).$$

The uncertainties in each of these quantities are given by

$$dp = p^{-1} \sqrt{dQ^2 Q^2 + dU^2 U^2}, \quad d\theta = \frac{28.6}{\sigma_p}.$$

where $p$ is the signal-to-noise ratio in $p$, or $p/dp$.

The polarization percentages were debiased according to the expression

$$p_{\text{db}} = \sqrt{p^2 - dp^2}.$$

B1-c was identified as a potential star-forming core in dust emission polarimetry measurements of the main molecular core of Barnard 1.

The core shows strong polarization, with evidence that the polarization percentage is constant to the core centre.

Typically, polarization percentage diminishes toward peaks in intensity, resulting in what are called ‘polarization holes’. (Matthews and Wilson 2002)

850 μm polarization pattern toward the main core of Barnard 1 overlaid on the Stokes I map. All vectors are associated with Stokes I values greater than 20% of the B1-d peak flux. Red vectors have polarization percentage, $p$, less than 1%. The vectors are accurate in position angle to better than 10°. (Matthews and Wilson 2002)
Formation of stellar systems: The evolution of SED (low mass star formation)

- **Class 0**
  - The core is cold, 20-30K

- **Class I**
  - An infrared excess appears

- **Class II**
  - Classical T Tauri Star (CTTS)
    - The peak shifts as a disk forms

- **Class III**
  - Weak-lined T Tauri Star (WTTS)
    - The disk dissipates

André, 2002, EAS(vol. 3)
• **Class II**
  Based on the slope (as previous slides), $-1.5 < \alpha_{IR} < 0$
  - Classical T Tauri Star (CTTS)
  - After envelope infall has ceased, **dusty disk produce** IR emission. SED is much **broader** than a single blackbody

• **Class III**
  Based on the slope (as previous slides), $\alpha_{IR} < -1.5$
  - Weak-lined T Tauri Star (WTTS)
  - The **disk dust is dissipated or coagulated**. In the SED, emission from the central star and only small contribution from the disk are detectable

- **Disk Property**
  153 disks in the Taurus-Auriga star formation region (Andrews & Williams, 2008) Based
Formation of stellar systems: CTTS (Class II)

- Classical T Tauri Star
  - T Tauri Star with strong Hα emission line - much brighter than other stars of similar T in IR
    - Dust in disk absorb light from central star and reradiate @ IR
    - ‘reprocessing’ or ‘irradiated’ (or ‘passive’) disk
    - Disk accretion produces jets and winds

- Figure (GM Aur star)
  - Scattered light (top left), Model (top right), and SED (below)
  - Model for the stellar photosphere emission (dark line)
  - Model combining the stellar emission and the disk excess emission (white line)

Watson et al., ‘Multi-wavelength imaging of YSO disks...’
Formation of stellar systems: Property of the disk

- Flat disk?
  Or Flared disk?

  - ‘Flared disk’: generally become proportionately thicker with increasing radius

---

Spectral Index (s) is from $\lambda L_\lambda \propto \lambda^s$

Lee Hartmann, 2000, *Accretion Processes in Star Formation*

Milla-Gabet et al., ‘The circumstellar environments of ..’
Formation of stellar systems: WTTS (Class III)

- Weak-lined T Tauri Star: T Tauri Star which has no strong optical excess emission @ NIR
  - Emission from the central star is only detectable
  - No jets or massive outflows
  - Solar-type magnetic activity with low-mass, pre-main-sequence star
  - Nominal definition: Young star with
    \[ W_\lambda (H\alpha) < 10 \text{ A} \]
  - Narrow H\alpha emission line

Lee Hartmann, 2000, *Accretion Processes in Star Formation*
Formation of stellar systems: CTTS & WTTS (1)

- Hα emission at 656.2 nm
- Figure: Hα profiles (as a function of the velocity shift from line center)
- Wide velocity width (± 200 km/s) is contributed to wind expansion
  - DF Tau: wide and strong Hα profile
    → CTTS
  - DI Tau: narrow and weak Hα profile
    → WTTS

Lee Hartmann, 2000,
Accretion Processes in Star Formation
WTTS nominal definition
Young star with \( W_\lambda(H_\alpha) < 10 \, \text{Å} \)

Redding-corrected
(K-L) colors
(flux ratios of
3.5 \( \square \) emission
to 2.25 \( \square \) emissions)

- \((K-L) < 0.3 \quad W_\lambda(H_\alpha) < 10 \, \text{Å} \) \rightarrow WTTS

- Otherwise, \rightarrow CTTS
• **Aurora Sicilia-Aguilar, ‘Disk Evolution at the Ages of Planet Formation’ (doctoral dissertation)**

• **Two young clusters in Cep OB2 association**
  – **Tr 37 (red)** is placed at the edge of bubble (Cep OB2) containing bright O6 star HD 206267
    - ~165 low-mass candidate
    - The age of the cluster members around 4 Myr
  
  – **NGC 7160 (blue)** lies near the center of a bubble.
    - ~50 low-mass candidates
    - The age of the cluster members around 10 Myr

![Diagram showing Tr 37, NGC 7160, and HD 206267 within Cep OB2 association]
Formation of stellar systems: Observations (2)

- **Tr 37** members (upper) and **NGC 7160** members (below)
  - WTTS (green)
  - CTTS (red)
  - HD 206267 (blue) in **Tr37**
  - High- and intermediate-mass stars (open stars)

- Filled symbols: confirmed or those where Li absorption is detected
- Open figures: Li abs. is not detectable due to poor signal-to-noise
Formation of stellar systems: Observations (3)

- SEDs of low-mass stars in **Tr 37 (avg 4 Myr)** → about 40-45% of the members in Tr 37 (total ~165) have accreting, circumstellar disks
  
  - black dotted line: similar spectral type derived from Kenyon & Hartmann (1995)
  - magenta dashed line: the median disk
  - light blue line: the median disk emission

- **ex. CTTS**
  according to Hα, even if it shows no sign of a disk

- **ex. WTTS**
  according to Hα

- **ex. CTTS**
  disk @ 5.8 μm and longer
  = presence of outer disk
Formation of stellar systems: Observations (4)

- **SEDs of low-mass stars in NGC 7160 (about 10 Myr)** (according to the same procedures followed for Tr 37) → only 1 sample shows indications of active accretion (CTTS).
  - black dotted line: similar spectral type derived from Kenyon & Hartmann (1995)
  - magenta dashed line: the initial disc mass in Taurus
  - light blue line: the media

![Diagram showing examples of ex. CTTS and ex. WTTS according to Hα]
• Evolution of protoplanetary disks
  – Theoretical isochrones from Siess et al. (2000) for 1, 10 and 100 Myr are shown together
  – CTTS (red circle)
  – WTTS (green triangle)

→ Significant difference in the apparent age scatter between
  Tr 37 (dispersion) and
  NGC 7160 (mostly located along the 10 Myr)
• Low-mass SED from the ages 1 to 10 Myr.
  – Taurus (1-2 Myr) - red
  – Tr 37 globule (1 Myr) - cyan
  – Tr 37 (avg ~4.5 Myr) - violet
  – NGC 7160 (10 Myr) – pink
  – TW Hya (10 Myr) - black

  – The steep line slopes of the SED are shown in NGC 7160 and TW Hya (these are comparably older than others)
Equilibria of gas planets

A. González, P. Kollmann & K. Jockers

30 June - 2 July 2009, Course on Origin of solar system

Max-Planck Institute for Solar System Research, Katlenburg-Lindau
Outline

- Giant planet formation
- Possible protoplanets and their probabilities
Outline:

Part 1:
The giant planets of our solar system
Disk instability versus nucleated instability hypothesis
Quasi-static models of giant planets and their limitations

Part 2:
Thesis of Christopher Hans Broeg: What kind of quasi-static planets can exist around other stars?
The giant planets of our solar system
Energy balance within giant planets

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Jupiter</th>
<th>Saturn</th>
<th>Uranus</th>
<th>Neptune</th>
</tr>
</thead>
<tbody>
<tr>
<td>Effective temperature (K)</td>
<td>124.4±0.3</td>
<td>95.0±0.4</td>
<td>59.1±0.3</td>
<td>59.3±0.8</td>
</tr>
<tr>
<td>Energy balance</td>
<td>1.67±0.09</td>
<td>1.78±0.09</td>
<td>1.06±0.08</td>
<td>2.61±0.28</td>
</tr>
<tr>
<td>Internal energy flux (W m⁻²)</td>
<td>5.44±0.43</td>
<td>2.01±0.14</td>
<td>0.042±0.047</td>
<td>0.433±0.046</td>
</tr>
<tr>
<td>Internal power/unit mass (10⁻¹¹ W kg⁻¹)</td>
<td>17.6±1.4</td>
<td>15.2±1.1</td>
<td>0.392±0.441</td>
<td>3.22±0.34</td>
</tr>
</tbody>
</table>

Energy radiated/Energy received from the Sun

For Jupiter and Neptune the value is consistent with the assumption that their excess radiation is caused by cooling of their interior since their origin.
Derived from interior models matching the observed gravitational fields.

Each of the four giant planets of our Solar system are believed to consist of a central, dense core and surrounding envelope composed of hydrogen, helium, and small amounts of heavy elements.

Inside Jupiter and Saturn, hydrogen, which is in molecular at low pressures, is thought to become metallic in the 1 – to 3-Mbar region. This transition could be abrupt or gradual.

Uranus and Neptune, contain, in a relative sense, more heavy elements.
Note that Jupiter and Saturn may not have a core (judging from these models).

Figure 2. Limits on the abundances of heavy elements in the four jovian planets in our solar system. For each planet, the point on the left represents the total amount of high-Z material, whereas the (lower) point on the right shows the amount of heavy elements segregated into the planet's core. For Jupiter and Saturn, the thick lines represent solutions with additional constraints obtained from evolution models. Note the high level of uncertainty, especially regarding the core masses of Jupiter and Saturn. Models of Jupiter with small cores (i.e., less than 2 M⊕) require significant enrichments in heavy elements (i.e., more than 20 M⊕).
In principle planets are made in a similar way as protostars:

An instability is needed to initiate planet formation. There can be a non-stationary collapse phase.

But:
The amount of matter available to make the planet is limited by the presence of the host star and its protoplanetary disk.
Hill sphere is the region where the gravitation of a protoplanet dominates. To start accretion the planet must *already* have its own region of gravitational influence.

The Hill sphere is the area between Lagrange points $L_1$ and $L_2$ made spherical.

$$R_H = \left( \frac{m_2}{3(m_1 + m_2)} \right)^{1/3} a,$$
GAS ACCUMULATION THEORIES_2

Typical nebula densities are more than two orders of magnitude below the stability limit, i.e. a local density enhancement necessary to form a planet will be immediately smoothed out again by the tidal forces of the surrounding nebula.

There are theories proposed that create giant gaseous protoplanets out of disk instabilities (Boss and coworkers). But in this lecture we will follow the more likely idea that the initial gravitational disturbance is created by a core of coagulated solid planetesimals (nucleated instability hypothesis).

*Let's assume that a core of coagulated solid planetesimals can form in the disk despite of the presence of the host star.*
Core accretion and gas capture

The core mass as a function of total mass for a protoplanet consisting of a solid core + gaseous envelope calculated according to the equations of stellar structure.

The masses are measured in units of the Earth’s mass.

Critical core mass

Jupiter’s mass

Total mass

\( f = \text{ratio of the assumed grain opacities to the interstellar values.} \)

Dashed lines: hydrodynamically unstable.
NUCLEATED INSTABILITY
Critical masses of protoplanets as a function of nebula mid-plane density

- minimum mass
- arbitrarily enhanced nebula
- solar nebula
- convective envelope
- static envelope

Critical density of a Jupiter-mass nebula fragment at Jupiter’s position. The value is the mean density of a condensation that is Jeans critical and fits into the Hill sphere.
Pollack et al. (1996) constructed models in which they simulated the concurrent accretion rates of both the gaseous and solid components of giant planets. They used an evolutionary model having three major components:

- A calculation of the three-body accretion rate of a single dominant-mass protoplanet surrounded by a large number of planetesimals.
- A calculation of the interaction of accreted planetesimals with the gaseous envelope of the growing giant protoplanet.
- A calculation of a sequence of quasi-hydrostatic models having a core/envelope structure.

In these models, there are three main phases to the accretion of Jupiter and Saturn.

1. **Phase 1** is characterized by rapidly varying planetesimal and gas accretion.
2. **Phase 2** of accretion is characterized by relatively time-invariant values of $dM_z/dt$ and $dM_{xy}/dt$, with $dM_{xy}/dt > dM_z/dt$.
3. **Phase 3** is defined by rapidly increasing rates of gas and planetesimal accretion, with $dM_{xy}/dt$ exceeding $dM_z/dt$ by steadily increasing amounts.

Fig. 1. Evolutionary path of a $1 \, M_\odot$ protostar in an infrared HR diagram (solid line). The numbers indicate the time (in years) since the formation of the (final) hydrostatic core. For comparison, the evolutionary path of a conventional fully hydrostatic $1 \, M_\odot$ pre-main sequence star is also included (broken line).
HYDRODYNAMICAL MODELS OF GIANT PLANET FORMATION NEAR STARS

• A major result of the hydrodynamical studies is that the proto-giant planets may pulsate and develop pulsation-driven mass loss. Only if the pulsations are damped can gas accretion produce Jupiter-mass envelopes. Extra-solar planets with a minimum of $0.5 \, M_J$ up, probably require efficient gas accretion and should satisfy the convective outer envelope criterion (Wuchterl 1993).

• Wuchterl (1993) has shown that almost all nebula conditions, from a literature collection of nebula models, result in radiative outer envelopes at the critical mass. Nonlinear radiation hydrodynamical calculations with zero-entropy gradient convection show that Uranus/Neptune-type giant planets are produced under such circumstances. Jupiter-mass planets should then be the exception.
Part 2:

Part 1:
The gas planets of our solar system
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Quasi-static models of giant planets and their limitations

Part 2:
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The model protoplanet

- A **protoplanet** is a spherical object that consists of a **solid core** surrounded by an **envelope** consisting of gas and dust.
- The protoplanet is embedded in a **nebula**. In this case the protoplanet is defined as the core and the region gravitationally bound to the core (called the **Hill sphere**).
- At its outer boundary the protoplanet is in **thermal and pressure equilibrium** with the nebula.
- It accretes **solid planetesimals** at a constant rate. The energy gained in this way must be transported back to the planet’s surface and radiated away. Therefore the **luminosity** of the protoplanet is determined by the **inflow of planetesimals**, assumed constant.
- As the core mass of the planet increases, so does its Hill radius, i.e. its gaseous envelope.
- When the critical core mass is reached the star **looses hydrostatic equilibrium**. Hydrodynamic calculations are necessary to determine the fate of the planet.
We want to get an overview about the possible planets of a given host star. For simplicity we restrict ourselves to static planets.

\[ \rho_c = 5500 \text{ kg m}^{-3} \]

\[ r_{\text{core}} = \sqrt[3]{\frac{3}{4\pi} \frac{M_c}{\rho_c}}. \]

\[ r_{\text{Hill}} = a \sqrt[3]{\frac{M}{3M_\ast}}. \]

Outer radius = Hill radius

\[ L = -(\Phi - \Phi_0) \dot{M}, \quad \text{with} \quad \dot{M} = \text{konst} \quad (\text{z. B. } 10^{-6} M_\oplus a^{-1}) \]

\[ \Phi_0 = \Phi(r_{\text{Hill}}). \]

\[ T_{\text{neb}} = 280 \cdot \left( \frac{a}{1\text{AU}} \right)^{-1/2} \left( \frac{L_\ast}{L_\odot} \right)^{1/4} \text{ K.} \]

Equilibrium temperature in the radiation field of the host star

The nebula pressure remains unspecified, core pressure and core mass variable.
Calculations

- System of first order differential equations describing mass, pressure, temperature etc. with distance to the core.
- Similar to equations of stellar structure. They include the possibility of convection.
- Assumption of nebula of hydrogen and helium.
- Radiative and pressure equilibrium at layer between protoplanet and nebula.
- Infall of planetesimals onto the core releases gravitational energy which leads to the luminosity of the protoplanet.
- Independent parameters are mass of core, pressure at core, planetesimal accretion rate, distance to star, luminosity of star. The accretion rate depends on the distance to the star and will be an estimate.
- Solve for equilibrium solutions.
Possibilities

- vary core mass $M_c$ and pressure at core $P_c$
- fix all other parameters
- shown is calculated mass of envelope
Region I: fully developed terrestrials

- no additional mass outside core
- finite pressure only at the core

- envelope mass increases with pressure
- they are big cores
- huge variety of surface pressures possible
Region II: fully developed Jovians

- most of the mass outside core
- pressure vanishes outside the planet
- envelope mass dominates total mass
- high pressure at the core
- different core sizes possible
Region III: loose planetesimals

- no strong increase of mass outside core
- empty environment
- cores are too small
Region IV: real protoplanets

- most of the mass outside the core
- finite pressure even at huge distances

- envelope mass is independent of $M_c$ and $P_c$
- thick envelopes impossible if $M_c$ too big
- envelopes teared away by tides if $P_c$ too low
Effect of Boundary Conditions

plots are for \textit{Jupiter's distance} to the star

- regions I and II are very small because there are \textit{no} equilibrium solutions

- in that distance fully developed planets do only form quasistatically until they reach \textit{critical mass}

fixed values of

- protoplanet's \textit{luminosity}
- \textit{temperature} at core

protoplanet's luminosity \textit{depends on planetary mass} and fixed accretion rate

- \textit{temperature at outer envelope} layer is used and calculated from fixed star's luminosity
Probabilities

assume
• all equilibria can be reached are \emph{equally probable} and
• all equilibria are \emph{stable}
• \(M_c\) and \(P_c\) are \emph{evenly distributed} on a log-scale
then
• \emph{number of equilibrium states} with a certain mass is a measure for the \emph{probability of protoplanets} with this mass \(\Rightarrow\) mass spectra
• mass spectra can be used to classify protoplanets

In the following: \(M_\star = 1 M_\odot\) \(\dot{M} = 10^{-4} M_\odot a^{-1}\).
Class G Protoplanets

\( G = \text{„ganz heiß“} \) (german) = \( \text{„very hot“} \)

this distribution describes all protoplanets very close to the star, with orbit periods < 4 days

- peak is created by region II solutions (fully developed jovians)
- many planets with masses > Jupiter
- peak is due to region IV solutions (protoplanets)
- many „hot Neptunes“
Class H Protoplanets

H = hot planets

this distribution describes all protoplanets at close to the star, with orbit periods between 4 and 16 days.

- mass of heavy population < Jupiter
- peaks merge with increasing distance

4 days period

16 days period
Class J Protoplanets

J = jovian planets

this distribution describes all protoplanets in \textit{relatively big distance} to the star, with orbit periods > 16 days

- only one dominant peak
Jupiter

this distribution describes protoplanets in the *outer region* of a solar system, with Jupiter‘s orbit period

- peak is not at Jupiter‘s mass

because Jupiter and class J protoplanets are not quasistatic for all times

→ Peak at Jupiter‘s mass can not be found by used method

→ class J mass spectra does mainly represent protoplanets

accretion rate smaller as before, which is realistic for Jupiter‘s orbit
Protoplanets can be formed...

- *in situ*
  (at the position of the final planet)
- *quasistatically*
  (protoplanet is at hydrostatic equilibrium for all times)

This will be the assumptions of the upcoming considerations.
The following slides will show that these assumptions are reasonable and true for at least one case.
Extrasolar planets: Many “hot Jupiters”

Can “hot Jupiters” form at the places they are found today or were they formed at larger distances to their host star and migrated to the place where they are found now?
The case of exoplanet “HD149026b”
A very hot Jupiter

• 0.4 $M_J$ mass
• high density
• 0.042 AU distance to star

$HD149026b$ was not considered to be formed in situ.
Using the probability manifolds discussed previously, one can represent this planet with the following parameters:

- $a = 0.042 \text{ AE}$,
- $M_* = 1.3 M_\odot$,
- $T_X = 1754 \text{ K}$.

Very high accretion rate!

Hydrodynamical calculations of the time evolution of this planet show that this equilibrium can be reached and is stable.
Summary
Part 1:

With the exception of Uranus the gas giants of our solar system radiate more energy than they receive from the Sun. This excess radiation is considered to be primordial.

Uranus and Neptune have cores consisting of ice and rock. For Saturn, and in particular for Jupiter such cores are likely but the core masses are uncertain.

To make a gaseous planet a significant excess of density is required at the start. It is likely that gaseous planets grow from a core consisting of coagulated solid planetesimals (ice and/or rock).

A gaseous planet can grow from a core to a maximum total mass. If this mass is exceeded, rapid non-hydrostatic accretion will set in. It is not clear how much mass the planet can gain in this phase as instabilities caused by opacity (kmechanism) may occur.
Summary
Part 2:

In his ph. d. thesis Ch. Broeg has studied the manifold of possible hydrostatic models of gas planets. He finds 4 types of hydrostatic planets: “fully developed terrestrials”, “fully developed Jovians”, “loose planetesimals”, and “real protoplanets”. The rate of accretion of solid planetesimals must be known and it is assumed that “there are always enough solid planetesimals available”. Probability distributions of planets as function of their distance from the host star are calculated.

For the so-called “hot Jupiters” (planets very close to their host star) a bimodal distribution of such planets is found, with a narrow, high mass peak of type II planets and a wide, lower mass peak of type IV planets. For slightly larger distances the two peaks merge into a single peak.

Exoplanet “HD149026b” (0.4 MJ, 0.04 AU distance) can be formed at the place it is observed, if the accretion rate is high enough. It is hydrodynamically stable.
Origin of solar systems
30 June - 2 July 2009
by Klaus Jockers (jockers@mps.mpg.de)
Max-Planck-Institut of Solar System Science
Katlenburg-Lindau

Part 5
Condensation and growth of solid bodies in protoplanetary disks
Outline

Theoretical considerations concerning protoplanetary disks:
• Why is the disk expected to be tapered (scale height versus distance from host star)
• Angular momentum in the disk

Condensation and growth of solid bodies
• Time scales of planetesimal formation
• Infall of grains onto the disk
• Growth of sub-meter particles by coagulation
• The drift problem in a disk partially supported by pressure
• Growth of planetesimals > 1 km (gravitational regime)
The extent of a Keplerian disk perpendicular to the disk plane:
Pringle, J.E., accretion disks in astrophysics,

Keplerian disk: Mass of disk negligible as compared to mass of central star.
Assumption: no forces except gravitation.

Hydrostatic equilibrium perpendicular to the disk plane:
\[
\frac{1}{\rho} \frac{\partial p}{\partial z} = \frac{\partial}{\partial z} \left[ \frac{GM}{\sqrt{r^2 + z^2}} \right] ; \text{ for } z \ll r: \frac{1}{\rho} \frac{\partial p}{\partial z} = -\frac{GMz}{r^3}.
\]

Replace the density \( \rho \) by the pressure \( p \) using the ideal gas law:
\[
p = n \cdot kT, \quad p = \rho \cdot \frac{kT}{m}
\]

\[
\frac{1}{p} \frac{dp}{dz} = -\frac{m}{kT} \frac{GM}{r^3} z
\]

Integrate, separating the variables:
\[
p = p_0 \exp\left(-\frac{m}{kT} \frac{GM}{r^3} z^2\right)
\]

Introduce Gaussian scale height \( H \):
\[
p = p_0 \exp\left(-\left(\frac{z}{H}\right)^2\right)
\]
Note \( z^2 \), not \( z! \)

By comparison:
\[
H = \sqrt{\frac{2kT}{m} \cdot \frac{r^3}{GM}}
\]

For physically reasonable temperature distributions, like \( T \sim r^{1/2} \), \( H \) rises with distance from the central star, i.e. the disk is tapered.
Outline

Theoretical considerations concerning protoplanetary disks:
• Why is the disk expected to be tapered (scale height versus distance from host star)
• Angular momentum in the disk

Condensation and growth of solid bodies
• Time scales of planetesimal formation
• Infall of grains onto the disk
• Growth of sub-meter particles by coagulation
• The drift problem in a disk partially supported by pressure
• Growth of planetesimals > 1 km (gravitational regime)
Internal dynamical evolution of the disk:

Redistribution of angular momentum can provide additional mass to the central star.

*Magnetic torque* can reduce rotation of star if ionization is high (frozen-in magn. field). Process needs ionized gas that may not be available at large distances from the proto-Sun.
Protoplanetary disks apparently do not extend all the way down to the surface of the star. Magnetic interactions at the corotation point funnel some of the disk’s gas onto the star and expel other gas in rapid centrifugally driven bipolar outflow which carries with it a substantial amount of angular momentum (?)..

*Gravitational torques*: Nonaxisymmetric local instabilities can create spiral density waves like in galaxies or in the Saturnian ring system which limit the allowed mass of the disk.
Large protoplanets may clear annual gaps surrounding their orbits, excite density waves transporting angular momentum outwards. Similar effects can arise if protostar rotates sufficiently rapidly to become triaxial. But as observations indicate, such high rotation rates seem to be unlikely.
Internal dynamical evolution of the disk (continued):

**Viscous torques:** Molecules move on Keplerian orbits, i.e. its transverse speed decreases outwards. Collisions (or turbulence) transfer mass inward and angular momentum outward (Lynden-Bell and Pringle MNRAS 168, 603-637, 1974).

Evolution on diffusion timescale:  
\[ t_d = \frac{\ell^2}{\nu_v}, \]

\( l = \) radius of the disk.

Viscosity largely unknown.

Molecular viscosity:

\[ \nu_v \sim \ell_{fp} c_s, \]

\( \ell_{fp} \) mean free path of molecules and \( c_s \) sound speed.

Turbulent viscosity:

\[ \nu_v = \frac{2}{3} \alpha_v c_s H_z, \]

\( 10^{-4} < \alpha_v < 10^{-2}. \)

\( H_z \) depends on temperature and the disk temperature on opacity perp. to midplane. Well inside the orbit of Mercury the interstellar dust grains are all evaporated and opacity is caused by \( H_2O \) and CO molecules and H ionization. At larger distances from the star the temperature is below 2000K and micrometer-sized dust is the dominant source of opacity.
Outline

Theoretical considerations concerning protoplanetary disks:
• Why is the disk expected to be tapered (scale height versus distance from host star)
• Angular momentum in the disk

Condensation and growth of solid bodies
• Time scales of planetesimal formation
• Infall of grains onto the disk
• Growth of sub-meter particles by coagulation
• The drift problem in a disk partially supported by pressure
• Growth of planetesimals > 1 km (gravitational regime)
Condensation and growth of solid bodies

Timescales for planetesimal formation:

The age of most chondrites (primitive meteorites) is 4.56 Gyr and they formed within a period ≤ 20 Myr

Evidence from extinct $^{26}$Al ($t_{1/2} = 0.72$ Myr) in carbonaceous chondrites suggests that first solid material formed only a few million years after the last injection of freshly nucleosynthesized matter (but other explanations exist). This timescale is the timescale of collapse discussed earlier. Evidence is based on observation of the daughter product $^{26}$Mg close to $^{27}$Al.

The freshly nucleosynthesized matter could come from stellar winds produced by a nearby asymptotic giant branch (AGB) star.

Isotope analysis of primitive meteorites indicates that they still contain interstellar grains.
Outline

Theoretical considerations concerning protoplanetary disks:
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Protoplanetary disk:
Infall stage
Gas and Dust
Duration of infall stage comparable to free-fall collapse time of the core $\sim 10^5$-$10^6$ yr. Matter with low specific angular momentum falls into the central star. Matter with high specific angular momentum cannot reach the central star and forms the disk. In the following it is shown that for such matter exactly half of the gravitational energy gained during infall goes into the kinetic energy of the orbiting body and the other half is converted to heat.

Kepler velocity: Body (mass $m$) on circular orbit around central star with mass $M$:
Parcels of gas fall from both sides of the disk from infinity to a circular orbit at heliocentric distance $r_{\text{SUN}}$ and meet there.

Gravitational energy gained:

$$\frac{GMm}{r}$$

Kinetic energy:

$$\frac{m}{2}v^2 = \frac{m}{2} \cdot \frac{GM}{r}$$

$$\Rightarrow \frac{GM_{\text{protostar}}}{2r_{\odot}} = \frac{v_c^2}{2}$$

Half of gravitational energy is converted to orbital kinetic energy, the other half per unit mass, is available for heat.
Equate half of gravitational energy with thermal energy per particle and find:
At 1 AU and 1 solar mass $v_c = 30 \text{ km s}^{-1}$ and the temperature in a hydrogen gas $\sim 7 \times 10^4 \text{ K}$.
But temperature falls quickly because of radiative cooling.
When two clouds meet from both sides of the forming disk, shock fronts form with temperatures $\sim 1500 \text{ K}$ at 1 AU and $\sim 100 \text{ K}$ at 10 AU.
Infall of grains onto the disk:

Acceleration of grains toward disk: $\rho_g$ density of gas, $\rho$ grain density, $c_s$ local sound speed, $R$ grain radius.

$$\frac{dv_z}{dt} = -\frac{\rho_g c_s}{R \rho} v_z - n^2 z,$$

$n$ is Keplerian orbital angular velocity. (mean motion)

$$n = \sqrt{\frac{GM_\odot}{r^3}}$$

Equilibrium settling speed:

$$v_z = \frac{n^2 z \rho R}{\rho_g c_s}.$$

At 1 AU $T = 500$–800K, $\rho_g = 10^{-9}$ g cm$^{-3}$, $c_s = 2.5 \times 10^5$ cm s$^{-1}$, $v_z = 0.03$ (z/H$_2$) cm s$^{-1}$. (H$_2$ nebula).

For 1 µm grain with $\rho = 1$ g cm$^{-3}$ it takes $10^6$ years to fall halfway toward midplane, $10^7$ years for 99.9% of the distance.

Coagulation is needed to form the disk in the available time.
Outline

Theoretical considerations concerning protoplanetary disks:
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FIGURE 12.9 Examples of fractal aggregates produced by numerical simulations. In a ballistic particle-cluster agglomeration (BPCA) process, a seed particle grows by the accumulation of single particles that collide with random impact parameters and from random directions on linear trajectories (hit-and-stick process). A ballistic cluster-cluster agglomeration (BCCA) process proceeds through the coagulation of equal-mass aggregates (again on linear trajectories with random impact parameters and from random directions). (a) BPCA with 1024 monodisperse spherical particles; the simple BPCA process leads to aggregates with a fractal dimension of 3.0. (b) BCCA with 1024 monodisperse spherical particles; these aggregates have a fractal dimension of 1.9. (c) BPCA with 2001 spherical constituent particles, following a power law size distribution with an exponent of $-3.15$. (d) BPCA with 2000 monodisperse spherical particles, aggregated onto a large spherical core. (Blum et al. 1994)
Coagulation

Coagulation is different for fluffy and smooth particles. It depends on electric charge, electric conductivity in the grain and molecular forces. Presently experiments are performed in space under microgravity and in the laboratory.

Planetesimal formation starts with the growth of fractal dust aggregates, followed by compaction processes. As the dust-aggregate sizes increase, the mean collision velocity also increases, leading to the stalling of the growth and possibly to fragmentation, once the dust aggregates have reached decimeter sizes.

Current models indicate a settling time into mm-sized bodies in $10^4$ years.

For more details see:
Outline

Theoretical considerations concerning protoplanetary disks:
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• Growth of planetesimals > 1 km (gravitational regime)
The problem of inward drift in a partially pressure supported disk:
Gas circles star slightly less rapidly than Keplerian rate.

Effective gravity felt by gas:
In circular orbits, the effective gravity is balanced by centrifugal forces $r_{\text{sun}} n^2$.
Since the pressure is much smaller than gravity we can approximate the angular velocity $n_{\text{gas}}$ as

$$g_{\text{eff}} = -\frac{G M_{\odot}}{r_{\odot}^2} - \frac{1}{\rho g} \frac{dP}{dr_{\odot}}.$$

$$n_{\text{gas}} \approx \sqrt{\frac{G M_{\odot}}{r_{\odot}^3}} (1 - \eta),$$

where

$$ \eta \equiv \frac{-r_{\odot}^2}{2 G M_{\odot} \rho g} \frac{dP}{dr_{\odot}} \approx 5 \times 10^{-3}. $$

For estimated protoplanetary disk parameters the gas rotates 0.5% slower than the Keplerian speed. But large particles must move with Keplerian speed, otherwise they will fall into the protostar!
Radial inward drift of planetesimals:

Particles moving at (nearly) the Keplerian speed encounter a headwind which removes part of their orbital angular momentum and causes them to spiral inward towards the star.

- Small particles are strongly coupled to the gas and therefore drift very slowly.

- Kilometer-sized bodies also drift inwards very slowly, because their surface to mass ratio is small.

- Peak inward drift rates occur for particles that collide with roughly their own mass within one orbital period.

At 1 AU a meter-sized body drifts inwards at the fastest rate $\sim 10^6$ km yr$^{-1}$. Because of the difference in (both radial and azimuthal) velocities, small (subcentimeter) grains can be swept on by larger grains which in turn move toward proto-Sun.
Particles must grow through this size range quickly, otherwise they will be lost.

FIGURE 12.11 The inward radial drift rates of solid particles in a protoplanetary disk as a function of size for three values of density: 0.5 (solid line), 2.0 (dashed line), and 7.9 (dotted line) g cm$^{-3}$. Gas parameters are the same as for Figure 12.10. Small particles, with small mass/surface area ratios, are strongly coupled to the gas and compelled to move with (nearly) its angular velocity. As this is less than the keplerian orbital rate, they feel a residual component of the Sun’s gravity, and settle inward at a terminal velocity at which gas drag balances this radial acceleration. Thus, larger and/or denser particles drift more rapidly in this regime. Bodies with large mass/surface area ratios travel in (nearly) keplerian orbits, moving faster than the gas. They experience a ‘headwind’ that causes their orbits to decay; larger and/or denser bodies are less affected by this drag, so the decay rate decreases with increasing particle radius. The radial velocity reaches a peak at the transition between these regimes, at sizes of about a meter. The abrupt changes in slope result from transitions between drag laws for different Knudsen and Reynolds numbers. (Courtesy: Stuart J. Weidenschilling)

No solution of this problem exists at present.
Outline

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• Growth of planetesimals > 1 km (gravitational regime)
Growth from planetesimals to planetary embryos, gravitational regime:

For bodies > 1 km major forces are gravitational interaction and physical collisions and gas drag.

Collision between planetesimals:
\[ v_i = \sqrt{v^2 + v_e^2}, \quad v_e = \left( \frac{2G(m_1 + m_2)}{R_1 + R_2} \right)^{1/2}. \]

\( v \) relative speed at large distances
\( v_e \) escape speed
Impact velocity \( v_i \geq v_e \)
\( v_i \geq 6 \text{ m s}^{-1} \) for rocky 10 km body.

Some of the kinetic energy of the colliding particle must be dissipated \( \rightarrow \)
Rebound velocity = \( v_i \varepsilon \) with \( \varepsilon \leq 1 \). If \( v_i \varepsilon \leq v_e \) particle accretes sooner or later.

The relative speed between planetesimals is critical:

- Only if \( v \ll v_e \) probability for capture of the particle high.

- If relative speed is too high, small grains will not accrete on large grains \( \rightarrow \) instead sandblasting of growing planetesimals.
At a larger scale, growth from cm-sized to kilometer sized planetesimals depends primarily on the relative motions between the various bodies.

**FIGURE 12.10** Contours of constant relative velocity (in cm s\(^{-1}\)) between pairs of particles of density 2 g cm\(^{-3}\) orbiting within a partially pressure-supported gaseous protoplanetary disk are displayed as a function of particle size. Sizes from 1 μm to 10 km are shown; relative velocities are due to thermal motions (dominant at sizes < 10 μm), as well as radial and transverse velocities induced by gas drag. Disk parameters are for the midplane at 1 AU in a nonturbulent minimum mass solar nebula: gas density \(\rho_g = 3.4 \times 10^{-9}\) g cm\(^{-3}\), \(T = 320\) K, \(\Delta v = 61.7\) m s\(^{-1}\). The narrow ‘valley’ in the contour plot results from the fact that equal-sized bodies have identical velocities relative to the gas. (Courtesy: Stuart J. Weidenschilling)

**Model calculations**

**Note:**
- Small relative speeds for small grains.
- Reduced relative speeds for large masses of similar size.
- Plateau for relative speeds of small and large bodies.
Growing planets in the gravitational regime:

Without proof:
Gravitational enhancement factor

$$F_g = 1 + \left( \frac{v_e}{v} \right)^2$$


Mass accretion, $\rho_s$ volume density of planetesimal swarm:
$R$ radius of planetary embryo.

$$\frac{dM}{dt} = \rho_s v \pi R^2 F_g,$$

For $v \leq v_e$, but $v \ll v_e$, $F_g \approx 1$ and
$$\frac{dM}{dt} \propto R^2.$$

For $v \ll v_e$, $F_g = \left( \frac{v_e}{v} \right)^2$

If we express $v_e$ by the radius $R$ of the planetesimal, we get:

$$v_e = \sqrt{\frac{2GM}{R}}$$

and

$$F_g = \frac{2GM}{R} \cdot \frac{1}{v^2} = \frac{8\pi}{3} \cdot \frac{\rho_p G}{v^2} \cdot R^2$$

$$\frac{dM}{dt} = \rho_s v \pi R^2 \cdot F_g = \frac{8\pi^2}{3} \cdot \frac{\rho_s \rho_p G}{v} \cdot R^4,$$

i.e. $\frac{dM}{dt} \propto R^4$.

i.e. for $v \approx v_e$ the mass grows $\sim R^2$ and for $v \ll v_e$ the mass grows $\sim R^4$ (runaway growth).
Growing planets in the gravitational regime (continued):

If one body is larger than all the others, it will not stir up the mean relative velocity $v$ as much as if all bodies have similar size. This will allow continued fast growth until three-body interactions become important.
Transfer to surface densities and calculation of planetesimal growth in radius:

Gaussian scale height (calculated earlier)

\[ H_z = \sqrt{\frac{2kT}{m} \cdot \frac{r^3}{GM}} = \sqrt{\frac{2kT}{m} \cdot \frac{1}{n}} = v_{th} \cdot \frac{1}{n} \]

\( n \) is mean motion.

Lissauer: If the proto-Sun’s gravity is dominant force in vertical direction and if the relative velocity between planetesimals is isotropic, then

\[ H_z = \frac{v}{\sqrt{3n}}. \]

The two scale heights are the same except of \( \sqrt{3} \).

Transfer to surface mass density of planetesimals \( \sigma_\rho \) (g cm\(^{-2}\)):

\[ \sigma_\rho = \sqrt{\pi} \rho_s H_z. \]

Growth of radius:

\( \rho_p \) density of planetary embryo

\[ \frac{dR}{dt} = \sqrt{\frac{3}{\pi} \frac{\sigma_\rho n}{4\rho_p}} F_g, \]
Growth time of planets:

\[ \frac{dR}{dt} = \sqrt{\frac{3}{\pi}} \frac{\sigma_p n}{4\rho_p} F_g, \]

For Earth \( F_g = 7, \sigma_p = 10 \text{ g cm}^{-2}, \ n = 2 \times 10^{-7} \text{ s}^{-1}, \ \rho_p = 4.5 \text{ g cm}^{-3} \), growth time \( 2 \times 10^7 \text{ yr} \), or better \( 10^8 \text{ yr} \), if depletion of planetesimals in later stage of accretion is considered.

Problems with outer planets. For Jupiter \( \sigma_p = 3 \text{ g cm}^{-2} \), heavy element mass 15-20 Earth masses, growth time \( > 10^8 \text{ yr} \). Surface density of solar nebula drops \( \sim r^{-3/2} \), growth time of Neptune is many times the solar system age.
\[ R_H = \left( \frac{m_2}{3(m_1 + m_2)} \right)^{1/3} a, \]

Fig. 3.9. The location of the Lagrangian equilibrium points (open circles) and associated zero-velocity curves for a mass \( \mu_2 = 0.01 \). The dashed line denotes the circle of unit radius centred on the mass \( \mu_1 \).
End of growth of planetary embryos

Area within reach of the growing embryo is \( \sim 4 \) times its Hill sphere.

Hill sphere: sphere of gravitational influence (limited by Lagrange points, previous view graph).
Radius \( R_H \) of Hill sphere:

\[
R_H = \left( \frac{m_2}{3(m_1 + m_2)} \right)^{1/3} a,
\]

Mass of planetary embryo which has accreted all mass within a ring of width \( 2\Delta r_\oplus \):

\[
M = \int_{r_\oplus - \Delta r_\oplus}^{r_\oplus + \Delta r_\oplus} 2\pi r' \sigma_\rho(r') dr' \approx 4\pi r_\oplus \Delta r_\oplus \sigma_\rho(r_\oplus).
\]

If \( \Delta r_\oplus = \varpi R_H \) we obtain maximum mass \( M_i \) (in g) to be accreted by a planetary embryo orbiting a star of 1 \( M_\odot \):

For Earth \( M_i = 5 \times 10^{26} \text{ g} \). 1 Earth mass = 6 \( \times 10^{27} \text{ g} \).

\[
M_i \approx 1.6 \times 10^{25} (r^2_{AU} \sigma_\rho)^{3/2},
\]
FIGURE 12.12 Snapshots of a planetesimal system on the $a$–$e$ plane. The circles represent planetesimals and their radii are proportional to the radii of planetesimals. The system initially consists of 4000 planetesimals whose total mass is $1.3 \times 10^{27}$ g. The initial mass distribution is a power with index $\xi = -2.5$ over the mass range $2 \times 10^{23}$ g $\leq m \leq 4 \times 10^{24}$ g. The system is followed using an $N$-body integrator, and physical collisions are assumed to always result in accretion. The numbers of planetesimals are 2712 ($t = 100\,000$ yr), 2200 ($t = 200\,000$ yr), 1784 ($t = 300\,000$ yr), 1488 ($t = 400\,000$ yr), and 1257 ($t = 500\,000$ yr). The filled circles represent planetary embryos with mass larger than $2 \times 10^{25}$ g, and lines from the center of each planetary embryo extend $5\,R_H$ outwards and $5\,R_H$ inwards. (Kokubo and Ida 1999)

Making planetary embryos close to the Earth, numerical calculations:
Final stages of planetesimal accumulation

The self-limiting nature of runaway growth implies that massive protoplanetary embryos form at regular intervals in semimajor axis.

Their random velocities are no longer strongly damped by energy equipartition with the smaller planetesimals. Therefore the embryos will pump up each other’s velocities. The orbital eccentricities will increase and the orbits possibly intersect.

Subsequent orbital evolution is governed by close gravitational encounters and violent, highly inelastic collisions. (The Earth’s moon is believed to be generated during such a collision).

The orbital evolution in the inner solar system can be studied with programs calculation the motion of N bodies in the gravitational field of the Sun.

During this process of mutual violent collisions the chemical composition of the resulting planets is averaged over some range of heliocentric distance.
FIGURE 12.13  (a) Simulation of the final stages of terrestrial planet growth in our Solar System using an N-body code that includes Jupiter and Saturn and that assumes all physical collisions lead to mergers. Planetary embryos are represented as circles whose radii are proportional to the embryos’ radii. The locations of the planetary embryos are displayed in semimajor axis–eccentricity phase space at the times indicated. (b) Synthetic
I. (b) Synthetic terrestrial planet systems produced by eight different $N$-body simulations of the final stages of planetary accretion. The final planets are indicated by filled circles centered at the planet's semimajor axis. The horizontal line through each circle extends from the planet's perihelion to its aphelion; the length of the vertical line extending upward from a planet's center is proportional to its inclination. The numbers under each circle represent the planet's final mass in $M_\oplus$. The results of the simulation shown in part (a) are presented in row E. (Courtesy: John Chambers)
Summary, part 1

In a Keplerian disk the Gaussian pressure scale height =

\[ H = \sqrt{\frac{2kT}{m} \cdot \frac{r^3}{GM}} \]

For reasonable temperature profiles like \( T \sim r^{1/2} \), \( H \) rises with distance from the central star.

Because of the need to conserve angular momentum, gas and dust do not fall directly on the protostar, but fall parallel to the momentum vector into the disk. The turbulent disk transports mass inward and momentum outward and in this way allows accretion onto the protostar.

For the solar system the time of planetesimal formation is known fairly accurately from the evidence of extinct \(^{26}\text{Al}\) (half life 0.72 Myr) in primitive meteorites.

Micron-sized particles can coagulate and grow to millimeter size in \(~10^4\) years.

Growth beyond meter size is hindered as particles of this size experience strong gas drag and may therefore be swept into the host star in \(~100\) years.
Summary, part 2

Km-sized planetesimals have sufficient gravity to grow by gravitationally attracting other bodies.

For \( v \approx v_e \) (escape velocity) \( \frac{dM}{dt} \sim R^2 \), and for \( v \ll v_e \) \( \frac{dM}{dt} \sim R^4 \) (runaway growth).

The Earth can be made in \( 2 \times 10^7 \) – \( 10^8 \) years, but, in accordance with Chapter 4 of this lecture, the gaseous planets cannot be formed by accretion of solid planetesimals alone.

Once massive terrestrial protoplanets have formed, they will pump up each other’s velocities. Violent collisions may occur until the planets have at last reached orbits that are stable for billions of years.
Origin of solar systems
30 June - 2 July 2009
by Klaus Jockers (jockers@mps.mpg.de)
Max-Planck-Institut of Solar System Science
Katlenburg-Lindau

Part 6
Equilibrium condensation of a solar nebula
Equilibrium condensation of a solar nebula


and

Lodders, Katharina, talk given on Planet Formation Conference, Tuebingen, Germany, March 2009

Part of this was covered by Roberto Bugiolacchi’s lecture.
Chemistry in the disk:

The initial chemical state of the disk depends upon the composition of the gas and dust in the interstellar medium and subsequent chemical processing during the collapse phase.

Comets, chondritic meteorites, and in some sense the planets as well, are relics from the planetesimal-forming era of the solar system’s protoplanetary disk.

The composition of the Sun and of the carbonaceous chondrites (unprocessed meteorites) tells us what the original elemental composition of the disk was, but not the chemical compounds these elements formed.

*All that is in the planets (meteorites, comets) was in the disk, but not all that was in the disk is in the planets (meteorites, comets).*
The chemical evolution of interstellar matter as it is incorporated into planetesimals via the disk process is fundamental to our understanding of planet formation.

Therefore our purpose is now to study the chemical reactions that took place during the cooling of the solar nebula.

Near the protostar T > 2000K and chemical equilibrium holds. When chemical reaction times become comparable to the cooling time, chemistry becomes more complicated (*freeze-out temperature*).

Examples CO/CH\textsubscript{4} and N\textsubscript{2}/NH\textsubscript{3}. CO and N\textsubscript{2} dominate the inner solar nebula while in the cold outer nebula CH\textsubscript{4} and NH\textsubscript{3} would be favored.

But there are N\textsubscript{2} and CO ices on Pluto and Triton.

For the time being we will assume chemical equilibrium anywhere in the disk. This is a traditional view but nevertheless provides important insight.
Calculations of the thermodynamically stable forms of C, O and N at the time when icy condensates began to form at 5 AU. The mass accretion rate of the protoplanetary disk for these calculations is $10^{-7}$ $M_\odot$ per year.
Dissociation and molecule formation:

At temperatures below about 2000 K, the overwhelming majority of the atoms of every element are neutral, not ionic.

\[ \text{H}_2 \rightleftharpoons \text{H} + \text{H} \]

Le Chatelier’s principle: “Any system, whenever possible, responds to an externally applied stress so as to minimize the internal effects of that stress.”

With rising pressure the reaction moves to the left side.

Halving the volume of the gas with a piston at constant temperature increases the internal pressure by less than a factor of two.
Formation of diatomic molecules

---

**Table IV.5  Formation of Diatomic Molecules**

<table>
<thead>
<tr>
<th>Reaction</th>
<th>Reaction</th>
<th>Reaction</th>
<th>Reaction</th>
</tr>
</thead>
<tbody>
<tr>
<td>H + H → H₂</td>
<td>S + S → S₂</td>
<td>P + O → PO</td>
<td>Sc + O → ScO</td>
</tr>
<tr>
<td>C + O → CO</td>
<td>N + O → NO</td>
<td>P + N → PN</td>
<td>Zr + O → ZrO</td>
</tr>
<tr>
<td>H + O → OH</td>
<td>N + S → NS</td>
<td>P + P → P₂</td>
<td>Y + O → YO</td>
</tr>
<tr>
<td>C + N → CN</td>
<td>H + F → HF</td>
<td>C + P → CP</td>
<td>Si + H → SiH</td>
</tr>
<tr>
<td>C + S → CS</td>
<td>H + Cl → HCl</td>
<td>Si + O → SiO</td>
<td>Mg + H → MgH</td>
</tr>
<tr>
<td>O + O → O₂</td>
<td>H + N → NH</td>
<td>Si + S → SiS</td>
<td>Ca + H → CaH</td>
</tr>
<tr>
<td>S + O → SO</td>
<td>H + C → CH</td>
<td>Al + O → AlO</td>
<td>Si + F → SiF</td>
</tr>
<tr>
<td>C + C → C₂</td>
<td>H + S → SH</td>
<td>Ti + O → TiO</td>
<td>Mg + F → MgF</td>
</tr>
<tr>
<td>N + N → N₂</td>
<td>H + P → PH</td>
<td>V + O → VO</td>
<td>Sr + F → SrF</td>
</tr>
</tbody>
</table>

Note that we deal here with equilibrium in collision dominated regions. The excess energy can always be transferred by collision to another particle and there is no restriction about the number of reaction products.
Formation of polyatomic molecules

Table IV.6  Formation of Polyatomic Molecules

<table>
<thead>
<tr>
<th>Reaction 1</th>
<th>Reaction 2</th>
<th>Reaction 3</th>
</tr>
</thead>
<tbody>
<tr>
<td>H + OH → H₂O</td>
<td>O + CS → COS</td>
<td>H + CH₂ → CH₃</td>
</tr>
<tr>
<td>H + SH → H₂S</td>
<td>O + CO → CO₂</td>
<td>H + CNO → HCNO</td>
</tr>
<tr>
<td>PO + O → PO₂</td>
<td>H + PH₂ → PH₃</td>
<td>H + CP → HCP</td>
</tr>
<tr>
<td>H + CO → CHO</td>
<td>H + NH → NH₂</td>
<td>H + CH₃ → CH₄</td>
</tr>
<tr>
<td>H + CH → CH₂</td>
<td>H + CHO → CH₂O</td>
<td>CH₂ + CH₂ → C₂H₄</td>
</tr>
<tr>
<td>H + CN → HCN</td>
<td>O + CN → CNO</td>
<td>CH₃ + CH₃ → C₂H₆</td>
</tr>
<tr>
<td>H + PH → PH₂</td>
<td>CH + CH → C₂H₂</td>
<td>AlO + H → AlOH</td>
</tr>
<tr>
<td>S + CO → COS</td>
<td>H + NH₂ → NH₃</td>
<td></td>
</tr>
</tbody>
</table>
H₂, He, Ne, Ar

Hydrogen, helium, and neon chemistry. The top line shows equal H and H₂ pressures; below it, H₂ is the dominant temperature for H₂, Ne, Ar, and He are illustrated. Triple point (TP) and critical point (CP) of H₂ and I–liquid II pseudo-triple point (“TP”). The horizontal Tₘᵣᵦ is the microwave background temperature of lower temperatures, although not wholly impossible, (or natural) refrigeration.

1 bar = 0.987 atm
Cooling of O, C, N from 3000 K downward

Digression on famous German chemists:

From chemistry textbook (Hollemann-Wiberg: Lehrbuch der anorganischen Chemie, Berlin 1958):

*Synthesis of ammonia gas (Haber-Bosch):*

\[ 3 \text{H}_2 + \text{N}_2 \leftrightarrow 2 \text{NH}_3 + 22.1 \text{ kcal} \]

a: exothermal \( \rightarrow \) equilibrium shifts toward \( \text{NH}_3 \) if temperature gets lower.
   at room temperature: equilibrium fully at the side of \( \text{NH}_3 \).
   But reaction speed is very low. A catalyst is required.

b: volume shrinks \( \rightarrow \) equilibrium shifts toward \( \text{NH}_3 \) if pressure gets higher.

*Synthesis of gasoline (Fischer-Tropsch):*

\[ n \text{CO} + (2n + 1) \text{H}_2 \leftrightarrow C_n\text{H}_{2n+2} + n\text{H}_2\text{O} \]
\[ n \text{CO} + 2n \text{H}_2 \leftrightarrow C_n\text{H}_{2n} + n\text{H}_2\text{O} \]

Note: CO and \( \text{N}_2 \) have the same electronic structure.
Cooling of O, C, N from 3000 K downward

Cooling of an H–O–C–N mixture from about 3000 K gives rise to the following sequence of equilibrium reactions:

\[
\begin{align*}
N + N & \rightleftharpoons N_2 & \text{(IV.99)} \\
H + H & \rightleftharpoons H_2 & \text{(IV.100)} \\
C + O & \rightleftharpoons CO & \text{(IV.101)} \\
H + O & \rightleftharpoons OH & \text{(IV.102)} \\
H + OH & \rightleftharpoons H_2O & \text{(IV.103)} \\
CO + 3H_2 & \rightleftharpoons CH_4 + H_2O & \text{(IV.104)} \\
N_2 + 3H_2 & \rightleftharpoons 2NH_3 & \text{(IV.105)} \\
H_2O & \rightleftharpoons H_2O(s) \quad \text{(solid)} & \text{(IV.106)} \\
NH_3 + H_2O(s) & \rightleftharpoons NH_3 \cdot H_2O(s) & \text{(IV.107)} \\
CH_4 + 7H_2O(s) & \rightleftharpoons CH_4 \cdot 7H_2O(s) & \text{(IV.108)}
\end{align*}
\]

Fischer-Tropsch
Haber-Bosch
Methane hydrate = clathrate
Higher pressure favors CH$_4$, NH$_3$, as said in the chemistry textbook!
Figure IV.15  Carbon, nitrogen, and oxygen chemistry. Atomic H, C, N, and O combine to form OH, CO, and N₂ and thence stepwise to form H₂O, CH₄, and NH₃ as the temperature falls. The condensation processes are, in order, condensation of water ice I, partial conversion of ice I to ammonia monohydrate, conversion of all remaining water ice to the methane clathrate hydrate, and condensation of the leftover methane. b illustrates the equilibrium regions of dominance of the various compounds of C, N, and O. Note the region of thermodynamic stability of graphite at low pressures (< 10⁻⁷ bar).
Mole fraction versus pressure at constant temperature

1000 K is equilibrium temperature at 0.1 AU in our solar system.

The pressure is less constrained ~ $10^{-4}$ bar at 0.1 AU

Figure IV.16  Mole fractions of H, C, N, and O gases along an isotherm. The crossover points for ammonia and methane at 1000 K are emphasized. Note also the CO$_2$ and H abundances.
Magnesium and Silicon

$3000 < T < 4000 \text{ K}: \text{SiO and SiS. At } 2000 \text{ K}: \text{Mg, SiO, SiS, and Si.}$

$$\text{SiO} + 2\text{Mg} + 3\text{H}_2\text{O} \rightleftharpoons \text{Mg}_2\text{SiO}_4(s) + 3\text{H}_2$$  \text{(IV.118)}

$$\text{SiS} + \text{H}_2\text{O} \rightleftharpoons \text{SiO} + \text{H}_2\text{S}$$  \text{(IV.119)}

$$\text{Si} + \text{H}_2\text{O} \rightleftharpoons \text{SiO} + \text{H}_2$$  \text{(IV.120)}

$$\text{SiO} + \text{Mg}_2\text{SiO}_4(s) + \text{H}_2\text{O} \rightleftharpoons 2\text{MgSiO}_3(s) + \text{H}_2.$$  \text{(IV.121)}

Forsterite is an olivine (nesosilicate), enstatite a pyroxene (inosilicate)
Olivines (refractory, common in “primitive solids”)

\( \text{Fe}_2\text{SiO}_4 \) (fayalite)
\( \text{Mg}_2\text{SiO}_4 \) (forsterite)
Most frequently we find mixtures of the two.

They are nesosilicate (nes = island, isolated tetrahedra)

Pyroxenes (more processed than olivines)

Pyroxenes (inosilicates) form single or double chains of tetrahedra.
As the oxygen atoms are shared between adjacent tetrahedra, the chemical formula is \( \text{SiO}_3 \) or \( \text{Si}_2\text{O}_6 \).
\( \text{MgSiO}_3 \) or \( \text{Mg}_2\text{Si}_2\text{O}_6 \) (enstatite, as compared to forsterite there is one more \( \text{SiO}_2 \) “anion” added)
Fig. 26.2  (a, b) Linkage of $\text{SiO}_4^{4-}$ tetrahedra in orthosilicates with (a) isolated tetrahedra as in olivine and (b) groups of two linked tetrahedra as in lawsonite. (c) Linkage of $\text{SiO}_4^{4-}$ tetrahedra in ring silicates with rings of six as in beryl. (d, e) Linkage of $\text{SiO}_4^{4-}$ tetrahedra in chain silicates with (d) a single chain as in pyroxenes (translational repeat is indicated) and (e) a double chain as in amphiboles (translational repeat is indicated). (f) Linkage of $\text{SiO}_4^{4-}$ tetrahedra in sheet silicates with an infinite two-dimensional sheet. (g) Linkage of $\text{SiO}_4^{4-}$ tetrahedra in framework silicates as in tridymite (with the c-axis vertical). Oxygen atoms are in the corners of tetrahedra; silicon atom is in the center of tetrahedra.
“Weathering” of enstatite

\[ 4\text{MgSiO}_3 + 2\text{H}_2\text{O} \rightleftharpoons \text{Mg}_4\text{Si}_4\text{O}_{10}(\text{OH})_4. \]  

Through such reactions \textit{water may be retained} in a chemically bound state in minerals such as talc, serpentine, or chlorite at temperatures far too high for direct condensation of water or ice.
Iron

The only gaseous iron species of any consequence is atomic iron vapor.

At ~1500 K this vapor condenses directly to metallic iron. At total pressure of several bars iron condenses above its melting temperature, and liquid iron is the stable phase.

At lower temperatures:

Fe(s) + H₂O ←→ FeO(s) + H₂

Mg → Fe substitution:
By an interesting coincidence of nature the Fe²⁺ ion is an almost perfect match for Mg²⁺ in both ion charge and ionic radius. In addition, Mg, Si, and Fe have very similar cosmic abundance. Complete oxidation of Fe to FeO (below ~400 K) provides SiO₂, MgO, and FeO in nearly equimolar amounts, precisely what is needed to make MgFeSiO₄. Olivines are minerals of this stoichiometry and structure, regardless the Mg:Fe ratio: Fe₂SiO₄: fayalite, Mg₂SiO₄: forsterite.

A similar substitution is possible in enstatite, but pure FeSiO₃ (ferrosilite) is unstable with respect to decay into FeO + SiO₂.
It is possible for the silicates to accommodate all the iron as FeO compounds in solid solutions before the temperature drops below 490 K. Thus metallic iron must disappear by this temperature, and pure wüstite (FeO mineral) never forms.

The oxidation of Fe metal involves incorporation of FeO in silicates by

\[
\text{Fe(s)} + \text{H}_2\text{O} + \text{MgSiO}_3(s) \rightarrow \text{FeMgSiO}_4(s) + \text{H}_2
\]

\[
\text{IV.131}
\]

\[
\text{FeMgSiO}_4(s) + \text{MgSiO}_3(s) \rightarrow (\text{Mg, Fe})_2\text{SiO}_4(s) + (\text{Mg, Fe})\text{SiO}_3(s).
\]

\[
\text{IV.132}
\]

\[
(\text{Fe, Mg})\text{SiO}_3(s) + 2\text{H}_2\text{O} \rightarrow \text{SiO}_2(s) + (\text{Fe, Mg})_3\text{Si}_2\text{O}_5(\text{OH})_4.
\]

\[
\text{IV.134}
\]
Figure IV.18  Iron oxidation and FeO entry into silicates. Freshly condensed magnesium silicates (mostly enstatite) are virtually devoid of FeO. Most oxidation of metal occurs below 600 K, and metal is exhausted near 490 K.
Below 490 K the ferromagnesian silicates olivine (now dominant) and pyroxene (now minor) can be altered by reaction with water vapor in much the same way as enstatite alteration produces talc. Production of serpentine can be written as follows:

\[
(\text{Fe, Mg})\text{SiO}_3(s) + 2\text{H}_2\text{O} \quad \text{pyroxene}
\]

\[
\rightleftharpoons \text{SiO}_2(s) + (\text{Fe, Mg})_3\text{Si}_2\text{O}_5(\text{OH})_4. \quad \text{quartz} \quad \text{serpentine}
\]

Graphite cannot form as single mineral but C can be implemented into liquid Fe as solution.
Figure IV.17  Silicon, magnesium, and iron reactions. The major high-temperature event is the reaction and condensation of SiO, Mg, Fe, H$_2$O, and other gases to produce metallic iron and the magnesium silicates. The endpoint in the oxidation of metallic iron to form FeO-bearing silicates and the point of conversion of ferromagnesian silicates to serpentine are also given.
Sulfur

Near 2000 K, sulfur is found largely as SiS and the SH radical, although traces of SO, COS and CS are also present. Solid sulfides of silicon are unstable, hence SiS does not condense. Instead SiS (and SH) are converted to H₂S during cooling.

The first and most important sulfide to form is FeS (troilite), which is produced by corrosion of metallic Fe:

$$\text{Fe(s)} + \text{H}_2\text{S} \rightleftharpoons \text{H}_2 + \text{FeS(s)}, \quad (\text{IV.135})$$

FeS formation does not depend on pressure and is complete at 600K. As the sulfur abundance is less that the iron abundance, H₂S is fully removed from the gas.
Iron exhibits siderophile, lithophile and chalcophile geochemical traits simultaneously. Iron is by mass the second most important element in the terrestrial planets.

**Figure IV.19** Distribution of iron between its minerals. The fraction of total iron in the monatomic iron vapor, in solid metallic iron, in solid FeS, and in FeO solid solutions in silicates are shown from above the condensation point of metal down to 400 K.
Aluminum and Calcium

Calcium and aluminum oxides are very refractory substances and, therefore, among the earliest condensates to form during the cooling of solar material.

Their dominant gaseous species near 2000 K, AlO and Ca, react near 1800 K to produce a complex series of refractory oxide condensates that are very poor in Si and Fe. Corundum Al$_2$O$_3$, spinel MgAl$_2$O$_4$, perovskite CaTiO$_3$, and several Ca aluminosilicates such as gehlenite (Ca$_2$Al$_2$SiO$_7$) and anorthite (CaAl$_2$Si$_2$O$_8$) condense over a temperature range of only 200 K.

Below the condensation temperature of the magnesium silicates it is possible for CaO to enter extensively into the pyroxenes as the end member CaSiO$_3$ (wollastonite).
Olivine and pyroxene nomenclature. Pure ferrosilite is unstable (shaded region). High-temperature condensates lie deep in the diopside and enstatite corners of the dihefsen pyroxene quadrilateral. At lower temperatures, oxidation of metallic iron by water vapor produces bronzite and hypersthene, and alters much of the pyroxene to olivine of intermediate composition. Minerals close to wollastonite composition, which crystallize in the triclinic system, are found only in highly alkaline rocks and metamorphosed limestones. Pyroxenes found in meteorites generally lie in the three regions labeled augite, pigeonite, and orthopyroxene. It is common for all three of these compositions to coexist. See text for further explanation. (See Fig. IX.22.)

At much lower temperatures, Ca-bearing minerals may react with water vapor to produce hydroxyl silicate especially the very stable amphibole Ca$_2$Mg$_5$Si$_8$O$_{22}$(OH)$_2$, tremolite. It appears that tremolite is the first phase capable in retaining water in a cooling system of solar composition. But serpentinization of the ferromagnesian materials is a more important water trap, as Ca has a low abundance.
Sodium and Potassium

Atomic sodium and potassium vapor, Na and K, remain in the gas phase until the completion of the condensation of enstatite. At slightly lower temperatures it becomes possible for the alcali metal vapors to react with aluminium-bearing minerals to produce alkali aluminosilicates.

\[
\text{MgAl}_2\text{O}_4(s) + 13\text{MgSiO}_3(s) + 2\text{Na} + \text{H}_2\text{O} \\
\text{spinel} \quad \text{enstatite} \\
\Leftrightarrow 7\text{Mg}_2\text{SiO}_4(s) + 2\text{NaAlSi}_3\text{O}_8(s) + \text{H}_2. \\
\text{forsterite} \quad \text{albite} \\
\text{(IV.138)}
\]

Anorthite (CaAl_2Si_2O_8) can form solid solutions with both albite (NaAlSi_3O_8) and orthoclase (KAlSi_3O_8). Such aluminosilicate solid solutions are collectively called feldspars. The great mutual solubility of anorthite and albite occurs despite their superficially very different formulae. This is because replacement of Ca^{++} by Na^{+} combined with the replacement of Al^{3+} by Si^{4+} leads to no net change in electric charge, and the ionic size differences are not as large as that between Na^{+} and K^{+}.

The low-temperature solubility of orthoclase in the soda-lime feldspar is, however, not large, because of the large ionic size of potassium.
Figure IV.22 The feldspars. Natural feldspars form a single solid solution at elevated temperatures; however, assemblages equilibrated at low temperatures often contain a potassium-rich feldspar of approximate orthoclase composition coexisting with a plagioclase (albite–anorthite solid solution) phase. The large ionic size of K discourages extensive substitution of orthoclase into plagioclase. However, solar proportions of K, Na, and Ca (marked X in the diagram) contain so little K that the low-temperature solubility limit for K-spar in plagioclase is not exceeded. Because Ca is also an important component of pyroxenes, the ratio an/(ab + or) in the feldspars is shifted along the dashed line away from the anorthite corner. Meteoritic feldspars (see Chapter VIII) are generally found in the shaded area of oligoclase composition. Where differentiation and fractional crystallization processes have greatly enhanced the or:ab ratio, a second feldspar of nearly pure orthoclase composition (lightly shaded area near the or corner) can coexist with plagioclase.
Retention of potassium is of great importance to the terrestrial planets, because $^{40}$K decay by electron or positron emission is a very large source of heat, sufficient by itself to heat solid planetary material up to the melting point.

The half-life for $^{40}$K decay is $1.3 \times 10^9$ years, and thus both the abundance and the rate of decay of $^{40}$K were greater by a factor $2^{(4.5/1.3)} = 11$ at the time of origin of the solar system.

The $^{40}$K nuclide is a product of excessive silicon burning in supernova explosions; it introduces the possibility of planetary thermal evolution extending over periods of billions of years.

$^{40}$K decay provides us with a useful nuclear clock for dating geological events.
Nickel and Cobalt

The dotted line traces the locus of the composition of the metal during cooling from first condensation to the disappearance of the last metal at 490 K. When the dotted line lies within the $\alpha+\gamma$ two-phase region these two phases form an intricately intergrown structure.

Figure IV.23  The iron–nickel system. The stability fields of liquid metal, delta, gamma, and alpha iron are shown for 0.01 atm. Freshly condensed metal (1540 K) is slightly Ni-rich. The dotted line traces the metal composition during cooling of a solar-composition gas. Note that the metal enters the taenite-plus-kamacite two-phase region near 870 K and probably stays inside it through FeS condensation and substantial oxidation of iron. The last remaining metal phase is very Ni-rich, with 60% Ni (kamacite) composition. This metal phase disappears due to entry of nickel into sulfides such as pentlandite, (Fe, Ni)$_9$S$_8$, near 490 K.
Dissolved carbon in Fe-Ni alloys

Figure IV.24 Dissolved carbon in Fe–Ni alloys. The equilibrium concentrations of carbon in ppm by weight are given over the entire pressure–temperature range of stability of the free metal phases. Note the maximum in carbon content near the CO/CH₄ equal-abundance line.
**Table IV.7  Mineral Assemblages for Equilibrium Condensation of Major Elements**

<table>
<thead>
<tr>
<th>Step</th>
<th>Materials present</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>Ca$_2$Al$_2$SiO$_7$, MgAl$_2$O$_4$</td>
</tr>
<tr>
<td>2</td>
<td>List 1 + (Fe, Ni)</td>
</tr>
<tr>
<td>3</td>
<td>List 2 + MgSiO$_3$</td>
</tr>
<tr>
<td>4</td>
<td>(Fe, Ni) + MgSiO$_3$ + CaMgSi$_2$O$_6$ + plagioclase</td>
</tr>
<tr>
<td>5</td>
<td>List 4 + FeS</td>
</tr>
<tr>
<td>6</td>
<td>(Fe, Ni) + MgSiO$_3$ + plagioclase + FeS + tremolite</td>
</tr>
<tr>
<td>7</td>
<td>(Fe, Mg)$_2$SiO$_4$ + plagioclase + (Fe, Ni)$_3$S$_8$ + amphibole</td>
</tr>
<tr>
<td>8</td>
<td>Serpentine + plagioclase + sulfide + amphibole = rock</td>
</tr>
<tr>
<td>9</td>
<td>List 8 (rock) + H$_2$O</td>
</tr>
<tr>
<td>10</td>
<td>Rock + H$_2$O + NH$_3$ · H$_2$O</td>
</tr>
<tr>
<td>11</td>
<td>Rock + NH$_3$ · H$_2$O + CH$_4$ · 7H$_2$O</td>
</tr>
<tr>
<td>12</td>
<td>Rock + NH$_3$ · H$_2$O + CH$_4$ + 7H$_2$O + CH$_4$</td>
</tr>
<tr>
<td>13</td>
<td>Rock + NH$_3$ · H$_2$O + CH$_4$ · 7H$_2$O + CH$_4$ + Ar = rock + ice</td>
</tr>
<tr>
<td>14</td>
<td>List 13 (rock plus ice) + Ne</td>
</tr>
<tr>
<td>15</td>
<td>Rock + ice + Ne + H$_2$</td>
</tr>
<tr>
<td>16</td>
<td>Rock + ice + Ne + H$_2$ + He = everything (solar)</td>
</tr>
</tbody>
</table>

Note. Elements considered for this list: H, He, O, C, N, Ne, Si, Mg, Fe, S, Ar, Ca, Al, Na, Ni, and K. The total abundance of all other elements is less than the abundance of nickel.

1. “Firebrick (Schamotte)”
   Al$_2$O$_3$ corundum
   MgAl$_2$O$_4$ spinel (1 divalent, 2 trivalent cations). Another example:
   Fe$^{2+}$(Fe$^{3+}$)$_2$O$_4$ magnetite
   CaTiO$_3$ perovskite
   CaAl$_2$SiO$_7$ gehlenite
   CaAl$_2$Si$_2$O$_8$ anorthite (feldspar)
   NaAlSi$_3$O$_8$ albite (feldspar)
   albite↔anorthite: plagioclase (feldspar)

2. kamacite (α-iron), taenite (γ-iron)

3. MgSiO$_3$ enstatite (pyroxene)

4. CaMgSi$_2$O$_6$ diopside (pyroxene)

5. FeS troilite

6. Ca$_2$Mg$_5$Si$_8$O$_{22}$(OH)$_2$ tremolite (amphibole): first phase capable of containing water

8. (Fe,Mg)$_3$Si$_2$O$_5$(OH)$_4$ serpentine (water addition to pyroxenes)
Equilibrium condensation

1. Ca$_2$Al$_2$SiO$_7$, MgAl$_2$O$_4$
2. List 1 + (Fe, Ni)
3. List 2 + MgSiO$_3$
4. (Fe, Ni) + MgSiO$_3$ + CaMgSi$_2$O$_6$ + plagioclase
5. List 4 + FeS
6. (Fe, Ni) + MgSiO$_3$ + plagioclase + FeS + tremolite
7. (Fe,Mg)$_2$SiO$_4$ + plagioclase + (Fe, Ni)$_8$S$_8$ + amphibole
8. Serpentine + plagioclase + sulfide + amphibole = rock
9. List 8 (rock) + H$_2$O
10. Rock + H$_2$O + NH$_3$ · H$_2$O
11. Rock + NH$_3$ · H$_2$O + CH$_4$ · 7H$_2$O
12. Rock + NH$_3$ · H$_2$O + CH$_4$ + 7H$_2$O + CH$_4$
13. Rock + NH$_3$ · H$_2$O + CH$_4$ · 7H$_2$O + CH$_4$ + Ar = rock + ice
14. List 13 (rock plus ice) + Ne
15. Rock + ice + Ne + H$_2$
16. Rock + ice + Ne + H$_2$ + He = everything (solar)

Figure IV.26 Major element flow chart for equilibrium condensation. The sequence of reactions of gases (above the solid line) and condensates (below the solid line) in solar material cooling at equilibrium. A pressure of 10$^{-2}$ bar is assumed.
High-temperature equilibrium condensation

Equilibrium condensation is path independent, materials need not even to have ever been in the gas phase!

For a wide range of pressures, the condensation sequence does not change much.

Figure IV.27 High-temperature equilibrium condensation. The minerals formed by the major elements are sketched. The three different temperature-pressure profiles correspond to isobaric (constant pressure), isopycnic (constant density), and adiabatic (constant entropy) structures. Note the extreme similarity of the condensation sequences for these very different sets of conditions. Thus condensate composition can be used to place strong constraints on temperature, but some other source of information is needed to constrain the pressures in the nebula.
Low-temperature equilibrium condensation

Figure IV.28 Low-temperature equilibrium condensation. The ice condensates formed at equilibrium in a system of solar composition are shown. Note the eutectic melting of water ice I plus ammonia monohydrate at 173 K. Pure water ice melts at 273 K, just off the right-hand edge of the diagram.
Bulk density and total condensed mass

Tell me your density and I will tell you the temperature at which you condensed and your chemical composition!

Figure IV.29  Bulk density of equilibrium condensate and percentage of total mass condensed. The uncompressed density of the total condensate along the equilibrium condensation sequence as listed in Table IV.7 is shown by the solid line, and the total condensed mass as a percentage of total solar-composition mass is shown as the dashed line. The temperatures given are for a particular adiabatic model of the nebular structure (600 K at 10^{-4} bar), but the results look very similar for a wide range of assumptions regarding the pressure distribution.
Thermochemical equilibrium stability fields for condensed material in a solar composition medium. The species diagrammed are in primarily solid form below the lines and primarily gaseous at higher temperatures. The dashed lines are estimated temperature-pressure profiles for the circumsolar and circumjovian disks (Prinn 1993).

Free parameter along dashed lines is time!
Equilibrium condensation versus nonhomogeneous accretion

Figure IV.27 High-temperature equilibrium condensation. The minerals formed by the major elements are sketched. The three different temperature-pressure profiles correspond to isobaric (constant pressure), isopycnic (constant density), and adiabatic (constant entropy) structures. Note the extreme similarity of the condensation sequences for these very different sets of conditions. Thus condensate composition can be used to place strong constraints on temperature, but some other source of information is needed to constrain the pressures in the nebula.

Figure IV.31 Condensation in the nonhomogeneous accretion model. This figure should be compared with the equilibrium condensation diagram in Figs. IV.27 and IV.28.
Minimum mass primitive solar nebula

Ascribe approximate compositions to each of the planets based on their observed densities.

Calculate the mass of the primitive solar material required to make that planet.

Spread the material out over an annulus bracketing the present orbit of the planet.
It is unlikely that this method will introduce errors as large as a factor of two for any planet.

<table>
<thead>
<tr>
<th>Planet</th>
<th>Mass ((10^{26}))</th>
<th>(F^a)</th>
<th>(M_{\text{Solar}}(10^{26} \text{g}))</th>
<th>(r_{\text{ann}}(10^{13} \text{cm}))</th>
<th>(A_{\text{ann}}(10^{26} \text{cm}^2))</th>
<th>(\sigma = M/A(\text{g cm}^{-2}))</th>
</tr>
</thead>
<tbody>
<tr>
<td>Mercury</td>
<td>3.3</td>
<td>350</td>
<td>1,160</td>
<td>0.33–0.83</td>
<td>1.82</td>
<td>637</td>
</tr>
<tr>
<td>Venus</td>
<td>48.7</td>
<td>270</td>
<td>13,150</td>
<td>0.83–1.29</td>
<td>3.06</td>
<td>4300</td>
</tr>
<tr>
<td>Earth</td>
<td>59.8</td>
<td>235</td>
<td>14,950</td>
<td>1.29–1.89</td>
<td>6.00</td>
<td>2500</td>
</tr>
<tr>
<td>Mars</td>
<td>6.4</td>
<td>235</td>
<td>1,504</td>
<td>1.89–3.20</td>
<td>20.95</td>
<td>72</td>
</tr>
<tr>
<td>Asteroids</td>
<td>0.1</td>
<td>200</td>
<td>20</td>
<td>3.2–6.0</td>
<td>80.9</td>
<td>0.25</td>
</tr>
<tr>
<td>Jupiter</td>
<td>19,040</td>
<td>5</td>
<td>95,200</td>
<td>6.0–11.0</td>
<td>267</td>
<td>355</td>
</tr>
<tr>
<td>Saturn</td>
<td>5,695</td>
<td>8</td>
<td>55,560</td>
<td>11.0–21.5</td>
<td>1072</td>
<td>42.4</td>
</tr>
<tr>
<td>Uranus</td>
<td>870</td>
<td>15</td>
<td>13,050</td>
<td>21.5–36.8</td>
<td>2802</td>
<td>4.7</td>
</tr>
<tr>
<td>Neptune</td>
<td>1,032</td>
<td>20</td>
<td>20,640</td>
<td>36.8–52.0</td>
<td>4240</td>
<td>4.9</td>
</tr>
<tr>
<td>Pluto</td>
<td>0.1</td>
<td>70</td>
<td>7</td>
<td>52–70</td>
<td>6900</td>
<td>0.001</td>
</tr>
</tbody>
</table>

\(^a\)\(F\) is the factor by which the planetary mass must be multiplied to adjust the observed material to solar composition.

Mercury has high uncompressed density.

Complete condensation of iron but incomplete condensation of silicates?
Figure IV.33  Mass distribution in the solar nebula. A mean slope of $r^{-1.5}$ to $r^{-2.0}$ is suggested. The inner and outer edges appear sharply truncated. The inner edge is certainly due to the infall of matter from that region into the forming Sun. The outer edge may be due to a finite scale size of the original nebular condensation at the time of its last Jeans instability.
The following two projections are from Lodders, Katharina, talk given on Planet Formation Conference, Tuebingen, Germany, March 2009

Composition of Sun, solar nebula, and the planets as a whole is dominated by H and He

- Minimum mass nebula
  - $\sim 0.01 \, M_{\text{sun}}$
  - $\sim 10 \, M_{\text{jup}}$
  - $\sim 3200 \, M_{\text{Earth}}$

H, He

All other elements
(1.5% of total)
The diagram illustrates the masses of different components in Earth masses. The masses are divided into several categories:

- **H, He**
- **C, N, O**
  - "ices"
- **Silicates**
- **Metal**

The masses are represented in a logarithmic scale, with the values ranging from 0.06 to 1.0 on the x-axis and from 10 to 100 on the y-axis. The diagram shows the mass contributions of various elements and phases, with specific values indicated for each category.
Role of stable CO in the condensation process (this and the next slide).

Here we see the solar system case (C/O = 0.43) where oxygen was more abundant than carbon, leaving the remaining O free to become bonded into molecules which may condense into solids.


Figure 7.2  Temperature-pressure phase diagram illustrating stability zones of major solids in an atmosphere of solar composition (adapted from Salpeter 1974, 1977; Barshay and Lewis 1976). Above the dashed curve, gas phase CO is stable and essentially all C is locked up in this molecule. The most abundant gas phase reactants which lead to the production of solids are Fe, Mg, SiO and H₂O. The curved arrow indicates the variation in physical conditions which may occur in the outflow of a typical red giant. Magnesium silicates and solid Fe condense below the upper curve (● ● ●). At much lower temperatures, Fe is fully oxidized to FeO (below curve marked + + +) and may then be absorbed into silicates. Hydrous silicates such as serpentine are stable below the curve marked ○ ○ ○. Finally, water-ice condenses below the continuous curve.
Case for $C/O > 1$.

Rings and PAHs may form via reactions

$$C_2H_2 + H \rightarrow C_2H + H_2$$

$C_2H$ (C=CH) is ring segment.

$$C_nH_m \rightarrow C_nH_{m-1} + H_2$$

$$C_nH_{m-1} + C_2H_2 \rightarrow C_{n+2}H_m + H$$

PAH = polycyclic aromatic hydrocarbon


**Figure 7.3** Temperature–pressure phase diagram illustrating stability zones of solids in a carbon-rich atmosphere (adapted from Salpeter 1974, 1977; Martin 1978). Solar abundances are assumed except that the abundance of C is enhanced to exceed that of O by 10%. Above the dashed line, gas phase CO is stable and essentially all O is locked up in this molecule. Other important gas phase carriers of C are labelled outside the dotted curve. The curved arrow indicates the change in physical conditions associated with a typical outflow from a red giant. Solid carbon is stable in the region enclosed by the dotted curve. Condensation curves for SiC and Fe$_3$C are also shown.
Non-equilibrium processes in the cold outer regions of the protoplanetary disk:

Low gas densities $\sim 10^{-9}$ g cm$^{-3}$. In interstellar medium 40% of carbon is in dust and 10% in PAHs, most gas-phase C is in CO.

Possibly CO and N$_2$ never converted to CH$_4$ and NH$_3$ in the cold outer regions. NH$_3$ in cometary ices may be of interstellar origin.
Summary

• The outer part of the present solar system was never heated as strongly as its inner part. This should be reflected in present day chemical abundances found in pristine bodies like carbonaceous chondrites and comets, and in planets. A useful approach, in particular for the inner solar system, is to model the equilibrium cooling of the primordial “soup” (i.e. gas cloud) and the gradual freezing out of substances of decreasing volatility.
• Because of the extremely high volatility of H and H₂, in the outer solar system NH₃ should dominate over N₂ and CH₄ over CO, but it is possible that interstellar N₂ and CO never transformed into NH₃ and CH₄.
• At distances from the Sun > approx. 3-5 AU (beyond the “snow line”) icy planetesimals may exist and facilitate the formation of gas planets.
• Hydrated minerals like talc or serpentine may have helped to conserve water closer to the Sun than the “snow line” and bring it to Mars and Earth.
• Minimum solar nebula: Supplement the abundances found in planets with volatile elements until you reach the standard abundance of elements as found in the Sun and in carbonaceous chondrites to reconstruct the solar nebula as function of distance from the Sun. A mean slope $r_{\text{hel}}^{-1.5} - r_{\text{hel}}^{-2.0}$ is suggested.
• If in a solar system C is more abundant than O (carbon stars), all oxygen will be bound in CO and there will be no oxygen left to form H₂O and the silicate minerals we have on Earth. A completely different tar-like chemistry will result.
Dynamical and physical properties of extrasolar planets

presented as part of the lecture „Origin of Solar Systems“

Ronny Lutz and Anne Angsmann
July 2, 2009
Outline

• Introduction, detection methods  (Anne)
• Physical properties, statistics  (Ronny)
• Dynamical properties, atmospheres  (Anne)
• Habitability of exoplanets  (Ronny)
Introduction

- Extrasolar planets (exoplanets) are defined as objects orbiting a star which have masses below $13.6 \, M_{\text{Jupiter}}$

- more precise definitions (until now only applicable in our solar system): spherical shape and ability to clear its neighbourhood

- large ranges of possible properties - mass (factor 5800 in our solar system), distance from host star, temperature, eccentricity, composition, ...

- interesting aspects, e.g. time-dependent heating for strongly eccentric orbits
Detection methods for exoplanets

- **Astrometry**: changes in proper motion of host star due to the planet’s gravitational pull

- **Radial velocity**
  - magnitude
  - period
- **Gravitational microlensing**
  - advantage: may not miss faint stars or planets
  - disadvantage: planet’s orbital plane needs to be detected and mass of planet can be derived

- **Gravitational microlensing**
  - advantage: may miss faint stars or planets
  - disadvantage: planet’s orbital plane needs to be detected and mass of planet can be derived
Detection methods for exoplanets

- **Transit**: planet passes in front of host star and causes decrease in brightness
  - Photometric measurements indicate **size** and **orbital period** of planet (and possibly even **atmospheric elements**)
  - duration of transit yields **orbital inclination** → in combination with Doppler method, total **mass** of planet can be determined
  - **mean density** from \( M \) and \( R \)

- **Direct observation**

  Fomalhaut b, the first exoplanet to be imaged directly in visible light (2008)
Fomalhaut

HST ACS/HRC

a = 115 AU, R \sim R_{Jup}, M \sim 0.05 - 3 M_{Jup}

young system (\sim 100 - 300 million years)
HR 8799, a system with three planets, discovered in 2007 in infrared light with the Keck and Gemini telescopes (Marois et al, 2008)

- young star (~ 60 million years), planets recently formed: detected IR radiation from planets is internal heat
- orbital motion of planets (anticlockwise) confirmed in re-analyzed multiple observations back to 2004
Planet Detection Methods


Existing capability
Projected (10-20 yr)
Primary detections
Follow-up detections

n = systems; ? = uncertain

Detectable planet mass
Pulsars
10M_J
M_J
10M_E
M_E

Timing (ground)
White dwarfs
Binary
eclipses/other
Radial velocity
Astrometry
Radio

Dynamical effects

Microlensing

Astrometric
Photometric

Photometric signal
Imaging
Disks
Free floating
Transits

Resolved imaging
Detection of Life?
Timing residuals

Ground
Space

Optical
Radio

Self-accreting planetesimals
Magnetic superflares
Radio emission
Reflected/blackbody

Accretion on star

1
2

240 planets (205 systems,
of which 25 multiple)

4 planets
2 systems

4
3
2
1
19
5

19
5

Ground
Space
Atmospheres of exoplanets

• Theoretical models
  • Hot Jupiters
  • theoretical spectra

• Observations
  • methods of investigating atmospheric properties of exoplanets
  • the Earth‘s spectrum seen from space
  • the spectrum of HD 209458 b
  • day-night brightness differences at HD 189733 b
  • the spectrum of HD 189733 b
Theoretical models

- atmospheric composition depends on initial species, reactions and various other processes, and temperature
- scale height of atmosphere related to mass and radius of planet:

\[
H = \frac{k_B N_A T}{\mu g} \\
g = \frac{G M}{r^2}
\]

(\(k_B\): Boltzmann constant, \(N_A\): Avogadro number, \(\mu\): mean molar mass of atmospheric gas (Ehrenreich et al. 2005))

- the atmospheres of less dense planets extend further outwards → easier detection
- atmospheric escape: complex process, depending on balance between heating by UV radiation from host star and infrared cooling by certain molecules, e.g. \(\text{H}_3^+\) (Koskinen et al., 2008)
- hydrodynamically escaping atmosphere brings heavier elements to the hot upper layers; easier to detect than stable atmosphere
Theoretical models

Hot Jupiters:

• presumably tidally locked to their host star, thus heat transport towards the dark side should be investigated
• observations are mixed: some planets exhibit large day-night contrasts, others don't - more data needed
• outer radiative zones expected due to strong external heating by stars; inhibition of convection
• stable atmospheres possible, depending on mass of planet, stellar irradiation and atmospheric composition
• prediction of water by models (Grillmair et al., 2008)
• planet-spanning dynamical weather structures predicted
Theoretical spectra

• theoretical spectra for transmission (transit) and emission/reflection have been developed
• emission and reflection spectra: later
• transmission spectra (Ehrenreich et al., 2005):
  • Earth-sized terrestrial planets
  • challenging as the expected drop in intensity is only $10^{-7}$ - $10^{-6}$
  • models include only water vapour, $\text{CO}_2$, ozone, $\text{O}_2$ and $\text{N}_2$, regarding the wavelength range 0.2 - 2 $\mu$m
  • separate into three types:
    a) $\text{N}_2/\text{O}_2$-rich (Earth-like)
    b) $\text{CO}_2$-rich (Venus-like)
    c) $\text{N}_2/\text{H}_2\text{O}$-rich („ocean planet“)
• calculate absorption, Rayleigh scattering etc.
Theoretical spectra

Earth-like planet: $N_2$, $O_2$

Water only detectable when present in substantial amount above the clouds
Theoretical spectra

Venus-like planet: CO$_2$
Theoretical spectra

Ocean planet: N₂, H₂O
Theoretical spectra

- vegetation: „red edge“
- rapid increase in reflectance of chlorophyll at $\lambda \geq 700$ nm
Investigating atmospheric properties

- **transit**: determination of atmospheric chemical composition (absorption features, transit radii at different wavelengths)

- **secondary eclipse**:
  - infrared emission from planetary atmosphere
  - deduction of **effective temperature** of planet
  - observations are easiest in *infrared light* because of better ratio between emission of planet and star
  - but: combining measurements in different wavelengths yields more information → atmospheric effects!

- **between transits**:
  - analysis of atmospheric chemical composition in planet‘s reflection spectrum / scattered light by substracting secondary eclipse brightness
  - differences between dayside and nightside
Investigating atmospheric properties

- atmospheric structure and dynamics: start by looking at the basic properties of planets in our solar system

stratosphere: rising temperature because of UV light absorption by ozone/hydrocarbon products

troposphere: linear increase in temperature with depth caused by convection of heat from the surface/deep interior

Marley et al., 2008
Reflection spectrum of the Earth's atmosphere (Turnbull et al., 2006)
cirrus ice particles at 10 km altitude

cumulus water cloud at 4 km
Reflection spectrum of the Earth‘s atmosphere  (Turnbull et al., 2006)

Comparison with models leads to the following conclusions:

- the Earth‘s spectrum clearly differs from those of Mars, Venus, the gas giants and their satellites:
- strong water bands → habitable planet
- methane and large amounts of oxygen → either biological activity or very unusual atmospheric and geological processes
- clear-air and cloud fractions required in models → dynamic atmosphere; changes in albedo
- periodic changes due to rotation: maps of surface (land/ocean)
- but: washing out of surface signals by clouds
- visibility of seasonal changes?
Reflection spectrum of the Earth’s atmosphere
The spectrum of HD 209458 b

- Properties: $M = 0.685\ M_{\text{Jup}}$, $R = 1.32\ R_{\text{Jup}}$, semimajor axis: 0.047 AU, orbital period: 3.5 days
- first exoplanet detected in transit (2000)

![Graph showing the spectrum of HD 209458 b](image-url)
The spectrum of HD 209458 b

- Charbonneau et al. (2002) reported on the detection of sodium lines during transit of HD 209458 b
- less sodium than expected (absorption features should be three times stronger); discussion of depletion, clouds etc.
- detection of HI (Lyα), OI and CII in 2004 (Vidal-Madjar et al.)
- large amounts of these species are too far outside to be gravitationally bound to the planet (models) → hydrodynamic escape; escape rate ≥ $10^{10}$ g/s
- temperature inversion leads to water emission lines (Knutson et al., 2007)
- H$_2$ Rayleigh scattering (Lecavelier des Etangs et al., 2008)
- absorption by TiO (titanium oxide) and VO (vanadium oxide) as possible cause for temperature inversion (Désert et al., 2008); absorption lines not clearly identified yet
The spectrum of HD 209458 b

Richardson et al. (2007), Jul 8
Deming et al. (2005)
Knutson et al. (2007c)
Richardson et al. (2003)

- Default, $P_n=0.3$
- $\tau_{\text{opt.}} \sim 3$, $P_n=0.3$
- '' $P_n=0.4$
- '' $P_n=0.5$

three models with stratosphere (absorber in upper atmosphere) and slightly different redistribution parameters $P_n$

model without extra absorber in upper atmosphere
Day-night contrast at HD 189733b
(Knutson et al., 2007)

- Properties: $M = 1.14 \, M_{\text{Jup}}$, $R = 1.138 \, R_{\text{Jup}}$, semimajor axis: 0.03 AU, orbital period: 2.2 days
Day-night contrast at HD 189733b (8 μm) (Knutson et al., 2007)

- distinct rise in flux from transit to secondary eclipse
- increment of (0.12 ± 0.02)% in total amplitude
- comparison with secondary eclipse depth → variation in hemisphere-integrated planetary flux: $F_{\text{min}} = (0.64 \pm 0.07) \times F_{\text{max}}$
- flux peak at $16 \pm 6$ degrees before opposition
- secondary eclipse yields brightness temperature $T_{\text{eff}} = (1205.1 \pm 9.3)$ K
- additional variations imply the hemisphere-averaged temperatures $T_{\text{max}} = (1212 \pm 11)$ K and $T_{\text{min}} = (973 \pm 33)$ K
- creation of a basic map of brightness distribution by using a simple model comprised of twelve slices of constant brightness
Day-night contrast at HD 189733b

(Knutson et al., 2007)

no extreme day-night difference: redistribution by atmosphere

offset of brightest spot from substellar point indicates presence of atmospheric winds
Day-night contrast at HD 189733b (24 μm) (Knutson et al., 2009)

- very similar findings at 24 μm (wavelength corresponding to atmospheric region with different pressure)
- circulation must be very similar in both regions
- only small differences in temperature between layers probed by 8 μm and 24 μm → no convection at these altitudes
- efficient transport of heat from day- to nightside by atmospheric winds at both probed altitudes
- the atmosphere of HD 189733b can be described accurately with models with no temperature inversion and water absorption bands, as opposed to HD 209458b
The dayside emission spectrum of HD 189733b (Grillmair et al., 2008)

Dayside emission spectrum of HD 189733b

Spectroscopy (this paper)
Photometry (Charbonneau et al. 2008)

$P_n = 0.3, \kappa_e = 0.0$
$P_n = 0.1, \kappa_e = 0.0$
$P_n = 0.15, \kappa_e = 0.035$

„water bump“: signature of vibrational states of water vapour
The dayside emission spectrum of HD 189733b (Grillmair et al., 2008)

- water bump, flux ratio at 3.6 and 4.5 μm and decrease of planet/star flux ratio below 10 μm indicate presence of water vapour (water also found in transmission)
- significant differences to previous observations → dynamical weather structures in the upper atmosphere which change the spectrum?
- comparison with models indicates weak heat redistribution to nightside
- but: nightside temperature is high, maybe internal energy source
- heat redistribution might depend on atmospheric depth; three-dimensional models necessary

- strong indications for H₂O, CO₂ and CO in transmission spectrum (Swain et al., 2009)
The dayside emission spectrum of HD 189733b (Swain et al., 2009)
Summary (Part 3)

- atmospheres of exoplanets are expected to display a large range of possible properties
- investigation of atmospheres in transit/secondary eclipse
- theoretical spectra resulting from models reproduce the Earth’s atmospheric spectrum quite well
- various elements have been detected in atmospheres of exoplanets, in transmission as well as in reflection
- day-night contrasts can be measured
- comparison with models is very helpful in the investigation of atmospheres
Theoretical models
Formation of atmospheres

- atmospheric composition and evolution: formation of atmospheres in three possible ways (Elkins-Tanton et al., 2008):
  - capture of nebular gases
  - degassing during accretion
  - degassing from tectonic activity

- **low-mass terrestrial planets** do not have sufficient gravity to capture nebular gases

- in the **inner solar system**, nebular gases may have dissipated already when final planetary accretion takes place

- hints for composition of planetesimals come from meteorites: chondrites (water contained as OH) and achondrites (very low water content)
Theoretical models
Formation of atmospheres - chondritic material

Chondritic material alone:
• water and iron react until the water reservoir is exhausted
• release (outgassing) of hydrogen to the atmosphere
• some non-oxidized iron remains in the surface
• very rare cases: all iron oxidized before water content depleted; then also release of water to the atmosphere

Chondritic material with added water:
• assumption of an amount of water exactly sufficient to oxidize all the iron
• same implications for the atmospheric composition as in first model (only hydrogen degassed)
• no metallic iron remaining
Theoretical models
Formation of atmospheres - achondritic material

Achondritic material alone:
• accretion of a protoplanet with mantle and core; silicate mantle fully melted (magma ocean)
• when cooling down, part of the water is trapped inside the solidifying mantle minerals

Achondritic material with added water:
• similar to preceding case, but with additional volatiles available in the magma ocean phase
Dynamical and physical properties of extrasolar planets

presented as part of the lecture “Origin of Solar Systems”

Ronny Lutz, Anne Angsmann  2009-07-02
Outline

• Part 1 (Anne):
  • Introduction, Detection methods

• Part 2 (Ronny):
  • Physical properties, Statistics

• Part 3 (Anne):
  • Dynamical properties, Atmospheres

• Part 4 (Ronny):
  • Habitability of exoplanets
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Physical properties - Statistics

All Catalogs
update: 23 June 2009

All Candidates detected

- Candidates detected by radial velocity or astrometry
  update: 23 June 2009
  353 planets
  278 planetary systems
  327 planets
  34 multiple planet systems

- Transiting planets
  update: 23 June 2009
  59 planetary systems
  59 planets
  0 multiple planet systems

- Candidates detected by microlensing
  update: 19 September 2008
  7 planetary systems
  8 planets
  1 multiple planet system

- Candidates detected by imaging
  update: 24 November 2008
  9 planetary systems
  11 planets
  1 multiple planet system

- Candidates detected by timing
  update: 25 November 2008
  4 planetary systems
  7 planets
  2 multiple planet systems

exoplanet.eu
some “exo”-world records...

- lowest mass → PSR 1257+12b: 0.02M_E (pulsar host); Gl 581e: 2M_E (M-dwarf host)
- highest mass → HD 43848b: 25M_J = 8065M_E (G-dwarf host)
- shortest period → CoRoT-7b: 0.85d (K0-dwarf host)
- longest period → Fomalhaut b: 320000d (A3-dwarf host)
- closest to us → eps Eridani b: 3.2pc (K2-dwarf host)
- lightest host star → 2M1207: M8-dwarf with 0.025M_S
- most massive host star → HD 13189: K2-giant with 4.5±2.5M_S
- most planets in a system → 55 Cnc with five planets (G8-dwarf host)

\[ 1 \text{ M}_J \approx 318 \text{ M}_E \]

Physical properties of exoplanets
**Physical properties - Statistics**

- **planet vs. brown dwarf** → deuterium burning limit $13.6M_J$
  
  → formation process (core accretion vs. core collapse)

- BD desert at $\approx 15-60M_J$

- **terrestrial (solid) planet vs. (gaseous) giant planet** → $10M_E$?, different formation?

---

Physical properties of exoplanets
Physical properties - Statistics

yellow circle: RV
red circle: transit
blue square: microlensing
purple square: pulsar timing
circle size prop. to orbital period

Fig.1

Physical properties of exoplanets
Physical properties - Statistics

- biases due to detection method and survey characteristics
  → mainly giant planets (RV)
  → mainly solar-like hosts
- well defined samples:
  (volume or magnitude limited)
  CORALIE, HARPS, Lick/Keck/AAT
  >1200 FGKM stars (main sequence)

stars with giant planets m > 0.2 $M_J$ and a < 0.1AU: $\approx$ 1%
stars with giant planets m > 0.2 $M_J$ and a < 5.0AU: $\approx$ 6%

Fig.2
Physical properties - Statistics

- mass distribution
  → steep rise towards lowest masses, but also obs. bias toward lowest masses
Physical properties - Statistics

- period distribution and semimajor axis distribution
  - connected via 3rd Kepler: \( M_{\text{tot}} P^2 = a^3 \), e.g. 1yr \( \rightarrow \) 1AU, 10d \( \rightarrow \) 0.09AU
  - period gap (no observational bias)
  - peak at 3d or 0.04AU (a result of migration?)
  - peak at \( \approx 250 \text{d} \) or \( \approx 0.8 \text{AU} \)

Fig.6
red: \( m \sin i < 0.75M_J \)
blue: \( m \sin i < 21M_E \)

Fig.7

Fig.8

Physical properties of exoplanets
Physical properties - Statistics

- mass vs. period

  → lack of massive planets on short period orbits with \( m > 2M_J \) and \( p < 100d \) (0.4AU)

  → lack due to mass transfer?

  → evaporation?

  → infall?

  → no obs. bias

  → short period peak due to low mass planets

  → no difference between single hosts and binary hosts

Fig. 9
**Physical properties - Statistics**

- **eccentricity**
  - circularization for $p < 6d$
  - median of 0.24 for $p > 10d$ (0.09AU)
  - no obvious difference in planet and stellar binary populations
  - Solar System: very low eccentricities

![Figure 10](image1.png)

![Figure 11](image2.png)

![Figure 12](image3.png)
Physical properties - Statistics

- host star metallicity
  - probability of finding a giant planet increases with host star metallicity
  - no such trend for low-mass (solid) planets, but bad statistics so far
  - metallicity excess due to external origin or conditions during formation?
  - no trend with semimajor axis

![Graph showing frequency of planets vs. [Fe/H]](image)

**Fig.13**

![Graph showing median [Fe/H] vs. semimajor axis (au)](image)

**Fig.14**

Physical properties of exoplanets
Physical properties - Statistics

- host star mass
  - probability for finding a planet is higher for higher mass host stars
  - observational bias due to sample limitations?
  - O,B,A stars too short lived?; 2\textsuperscript{nd} generation planets around evolved stars?
  - low mass planets seem to be more common around M-dwarfs
  - no correlation with semimajor axis

Fig.15

Fig.16

Physical properties of exoplanets
Physical properties - Statistics

- **multiple-planet systems**
  - high eccentricities for large orbits
  - mean motion resonances
  - orbital parameters seem to be indistinguishable from single planet systems
  - **problems**: low amplitude RV signals from long period planets, which may easily be absorbed in a single-planet Keplerian orbit solution...

horizontal lines: periapse and apoapse in eccentric orbits
numbers: minimum masses in Jupiter masses
Summary (Part 2)

- 353 exoplanets, most are giants, bias due to detection method
- earth-mass planets will be found in the near future
- steep rise towards lower masses
- period gap
- no close-in massive planets
- circularization for close orbits
- whole range of eccentricities for large orbits
- metallicity excess
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Habitability of exoplanets

- habitability $\rightarrow$ life, biology, chemistry
- exoplanets $\rightarrow$ astronomy, geology

 Astrobiology

Life (as we know it) is based on carbon and liquid water:

main building blocks: H, C, O, N, (P, S), liquid H$_2$O
Concept of the habitable zone (HZ)

- **HZ**: distance from a star, where liquid water can exist on the surface
  - determined by $T_S$
  - definition neither necessary nor sufficient
  - habitability does NOT imply that an environment is inhabited
- **CHZ**: region in which liquid water can be maintained throughout most of the star's lifetime
  - $L$ of the host star increases with time $\Rightarrow$ HZ moves outwards

---

Habitability of exoplanets
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Habitability of exoplanets
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- **GHZ**: habitable regions in the galaxy
  - thin disk excluding the innermost and outermost parts

---

Habitability of exoplanets
Concept of the habitable zone (HZ)

- boundaries for an earth-like rocky planet:
  - inner edge: water-loss, H-escape, runaway greenhouse
  - outer edge: condensation of atmospheric CO$_2$, formation of CO$_2$ clouds, increased albedo, runaway glaciation

![Graph showing the habitable zone](image)

**Fig. 18**

solar system: HZ $\approx$ 0.95 - 1.67 AU
CHZ $\approx$ 0.95 - 1.15 AU
**Biosphere of a habitable planet**

**Earth's atmosphere:**  IR-spectrum

- **continuum** $\rightarrow T_{\text{eff}}, T_\odot$ (255K, 288K)
- $\text{CO}_2$
- $\text{H}_2\text{O}$-vapour $\rightarrow$ strongest greenhouse gas
- $\text{O}_3 (\rightarrow \text{O}_2)$
- $\text{CH}_4$
- $\text{N}_2$
- chlorophyll

\[ \exists/\x_6/\x_{15}/\x_{13}/\x_{11}/\x_{12}/\x_{16}/\x_{14}/\x_{16}/\x_9/\x_{15}/\x_{18}/\x_9/\x_4/\x_9/\x_{12}/\x_4/\x_6/\x_{17}/\x_4/\x_3/\x_{16}/\x_{17} \]

**Fig.20**

Reduction of $\text{CO}_2$, increase of $\text{O}_2$

Biomass changes the atmosphere
Biosphere of a habitable planet

Fig. 21
Climate of a habitable planet

Earth's Energy Budget

Incoming solar energy 100%

- Reflected by atmosphere 6%
- Reflected by clouds 20%
- Reflected from Earth's surface 4%
- Absorbed by atmosphere 16%
- Absorbed by clouds 3%
- Conduction and rising air 7%
- Carried to clouds and atmosphere by latent heat in water vapor 23%
- Absorbed by land and oceans 51%
- Radiated directly to space from Earth 6%
- Radiated to space from clouds and atmosphere 64%

Various factors can disturb this budget and therefore the climate → the habitability!

Fig. 22

Surface emits in thermal IR

Habitability of exoplanets
Climate of a habitable planet

feedback processes affect the climate:

• positive → destabilizing due to amplification of perturbations
• negative → stabilizing response to perturbations

positive feedbacks:  
snow / ice cover → runaway glaciation  
atmospheric water vapour → runaway greenhouse

negative feedbacks:  
carbonate-silicate-cycle  
\( T_S \leftrightarrow F_{IR,\text{out}} \) interaction
Factors influencing habitability

**planetary properties:**

- climatic stability
- albedo → cloud cover / ice+snow cover / ocean to land ratio
- presence of an atmosphere → protection from UV and impacts / pressure for liquid water
- atmospheric composition → volatiles / albedo
- geological activity → plate tectonics / volcanism / internal heat
- magnetic field → protection from charged particles
- rotation → dynamo / MF
- obliquity → moderate seasons
- mass → >0.3M$_E$ → maintain an atmosphere / activity
  - <10M$_E$ → prevent massive H-He atmosphere
- interaction with other planets → orbital stability
- eccentricity of the orbit → climatic stability
Factors influencing habitability

external properties:

- host star \(\rightarrow\) type

Fig. 23
Factors influencing habitability

**external properties:**

- **host star → type**
- **presence of giant planets → destabilizing: migration through HZ (inner giants)**
  
  stabilizing: impact shielding / orbital stability due to outer giants ?
- **presence of a satellite → Moon stabilizes Earth's obliquity (23.4°)**
- **stability around binary / multiple hosts ?**
- **impacts of comets / meteorites → positive: delivery of volatiles / water ice / (and life?)**
  
  negative: destroying biospheres
- **tidal locking / synchronous rotation → extreme hemispheres**
Habitable exoplanets?

**examples:**

- *Gliese 876d*: 7.5$M_\oplus$ around 0.32$M_\odot$ M-dwarf (0.02 AU) → too hot

- *OGLE-05-390L b*: 5.5$M_\oplus$ at 2.1AU → too cold
Habitable exoplanets?

**Examples:**

- **Gliese 581**: 0.31M$_S$ M-dwarf, 3200K, 0.014L$_S$
  
  b: 15.6M$_E$ at 0.04 AU → outside HZ
  
  c: 5M$_E$ at 0.07 AU → only habitable for high albedo → high cloud cover needed
  
  d: 8.3M$_E$ at 0.25 AU → best candidate if M stays below 10M$_E$ (i unknown)
  
  e: 1.97M$_E$ at 0.03AU → outside HZ, lowest mass known so far (17.06.2009)

Habitability of exoplanets
100,000 stars simultaneously, 1.4m primary mirror, launched in Mar. 2009
Space missions - TPF (NASA)

- indefinitely postponed, IR-Interferometer array or optical telescope

TPF-I

TPF-C

Figs.: nasa.gov

Habitability of exoplanets
6 IR telescopes (>3m) as Interferometer, launch 2015 or later
• definition of a “habitable zone”
• atmospheric features as biomarkers
• various factors (planetary and external properties) can influence habitability
• Gliese 581 system as most promising candidate
References

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- www.exoplanet.eu

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Figures:
Fig.1: Udry&Santos (2007)  Fig.13: Udry&Santos (2007)
Fig.2: Mason (2008)  Fig.14: Mason (2008)
Fig.3: Udry&Santos (2007)  Fig.15: Johnson (2009)
Fig.4: Johnson (2009)  Fig.16: Mason (2008)
Fig.5: Udry&Santos (2007)  Fig.17: Johnson (2009)
Fig.6: Udry&Santos (2007)  Fig.18: Kasting et al. (1993)
Fig.7: Mason (2008)  Fig.19: shayol.bartol.udel.edu/~rhdt/diploma/lecture_9
Fig.8: Mason (2008)  Fig.20: astrobiology.arc.nasa.gov
Fig.9: Udry&Santos (2007)  Fig.21: markelowitz.com/exobiology.htm
Fig.10: Udry&Santos (2007)  Fig.22: Image courtesy: NASA's ERBE program
Fig.11: Johnson (2009)  Fig.23: www.atlasoftheuniverse.com
Fig.12: Mason (2008)  Fig.24: Selsis et al. (2007)
Biosphere and climate of a habitable planet

the greenhouse effect:

- raise of $T_S$ because of greenhouse gases ($H_2O$-vapour, $CO_2$, ...)

  ↓

- transparent in UV, V but absorbing in IR

  ↓

- emitted thermal IR radiation is re-emitted to the surface

without the greenhouse effect, our Earth would be frozen!

→ in a positive feedback, this effect may become runaway!
the carbonate-silicate-cycle:
- regulates the amount of atmospheric CO₂
- removed by silicate weathering
  - bound to carbonate rocks
  - subducted to high T, P (plate tectonics)
  - unbound again
- added by volcanic activity

negative feedback due to the dependence of the weathering rate on $T_S$ (liquid water):

$T_S$ rises → CO₂ concentration falls
$T_S$ falls → CO₂ concentration rises

→ stabilization of the climate

(in humans would not interact...
Atmospheric Absorption Bands

- Percent
- Wavelength (μm)
- Total Absorption and Scattering
- Water Vapor
- Carbon Dioxide
- Oxygen and Ozone
- Methane
- Nitrous Oxide
- Rayleigh Scattering

HABITABILITY OF PLANETS
Dynamics of comets and the origin of the solar system

Jean-Baptiste Vincent

Max-Planck-Institut für Sonnensystemforschung

Origin of solar systems - 30/06/2009
Introduction

Because of their random apparitions, comets were considered as atmospheric events during many centuries until Tycho Brahe proved that comets are celestial objects by measuring parallaxes in 1577.

The study of comet dynamics started then with E. Halley

1705 : Edmund Halley calculates the period of a comet
1758 : Observations confirm Halley’s predictions
Comets follow Keplerian orbits.

=> For any observed comet, we can calculate its orbital parameters and estimate its orbit.

This will allow us to sort the comets into different groups, the study of these groups leading to important insights on the early ages of our solar system.
Classification

The classification is mainly based on the orbital parameters of the comets like *orbital period* or *aphelion distance*.

We distinguish three main groups:

1. short period comets
2. long period comets
3. non-periodic comets

- Jupiter Family Comets (JFC)
- Halley Type Comets (HTC)
- Main Belt Comets (MBC)
Short Period Comets

<table>
<thead>
<tr>
<th>Period</th>
<th>&lt; 200 years</th>
</tr>
</thead>
<tbody>
<tr>
<td>a</td>
<td>&lt; 100 AU</td>
</tr>
<tr>
<td>i</td>
<td>± 10°</td>
</tr>
</tbody>
</table>

1/5 of the catalogued comets

Coming mainly from the Kuiper Belt region (beyond Neptune’s orbit). Orbits affected by perturbations from the giant planets (especially Jupiter)

Examples: 1P/Halley (p = 76 yrs), 2P/Encke (p = 3.3 yrs)
Tisserand’s Parameter

Tisserand’s parameter is a dynamic number which is conserved (or only slightly modified) during an encounter between two objects in a restricted 3-bodies system.

\[ T_J = \frac{a_J}{a} + 2 \varphi \left( \frac{a}{a_j} (1 + e^2) \varphi \cos(i) \right) \]

- \( a_j \) = semi-major axis of Jupiter = 5.20 AU
- \( a \) = semi-major axis of the comet
- \( e \) = eccentricity
- \( i \) = inclination
Short Period Comets

JFC:
- $2.5 < T_J < 3$
- Period < 20 years
- Aphelion ~ Jupiter distance from the Sun (5.2 AU)
- 2P/Halley, 9P/Tempel 1

HTC:
- $T_J < 2$
- $20 < \text{Period} < 200$ years
- Aphelion ~ Neptune distance from the Sun (30 AU)
- 1P/Halley, 8P/Tuttle
Short Period Comets

The Tisserand’s parameter is very different for JFC and HTC. Besides, some of the HTC have a retrograd orbit which can not be explained if they were formed were they are now.

HTC are believed to be long period comets captured by the giant planets and trapped in the inner solar system, whereas JFC were “born” as short period comets.
Main belt comets?

Asteroids pretending to be comets?

or

Comets pretending to be asteroids?
Long Period Comets

<table>
<thead>
<tr>
<th>Period</th>
<th>&gt; 200 years</th>
</tr>
</thead>
<tbody>
<tr>
<td>$a$</td>
<td>&lt; $10^5$ AU</td>
</tr>
<tr>
<td>$i$</td>
<td>± 90°</td>
</tr>
</tbody>
</table>

4/5 of the catalogued comets

Examples: Hale-Bopp, Comet McNaught

None of these parameters can be explained by simple gravitational interaction of short period comets with giant planets.
Long Period Comets

Distribution of $1/a$ shows a peak around $5.10^{-3}$, typical of orbits which have not been affected by interactions with the giant planets.

$\Rightarrow$ Most of the comets enter the inner solar system for the first time.

This let Oort to conclude (1950) about the existence of $\sim 10^{11}$ comets orbiting in a spherical “cloud” at distances of $10^4$-$10^5$ AU. This structure, called the Oort cloud, is the source of the long-period comets that appear today.
Non-periodic comets

These comets have an eccentricity \( \geq 1 \). Their orbit is a parabola or an hyperbola, thus they do not come back after the perihelion passage.

Spectrometric measurements have shown that they come from the solar system (same isotope ratio).

We expect similar comets coming from other solar system (~6/century) but unfortunately none has been observed so far.
Summary

- Short Period comets come from the Kuiper belt (30 – 55 AU)
- Long period & non-periodic comets come from the Oort cloud (50 000 AU)
- as a reference: $\alpha$ – Centauri (277 600 AU)
We have a problem...

Comets show a $^{12}\text{C}/^{13}\text{C}$ ratio \(\sim 100\), close to the solar system value (89) and quite different from the interstellar one (\(\sim 30-50\))

\[\Rightarrow\text{comets must have formed in the solar system}\]

Comets share a similar composition,

\[\Rightarrow\text{they must have formed in the same region of the early solar system, more or less beyond Neptune’s orbit.}\]
We have a problem...

But...

=> radius of accretion disk $\sim$ 200-1000 AU
=> radius of Oort cloud $\sim$ 50 000 AU

Comets come from the solar system, but cannot be formed where they are now... we need a dynamical model.

=> The “Nice” Model
The “Nice” Model

R. Gomes, H. F. Levison, K. Tsiganis & A. Morbidelli
The “Nice” Model

Still not perfect but best model available so far.

Main scenario:
• Original solar system much more compact than now, with the four giant planets orbiting between 5.45 and 14.2 AU, surrounded by a disk of planetesimals at 30-35 AU.
( now Jupiter is at 5.2, Neptune at 30.1, and the Kuiper Belt between 30 and 50 AU)
The “Nice” Model

- Planetesimals are dragged *inwards* by Neptune, which move then *outwards* to keep the angular momentum of the Solar system constant.

- The same process continues as the icy bodies reach Uranus and Saturn.

- The situation is different for Jupiter. As its gravitational force is much stronger, most of the small bodies are accelerated and ejected out of the inner solar system, which causes Jupiter to move *inwards* to balance the loss of angular momentum.
The “Nice” Model

• This slow migration goes on for ~800 Myrs, until Jupiter and Saturn cross their 1:2 mean-motion resonance orbit.

• The gravitational interaction between the two bodies causes an increase in Saturn eccentricity, which then affects the 2 other giant planets, eventually switching their orbits.

• In 4 Myrs, planets move from their original position to the current one. The disk of planetesimals “explodes” and loses 97% of it’s mass.

• Small bodies are ejected inwards (Late Heavy Bombardment), and outwards (Oort Cloud).
The “Nice” Model

100 Myr, beginning of planetary migration

882 Myr, just after reaching 1:2 MMR

879 Myr, just before reaching 1:2 MMR

1082 Myr, final state, only 3% of initial disk mass remains.

Figure 2 | The planetary orbits and the positions of the disk particles, projected on the initial mean orbital plane. The four panels correspond to four different snapshots taken from our reference simulation. In this run, the four giant planets were initially on nearly circular, co-planar orbits with semimajor axes of 5.45, 8.18, 11.5 and 14.2 AU. The dynamically cold planetesimal disk was 35M⊕, with an inner edge at 15.5 AU and an outer edge at 34 AU. Each panel represents the state of the planetary system at four different epochs: a, the beginning of planetary migration (100 Myr); b, just before the beginning of LHB (879 Myr); c, just after the LHB has started (882 Myr); and d, 200 Myr later, when only 3% of the initial mass of the disk is left and the planets have achieved their final orbits.
The “Nice” Model
Any evidence?

Source: Alessandro Morbidelli
Summary

The Nice model

- explains the actual position of the giant planets and the Kuiper belt,
- explains the origin of the Oort cloud,
- solves the mystery of the missing mass in the Kuiper belt,
- explains the LHB.

but...

- everything is based on the assumption of a more compact solar system at the beginning...
- Giant exoplanets have been observed very close to their star so... why not?
References:


- + all articles written by these authors
Dynamics of the Trans-Neptunian Population

By Yacine Saidi
Outline

- Definition of the Trans-Neptunian Population
- Subdivisions of TNOs
- Orbital distribution of TNOs
- Qualitative and quantitative information about TNOs
- Stability of Orbits in the Kuiper Belt
A **Trans-Neptunian Object** (TNO) is any small object in the Solar System that revolve around the Sun beyond the orbit of the planet Neptune (at ~30 AU) to around hundreds of AU. The TNOs are the source of short period comets.
The first astronomer to suggest the existence of a trans-Neptunian population was Frederick C. Leonard. In 1930, soon after Pluto's discovery.

The possibility that TNO might be present beyond Pluto was the subject of published speculation by Kenneth Edgeworth as early as 1943.

In 1951, Gerard P. Kuiper, wrote a review paper that included speculation about objects beyond Pluto.

The first TNO was discovered in 1992 by David Jewitt and graduate student Jane Luu. Designated 1992 QB₁, the body moves in a nearly circular orbit at a distance from the Sun of about 44 AU.

In 2006 after the IAU resolution...
TNOs Like Structure

In 2006, astronomers have resolved dust disks believed to be Kuiper belt-like structures around nine stars other than the Sun.

HD 53143  No radial variation in disk structure (width >55 AU)

HD 139664  Belt-like morphology with a dust peak 83 AU and distinct outer boundary at 109 AU

Subdivisions of Trans-Neptunian Population

TNOs are often classified according to their distance from the Sun and their orbit parameters to two distinct categories: **Kuiper Belt Objects** (KBOs) and **Scattered Disk Objects** (SDOs).

[Diagram of the Trans-Neptunian Population with categories: Kuiper Belt Objects and Scattered Disk Objects, further divided into Classical KBOs, Resonant KBOs, Cold, and Hot.]

[Based on Morbidelli, A. 2008 classification]
Basic Orbital Dynamics

All objects in the Solar System move along elliptical paths with the Sun at one of the two foci of the ellipse.

The shape of the ellipse can be completely determined by:

\[ e = \left(1 - \frac{\beta^2}{\alpha^2}\right)^{1/2} \]

- **a**: semimajor axis
- **e**: eccentricity

**Orbits is:**
- Circular \( e = 0 \)
- Elliptic \( 0 < e < 1 \)
- Parabolic \( e = 1 \)
- Hyperbolic \( e > 1 \)

**f**: true anomaly of the body
**E**: eccentric anomaly
**M**: mean anomaly \( M = E - e \sin \nu \)

\[ E = \nu + \tau \] with \( n \) is the orbital frequency

\[ n = \sqrt{\frac{GM\alpha^{-3}}{2}} \]
Basic Orbital Dynamics

To characterize the orientation of the ellipse in space, with respect to an arbitrary orthogonal reference frame centered on the Sun, we need to introduce three additional angles:

\( i \): inclination of the orbital plane
\( \Omega \): longitude of the ascending node
\( \omega \): argument of the perihelion

For the case where \( i=0 \) and/or \( e=0 \) it’s useful to introduce:

\( \varpi \): longitude of perihelion \( \varpi = \omega + \Omega \)
\( \lambda \): mean longitude \( \lambda = M + \varpi \)
Orbital Distribution of TNOs

Classical KBOs (cubewanos)
40 AU ~< a ~< 47 AU
e < 0.1

Resonant KBOs
Plutinos (2:3): a ~ 39 AU
e ~ 0.1 - 0.3
Twotinos (1:2): a ~ 47 AU
e ~ 0.1 - 0.3

Scattered disk
a > 50 AU
30 ~< q ~< 38 AU

Extended SDOs

[Morbidelli, A. 2008]
Orbital Distribution of Classical KBOs

Points with error bars show the model-independent estimation from limited subset of confirmed classical KBOs. Smooth line shows the best fit population model.

Cold classical KBOs $\Rightarrow$ $i<4^\circ$

Hot classical KBOs $\Rightarrow$ $i>4^\circ$

The cold and hot populations are roughly equal in number.

Outer Solar System

Jupiter, Saturn, Uranus & Neptune

Kuiper Belt
- Cubewanos
- Plutinos
- Pluto

Escaped from Kuiper Belt
- Periodic Comets
- Centaurs (all objects with 5.2<a<30 AU)
- SDOs

[Minor Planet Center]
Largest Known TNOs

- Eris
- Pluto
- Makemake
- Haumea
- Sedna
- Orcus
- Quaoar
- Varuna

Credit: NASA, ESA, and A. Feild (STScI)
**Most Distant Known TNOs**

- *Sedna* is the most distant solar system object ever discovered. It is twice as far from the sun as any other solar system object and three times farther than Pluto or Neptune.

- Highly elliptical orbit, with its **aphelion** estimated at 975 AU and its **perihelion** at about **76.16 AU**.

- Sedna's precise orbital period is not yet known, but it is calculated at between **10.5** and **12.0** thousand years.

Credit: NASA/JPL-Caltech/R. Hurt (SSC-Caltech)
**Orbital Elements: Mean Motion Resonance**

A **Mean motion resonance (MMR)** is a dynamical situation where the ratio of the orbital periods of two orbiting objects can be simply expressed as the ratio two small integer.

**MMR** can either stabilize or destabilize the orbit.

Stabilization occurs when the two bodies move in such a synchronized fashion that they never closely approach: E.g. Pluto and the plutinos in general are in stable orbits. This is because a 2:3 resonance keeps them always at a large distance from it.

Credit: Wikipedia
**Orbital Elements:** Secular resonance

A **Secular resonance** is a dynamical situation which occur when the precession of two orbits is synchronized (a precession of the perihelion, of the ascending node, or both).

A small body in secular resonance with a much larger one (e.g. a planet) will precess at the same rate as the large body. Over long times (a million years, or so) **a secular resonance will change the eccentricity** \( (e) \) **and inclination of the small body** \( (i) \).

**Resonance:**

\[ g = g_n \text{ (perihelion: affects e)} \]
\[ s = s_n \text{ (nodal: affects i)} \]

\( g \) : frequency longitude of perihelion  
\( s \) : frequency longitude of node

Credit: Mobidelli talk 2008
Stability of Orbits in the Kuiper Belt

The dynamical lifetime for small particles in the Kuiper Belt derived from 4 billion year integrations. The color of each strip represents the dynamical lifetime of the particle.

- Yellow: Stable
- Black: Unstable

Orbital distribution of the real objects:
- Green circles: objects with $i < 4^\circ$
- Yellow circles: objects with $i > 4^\circ$

[The Duncan, Levison & Budd (1995) numerical survey]
Most Distant Known TNOs

Stable orbits at large $e$ in MMRs

[The Duncan, Levison & Budd (1995) numerical survey]
Stability of Orbits in the Kuiper Belt

unstable region, caused by secular resonances

$q = 30\text{AU}$

$q = 35\text{AU}$

Initial Eccentricity

Initial Semi-major Axis

Dynamical Lifetime (Yrs)

$4 \times 10^9$

$3 \times 10^9$

$2 \times 10^9$

$10^9$

$0$
Most of the observed TNOs objects (with the exception of scattered disk bodies) are associated with stable zones. Their orbits do not significantly change over The age of the solar system.
Summary

Trans-Neptunian Population

- **Kuiper Belt Object**
  - Classical KBOs (cubewanos): 40 AU \(< a \leq 47\) AU \& \( e < 0.1 \)
    - Hot population \( i > 4^\circ \)
    - Cold population \( i < 4^\circ \)

Resonant KBOs
  - Plutinos (2:3) \( a \approx 39\) AU \& \( e \approx 0.1 - 0.3 \)
  - Twotinos (1:2) \( a \approx 47\) AU \& \( e \approx 0.1 - 0.3 \)

- **Scattered disk** with \( a > 50\) AU and \( 30 \leq q \leq 38\) AU

The dynamical structure of the Kuiper Belt has been sculpted by a combination of a mean-motion resonance and secular resonance during the formation of the giant planets.
References


Physical properties of Trans-Neptunian Objects

C. Ejeta
Outline

• Introduction
• Sizes and Albedos Of TNOs
• Surface Composition( colors & spectra)
• TNO Binaries
Introduction

- TNOs are populating the Kuiper Belt beyond Neptune and are believed to retain the least altered material of the Solar System.

Why do we study TNO population?

- provides the richest information about the evolution of the solar system as it carries the scars of the accretional & evolutionary processes that molded the current form of the outer solar system.
a) Size and Albedo:

(i) Size of TNOs can be estimated from their brightness:
\[ p R^2 = 2.235 \times 10^{16} \times 10^{0.4(M_F - M_F(1,1,0))} \]

(ii) From reflected sunlight & thermal radiation:

**Reflected sunlight:**

\[ F_{TNO} = \frac{F_{sun} \pi R^2 \alpha p(\varphi)}{\Delta^2 r^2} \]

- \( F_{TNO} = \) flux of TNO
- \( F_{sun} = \) solar flux
- \( R = \) radius
- \( 4(2) = \) fast (slow) rotator

**Thermal flux:**

\[ \frac{F_{sun} \pi R^2 (1-a)}{\Delta^2} = \sigma T^4 4(2)\pi R^2 \]

- \( a = \) albedo, \( p(\varphi) = \) phase function
- \( r = \) heliocentric distance
- \( \Delta = \) geocentric distance

when the object is at opposition, \( p(0) = 1 \) and \( \Delta = r-1 \) are assumed:

\[ F_{TNO} \sim \frac{\alpha R^2}{r^4} \quad T \sim (1-a)^{1/4} r^{-1/2} \]
• 2003EL₆₁ (surface H₂O ice abundant).

• 2005FY₉ & Eris (CH₄ ice surface dominated group)
b) Composition
1 Colors

- TNOs have surfaces showing different colors, from neutral to very red which can be an indicative for diversity in their inherent surface composition and different degree of processing they have undergone.

- Can be used to classify objects into groups- color taxonomy: derived based on color indices (B-V, V-R, V-I, V-J, V-H, & V-K) , identifies the presence of four groups: BB, BR, IR, & RR.
  - the trend from neutral (BB) to very red (RR) groups indicates the possible sequence of alteration processes (collisions, resurfacing, craters, UV and/or energetic particle bombardment, etc....)
Correlations between colors and orbital elements:

[Doressoundiram, 2005]
Classical TNOs with low $e$ & $i$ exhibit very red surface color

[Doressoundiram, 2005]
• Classical objects (between the 2:3 & 1:2 resonances) with high eccentricity (and inclination) are preferentially neutral / slightly red.

• No clear trend is evident for scattered TNOs (a > 50 AU), nor for the plutinos, which appear to lack any trends in their surface colors.

• No trend can be drawn regarding correlations of colors with size or heliocentric distances.

• At present, no fully convincing mechanism explaining the anisotropic color distribution within TNO population exists.
2 SPECTRA

- TNO population represents some of the faintest objects in the solar system, as a result only brightest ones (~10) have been observed spectroscopically.

**Diagnostic wavelength range: 0.4-2.5μm**

- Visible Wavelength range (0.4-1μm): Important for detecting organic materials and aqueously altered minerals, such as phyllosilicates on the surface.

- Near-Infrared wavelength range (1-2.5μm): is the most diagnostic region for determining the presence of ices. Some TNOs are featureless in this region.
1. Methane-ice dominated spectra:

-Eris, Pluto, Sedna, & 2005FY9 have IR spectra dominated by CH4 absorption.

[Licandro et al, 2006]
2. Water-ice dominated spectra:

2003 EL61 (D~2000Km)

[Trujillo et al., 2007]

50000 Quaoar (D~1300Km)

[Jewitt and Luu, 2004]
• Include the largest objects after the methane-dominated objects.

• The spectral feature in Quaoar (a steep, positively sloped continuum from 0.4 to ~1.3μm wavelength), is characteristic of a strong ultraviolet absorber, probably associated with an organic surface composition.

• The ice on these objects is known to be crystalline as it sows a narrow band at 1.65μm that's absent in the spectrum of amorphous ice, which indicates that it has been raised above the critical temperature for transformation (roughly 100/110) at some point in its history. The heating could have been occurred in the deep interiors of the TNOs provided that there is a way for the heated ice at the depth to be emplaced onto the surface.
3. Water-ice spectra with the presence of methanol-like features:

- TNO 55638 = 2002VE$_{95}$ show spectra with 2.27μm band characteristic of CH$_3$OH

[Barucci et al. 2006]
TNO Binaries

- Similar colors – common surface materials
- Most are equal-sized systems
- Origin:
  - formation-unlikely
  - capture (for similar-sized components)
  - impact (for small satellites of large TNOs)

- CCDOs have more binaries than the remainder TNO population, which suggests that diverse dynamical histories of the bodies have had an effect on the survival of binaries.
• Relative sizes:
\[ \frac{R_1}{R_2} = \sqrt{\frac{A_1}{A_2}} = \sqrt{\frac{p_2}{p_1}} 10^{0.2 \Delta_{mag}} \]

• Total mass of the system:
\[ M_{sys} = \frac{4\pi^2 a^3}{GT^2} \]

• Individual radii:
\[ R_1 = \left( \frac{3V_{sys}}{4\pi(1-10^{-0.6\Delta_{mag}})} \right)^{\frac{1}{3}} 10^{0.2 \Delta_{mag}} \]
\[ R_2 = R_1 10^{-0.2 \Delta_{mag}} \]

• Geometric albedo:
\[ p_\lambda = \left( \frac{C_\lambda}{R_{eff}} \right)^2 10^{-0.4 H_\lambda} \]
\[ R_{eff} = \sqrt{R_1^2 + R_2^2} \]
<table>
<thead>
<tr>
<th>object</th>
<th>a (km)</th>
<th>ε</th>
<th>T (days)</th>
<th>M (10^{18} kg)</th>
<th>ρs (g cm^{-3})</th>
<th>ρ (g cm^{-3})</th>
<th>J / P</th>
<th>refs</th>
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<tr>
<td>(136199) Eris</td>
<td>36,000</td>
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<td>14</td>
<td>16,400</td>
<td>0.86(7) (V)</td>
<td>2.26(25)</td>
<td>0.16</td>
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<td>(134340) Pluto/Charon</td>
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<td>19,571(4)</td>
<td>0.00000(7)</td>
<td>6.387230(1)</td>
<td>13,050(620)</td>
<td>0.51–0.71 (V)</td>
<td>2.03(6)</td>
<td>0.40</td>
<td>[2]</td>
</tr>
<tr>
<td>Nix</td>
<td>48,670(120)</td>
<td>0.002(2)</td>
<td>24.8562(13)</td>
<td>1.521(65)</td>
<td>0.38 (V)</td>
<td>1.65(6)</td>
<td></td>
<td>[2]</td>
</tr>
<tr>
<td>Hydra</td>
<td>64,780(90)</td>
<td>0.005(1)</td>
<td>38.2065(14)</td>
<td>0.1–2.7</td>
<td>0.01–0.35</td>
<td>2.0</td>
<td></td>
<td>[4]</td>
</tr>
<tr>
<td>(136108) 2003 EL61</td>
<td>49,500(400)</td>
<td>0.050(3)</td>
<td>49.12(3)</td>
<td>4.210(100)</td>
<td>0.71(1) (v)</td>
<td>2.9(4)</td>
<td>0.53</td>
<td>[3]</td>
</tr>
<tr>
<td>(47171) 1999 TC26</td>
<td>7,720(460)</td>
<td>0.22(2)</td>
<td>50.4(5)</td>
<td>14.4(2.5)</td>
<td>0.08(3) (v)</td>
<td>0.5(3/2)</td>
<td>0.31</td>
<td>[5]</td>
</tr>
<tr>
<td>2001 QC298</td>
<td>3,690(70)</td>
<td>0.34(1)</td>
<td>19.2(2)</td>
<td>10.8(7)</td>
<td>0.08 (V_{sun})</td>
<td>1.0</td>
<td>1.16</td>
<td>[6]</td>
</tr>
<tr>
<td>(26308) 1998 SM105</td>
<td>11,310(110)</td>
<td>0.47(1)</td>
<td>130(1)</td>
<td>6.78(24)</td>
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<td>0.7(3/2)</td>
<td>0.56</td>
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</tr>
<tr>
<td>(65489) Ceto/Phorcys</td>
<td>18,400(50)</td>
<td>&lt;0.014</td>
<td>9.557(8)</td>
<td>5.4(4)</td>
<td>0.08(2) (V)</td>
<td>1.4(6/3)</td>
<td>0.89</td>
<td>[8]</td>
</tr>
<tr>
<td>(66652) 1999 RZ212</td>
<td>46,600(170)</td>
<td>0.46(1)</td>
<td>46.263(6/74)</td>
<td>3.7(4)</td>
<td>0.17 (R)</td>
<td>1.0</td>
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</tr>
<tr>
<td>1998 WW31</td>
<td>22,300(800)</td>
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<tr>
<td>(88611) 2001 QT207</td>
<td>27,880(150)</td>
<td>0.241(2)</td>
<td>825(1)</td>
<td>2.51(5)</td>
<td>0.13 (r')</td>
<td>1.0</td>
<td>3.41</td>
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<tr>
<td>(42355) Typhon/Echidna</td>
<td>18,300(30)</td>
<td>0.53(2)</td>
<td>18.971(1)</td>
<td>0.96(5)</td>
<td>0.05(V)</td>
<td>0.47(18/10)</td>
<td>2.13</td>
<td>[12]</td>
</tr>
<tr>
<td>(58534) Logos/Zoe</td>
<td>8,010(80)</td>
<td>0.45(3)</td>
<td>312(3)</td>
<td>0.42(2)</td>
<td>0.37(4) (R)</td>
<td>1.0</td>
<td>2.65</td>
<td>[13]</td>
</tr>
<tr>
<td>2003 QY90</td>
<td>7,000–13,000</td>
<td>0.44–0.93</td>
<td>306–321</td>
<td>0.3–1.7</td>
<td>0.19–0.35 (r')</td>
<td>1.0</td>
<td>2.14</td>
<td>[14]</td>
</tr>
</tbody>
</table>
Summary

- TNOs show strong color diversity when compared to other solar system bodies.
- There is color-orbital inclination correlation for the Classical TNOs.
- It’s not clear whether the color diversity is the result of compositional diversity or whether it’s the result of some evolutionary processes (e.g., collisions or surface irradiation processes).
- There is a correlation between the presence of ice and the size of the TNOs.
- No general explanation for the spectral variation is yet available (due to either variation in the initial composition or in variation in the subsequent history)
References:

• The solar system beyond Neptune (Book) by M.A. BARUCCI, H.BOEHNHARDT, D.P.CRUİKSHANK, AND A.MORBIDElli; The University of Arizona Press, 2008

• Trans-Neptunian Objects and Comets (Book) by D.Jewitt, A. Morbidelli, H. Rauer. Springer-Verlag Berlin Heidelberg, 2008

• D.P.Cruikshank et al., 2007, Protostars and planets V, pp.879
Physical and chemical properties of comets

Manuela Lippi
Chemeda Tadesse
Comet structure


Comet scheme from Northern Arizona Univ. webpage: http://www4.nau.edu/meteorite/Meteorite/Book-GlossaryC.html
The Nucleus

- When a comet is far from the Sun, only the nucleus is visible.
- The nucleus is a small celestial body, several kilometers in diameter, constituted mainly of water ice and rocks.
- Observations of a comet nucleus far from the Sun (when the comet is inactive) are difficult because the nucleus is a small and faint object.
- Observations of the nucleus when the comet is close to the Sun (active comet) are impossible since the nucleus is completely hidden by the coma.

Physical properties of the Nucleus

- Shapes and sizes from light curves (see for example C. Tubiana PhD thesis):
  - The nucleus is small and far.
  - Sometimes activity is anyway present.
  - ‘in situ’ measurements. In the last two decades spacecraft fly-by of 6 comets:
    - 1P/Halley (1996, Giotto, Vega1, Vega2, Suisei, Sakigake)
    - 26P/Grigg-Skjellerup (1992, Giotto)
    - 19P/ Borrelly (2001, Deep Space 1)
    - 81P/ Wild (2004, Stardust)
- Albedo from radiometric methods: 0.02 ÷ 0.06.
- No Masses or Densities were measured directly. Density estimates from non-gravitational forces (~ 0.5 g/cm³).
- Models for the internal structure of the nucleus => Rubble pile.

Comet 81P/Wild, Image taken by Stardust mission.

An example of rubble pile model for a generic comet.
Chemical composition: Parent Molecules

- Parent Molecules are the molecules released directly from the nucleus during the sublimation process that occur when the comet get close to the Sun.
- Among these molecules, Water is the most abundant one.
- Parent molecules are also present in molecular clouds.
- Other small organics are found in comets. The same organic molecules are found in molecular clouds.
Comets Chemical

If we consider the relative abundances of CH-bearing molecules, comets can be classified in Organic Normal, Organic Depleted, Organic Enriched. Actually there is not a real sharp separation between the three classes, but differences between comets are clear.

Comparison of Comets with Interstellar Material

- The similarity of composition of comets and dense interstellar clouds suggests that possibly comets contain unaltered interstellar material.
- However, there are differences in the relative abundances of some cometary species when compared to their interstellar values, suggesting that the material was in part processed.
- Location, epoch and source of this processing is largely to be determined.

<table>
<thead>
<tr>
<th></th>
<th>High Mass Star-forming regions</th>
<th>Solar-Type Star-forming regions</th>
<th>Comets average</th>
</tr>
</thead>
<tbody>
<tr>
<td>H₂O</td>
<td>100</td>
<td>100</td>
<td>100</td>
</tr>
<tr>
<td>CO</td>
<td>9–20</td>
<td>5–50</td>
<td>1.8–30</td>
</tr>
<tr>
<td>CO₂</td>
<td>12–20</td>
<td>12–37</td>
<td>3–6</td>
</tr>
<tr>
<td>CH₃OH</td>
<td>0–22</td>
<td>0–25</td>
<td>1.8–2.5</td>
</tr>
<tr>
<td>CH₄</td>
<td>1–2</td>
<td>&lt;1</td>
<td>0.14–1.5</td>
</tr>
<tr>
<td>H₂CO</td>
<td>1.5–7</td>
<td>—</td>
<td>0.4–4</td>
</tr>
<tr>
<td>OCS</td>
<td>0–0.3</td>
<td>&lt;0.08</td>
<td>0.1–0.4</td>
</tr>
<tr>
<td>NH₃</td>
<td>0–5</td>
<td>—</td>
<td>0.5–1.5</td>
</tr>
<tr>
<td>HCOOH</td>
<td>0.4–3</td>
<td>—</td>
<td>0.09</td>
</tr>
<tr>
<td>C₂H₆</td>
<td>&lt;0.4</td>
<td>—</td>
<td>0.11–0.67</td>
</tr>
<tr>
<td>HCN</td>
<td>&lt;3</td>
<td>—</td>
<td>0.1–0.3</td>
</tr>
<tr>
<td>C₂H₂</td>
<td>—</td>
<td>—</td>
<td>0.1–0.5</td>
</tr>
</tbody>
</table>

Adapted from: Ehrenfreund et al., Interstellar Material to Cometary Particles and Molecules, Comets II, 2005.

Origin of cometary material.

- The presence or not of some very volatiles ices can be related at the temperature at which comets formed.
- Relative abundances of parent molecules may provide informations on the conditions under which cometary ices formed. Some cometary ices have similar characteristics to the interstellar ice:

\[
\begin{align*}
\frac{\text{CH}_4}{\text{CO}} &= \begin{cases} 
\sim 1 & \text{in condensed phase in the interstellar material} \\
\sim 0.001 & \text{in gas phase in the interstellar material}
\end{cases} \\
\geq 0.1 \text{ in Comets}
\end{align*}
\]

- Relative abundance can be used to understand in which part of the proto-planetary nebula comets formed: as example: \(\frac{\text{H}_2\text{CO}}{\text{CH}_3\text{OH}}\); Methanol can be converted to formaldehyde by photoprocessing:

\[
\begin{align*}
\Rightarrow & \begin{cases} 
\text{H}_2\text{CO} > \text{CH}_3\text{OH} & \text{Formation region close to the Sun} \\
\text{H}_2\text{CO} < \text{CH}_3\text{OH} & \text{Formation region far from the Sun}
\end{cases}
\end{align*}
\]
Other aspects: Isotopic ratios

<table>
<thead>
<tr>
<th>Isotopes</th>
<th>comet</th>
<th>method</th>
<th>cosmic ratio</th>
<th>comet ratio</th>
<th>ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>[D]/[H]</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>H₂O⁺</td>
<td>1P/Halley</td>
<td>M</td>
<td>1.5 x 10⁻⁵</td>
<td>3.08 ± 0.53 x 10⁻⁴</td>
<td>[1]</td>
</tr>
<tr>
<td>H₂O</td>
<td>C/1996 B2</td>
<td>R</td>
<td></td>
<td>3.02 ± 0.22 x 10⁻⁴</td>
<td>[2]</td>
</tr>
<tr>
<td></td>
<td>C/1995 O1</td>
<td>R</td>
<td></td>
<td>2.9 ± 1.0 x 10⁻⁴</td>
<td>[3]</td>
</tr>
<tr>
<td>HCN</td>
<td>C/1995 O1</td>
<td>R</td>
<td></td>
<td>3.3 ± 0.8 x 10⁻⁴</td>
<td>[4]</td>
</tr>
<tr>
<td>NH</td>
<td>C/1996 B2</td>
<td>V</td>
<td></td>
<td>2.3 ± 0.4 x 10⁻³</td>
<td>[5]</td>
</tr>
<tr>
<td>CH₃OH</td>
<td>1P/Halley</td>
<td>M</td>
<td></td>
<td>&lt; 3 x 10⁻²</td>
<td>[6]</td>
</tr>
<tr>
<td>[¹⁵O]/[¹⁶O]</td>
<td></td>
<td></td>
<td></td>
<td>&lt; 1 x 10⁻²</td>
<td>[7]</td>
</tr>
<tr>
<td>H₂O⁺</td>
<td>1P/Halley</td>
<td>M</td>
<td>2.0 x 10⁻³</td>
<td>1.93 ± 0.12 x 10⁻²</td>
<td>[1]</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>2.13 ± 0.18 x 10⁻³</td>
<td>[2]</td>
</tr>
<tr>
<td>[¹³C]/[¹²C]</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>CN</td>
<td>1P/Halley</td>
<td>V</td>
<td>1.1 x 10⁻²</td>
<td>1.05 ± 0.13 x 10⁻²</td>
<td>[8]</td>
</tr>
<tr>
<td>HCN</td>
<td>C/1996 B2</td>
<td>R</td>
<td></td>
<td>2.9 ± 1.0 x 10⁻²</td>
<td>[9]</td>
</tr>
<tr>
<td></td>
<td>C/1995 O1</td>
<td>R</td>
<td></td>
<td>1.11 ± 0.18 x 10⁻²</td>
<td>[10]</td>
</tr>
<tr>
<td>[¹⁵N]/[¹⁴N]</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>CN</td>
<td>1P/Halley</td>
<td>V</td>
<td>3.6 x 10⁻³</td>
<td>&lt; 3.6 x 10⁻²</td>
<td>[8]</td>
</tr>
<tr>
<td>HCN</td>
<td>C/1995 O1</td>
<td>R</td>
<td></td>
<td>3.1 ± 0.4 x 10⁻²</td>
<td>[11]</td>
</tr>
<tr>
<td>[¹³S]/[³²S]</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>atomic S</td>
<td>1P/Halley</td>
<td>M</td>
<td>4.2 x 10⁻²</td>
<td>4.5 ± 1.0 x 10⁻²</td>
<td>[12]</td>
</tr>
<tr>
<td>CS</td>
<td>C/1995 O1</td>
<td>R</td>
<td></td>
<td>3.7 ± 0.4 x 10⁻²</td>
<td>[11]</td>
</tr>
</tbody>
</table>

- Isotopic contents can be used as a proxy to compare comet material with solar, terrestrial material and interstellar medium.

- Oxygen, Carbon and Nitrogen isotopic ratios are consistent with the terrestrial values.

- One can conclude again that part of cometary material has interstellar origins while part was processed.

D/H ratio

• D/H ratio is notably higher in meteorites and comets.

• D/H ratio for comets is a factor of ten higher than the proto-solar nebula and about twice the one measured for the Earth.

• Hot Cores
  D/H = $30 \div 400 \times 10^{-5}$
  Interstellar origin of Cometary ices.

• D/H ratio to test the role that comets had in delivering water to the Earth.


Note: Also 8P/Tuttle show the same value (G. Villanueva et al. 2008)

The Dust Tail

• Dust is released from the comet during the sublimation of the volatile ices.

• Dust study by *in situ* measurements, spectroscopy from the ground or the space and laboratory analysis of interplanetary dust particles of probable cometary origin.

• Dimensions of the Dust grains are estimated to be between 2 and 10 μm.

• These grains are bigger than the interstellar grains.

Comet McNaught (C/2006 P1) tail.
Elemental composition of comets.

If we analyze the elements present in the cometary dust we can notice that for more volatiles elements the composition is more close to the ISM than to the Solar one. Probably comets come from the outer solar system and beyond and they where less processed than other bodies.

<table>
<thead>
<tr>
<th>element</th>
<th>Ikeya-Seki$^a$</th>
<th>P/Halley</th>
<th>Solar system</th>
<th>CI-chondrite</th>
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</thead>
<tbody>
<tr>
<td></td>
<td>dust</td>
<td>dust + ice</td>
<td></td>
<td></td>
</tr>
<tr>
<td>H</td>
<td>2 025.</td>
<td>4 062.</td>
<td>$2.6 \times 10^6$</td>
<td>492.</td>
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<tr>
<td>Li</td>
<td>†</td>
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<td></td>
<td>0.0053</td>
</tr>
<tr>
<td>C</td>
<td>814.</td>
<td>1 010.</td>
<td>0.</td>
<td>70.5</td>
</tr>
<tr>
<td>N</td>
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<td>95.</td>
<td>291.</td>
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</tr>
<tr>
<td>O</td>
<td>890.</td>
<td>2 040.</td>
<td>2 216.</td>
<td>712.</td>
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<tr>
<td>Na</td>
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<td>5.3</td>
</tr>
<tr>
<td>Mg</td>
<td>(=100.0)</td>
<td>100.0</td>
<td>=100.0</td>
<td>=100.0</td>
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<tr>
<td>Al</td>
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<td>6.8</td>
<td>7.9</td>
<td>7.9</td>
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<tr>
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<td>185.</td>
<td>93.</td>
<td>93.</td>
</tr>
<tr>
<td>P</td>
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<td>1.0</td>
<td></td>
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<tr>
<td>S</td>
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<td>72.</td>
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<td>48.</td>
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<tr>
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<td>0.35</td>
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<tr>
<td>Ca</td>
<td></td>
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<td>5.7</td>
<td>5.7</td>
</tr>
<tr>
<td>Ti</td>
<td>&lt;0.02</td>
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<td>0.22</td>
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<tr>
<td>V</td>
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<tr>
<td>Cr</td>
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<tr>
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<tr>
<td>Fe</td>
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<td>52.</td>
<td>84.</td>
<td>84.</td>
</tr>
<tr>
<td>Co</td>
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<td>0.21</td>
<td>0.21</td>
</tr>
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<td>Ni</td>
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<td>4.1</td>
<td>4.6</td>
<td>4.6</td>
</tr>
<tr>
<td>Cu</td>
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<td></td>
<td>0.049</td>
<td>0.049</td>
</tr>
</tbody>
</table>

$^a$ In addition, abundances relative to Na were determined to be $1.6 \times 10^{-3}$ for K and $< 2.5 \times 10^{-5}$ for Li [90].

Comets are object that live the majority of their life far from the Sun. They are less altered than other bodies by various processes that occurred and still occur in our Solar System.

The structure and chemical composition of comets clearly provide key data on the processes and conditions in the outer Solar System at the time of its formation.

Comets are probably formed from a mixture of interstellar and processed material, and their composition will differ according to their place of formation in the proto-planetary disk.

Contrary to TNOs, comets have eccentric orbits, and they can release part of their primordial material when they come close to the Sun.

Thank You!!!